Undergraduate Lecture Notes in Physics

Mark Gallaway

An Introduction to Observational Astrophysics

Springer

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Contents

Acronyms

2MASS	Two Micron Allsky Survey: An infrared astronomical survey.
AAVSO	American Association of Variable Star Observers: A group of variable star observers and also the catalogue they have created.
ADC	Analogue-to-digital converter: A device to convert an analogue signal such as voltage to a digital one such as count.
ADS	Aitken double star catalogue
Alt	Altitude: the vertical position of a telescope.
APT	Aperture photometry tool: A freeware multiplatform photometry.
AU	Astronomical unit: The mean distance between the Earth and the Sun, most often used in solar system physics.
Az	Azimuth: The horizontal position of a telescope.
CCD	Charge-coupled device: The detector in a modern astronomical digital camera.
CDS	Centre de données astronomiques de Strasbourg: The French body that hosts SIMBAD.
CMB	Cosmic microwave background: The afterglow, at microwave wavelengths, of the light that escapes during the era of recombination.
CMD	Colour-magnitude diagram: A plot of absolute magnitude against colour. The Hertzsprung-Russell diagram is an example if a colour-magnitude diagram.
CMOS	Complementary metal oxide semiconductor: A photon detector similar to the CCD.
CTE	Charge transfer efficiency: The rate at which electrons are lost during pixel-to-pixel transfer in a CCD.
CTIO	Cerro Tololo Inter-American Observatory: An international observatory based in Chile.
CYMK	Cyan, yellow, magenta, black: A colour scheme for colour printing.
Dec	Declination: The vertical location of an object on the celestial sphere.
DSS	Digital Sky Survey: A catalogue made from digitised optical Schmidt plates.
DSS	Deep Sky Stacker: An application for stacking frames.

E-ELT	European Extremely Large Telescope: A 39 m diameter telescope current under construction in Chile.
ESA	European Space Agency: The European Space body.
FFT	Fast Fourier transformation: A mathematical process used to apply a Fourier transformation.
FITS	Flexible Image Transport System: The standard format for astronomical images.
FOV	Field of view: The angular diameter of the area of the sky that can be seen by a telescope.
FWHM	Full width half maximum. The width of a curve measured form half its maximum value. In photometry, this is related to the seeing.
GMT	Greenwich Mean Time: The current time for an observer on the Greenwich meridian line - without adjustments for daylight saving.
GPS	Global Positioning System: A satellite positioning system operated by the US government.
GSC	Hubble Guide Star Catalogue. A large catalogue of bright stars used for the point of the Hubble Space Telescope.
GVCS	General Catalogue of Variable Stars: A catalogue of variable stars.
HD	Henry Draper Catalogue: A stellar catalogue.
HDC	Henry Draper Catalogue: An HD supplement.
HDEC	Henry Draper Catalogue: An HD supplement.
HI	Hydrogen 1: Monoatomic hydrogen.
HII	Hydrogen 2: Ionised hydrogen.
HR	Hertzsprung-Russell: As in the Hertzsprung-Russell diagram.
HST	Hubble Space Telescope: The first large spaceborne optical telescope.
IAU	International Astronomical Union: The international governing body for astronomy.
IC	Index Catalogues: A supplement to the NGC.
IFU	Integrated field unit: A device that allows whole-field spectrography.
IRPC	Infrared Processing Centre: NASA science center tasked for data processing, analysis and archiving of NASA's infrared data.
IRSA	Infrared Science Archive: An online database integrating various infrared astronomical catalogues.
ISO	Infrared Space Observatory: A spaceborne infrared telescope.
Kpc	Kiloparsec: 1000 parsecs.
LRGB	Luminance, red, green, blue: A colour image with a luminance frame.
LST	Local Sidereal Time: The current time at an observer's location based on the transit of astronomical objects across the local meridian.
MOS	Multi-object spectrography. A device for taking many spectra at once.
Mpc	Megaparsec: 1 million parsecs.
MPC	Minor Planet Center: The IAU body responsible for solar system objects. The MPC holds an extensive database of comets and asteroids.
NDF	Neutral density filter: A filter that reduces brightness without affecting the spectral ratio of the source.

NGC	New General Catalogue of Nebulae and Clusters of Stars: A large astronomical catalogue of non-stellar sources.
NOFS	Naval Observatory Flagstaff Station: A large observatory located in Arizona.
NOMAD	Naval Observatory Merged Astrometric Dataset: A large astronomical catalogue produced by, amongst others, the NOFS.
pc	Parsec: A unit of astronomical distance equating to approximately 3×10^{16} m. The parsec is the distance of an object that would have a 1 arcsecond parallax.
PNe	Planetary nebula: A type of nebula-related formation of a white dwarf.
PSF	Point spread function: In photometry the function that describes how the light from an object spreads over the frame. Most often a Gaussian.
QE	Quantum efficiency: The efficiency of a quantum device, for example, a CCD.
RA	Right ascension: The horizontal location of an object of the celestial sphere.
RBI	Residual bulk image: The appearance of a false star in an image caused by the saturation of a CCD.
RGB	Red, green, blue: A colour image.
RMS	Root mean square: The square root of the mean of the squares of the sample.
SIMBAD	Set of Identifications, Measurements, and Bibliography for Astronomical Data: An online searchable database of astronomical objects.
SNR	Signal-to-noise ratio: The ratio of the signal to the noise of the system. A general indication of uncertainty.
STScI	Space Telescope Science Institute: The organisation that manages the day-to-day operation of the Hubble Space Telescope.
UGC	Uppsala General Catalogue: A northern hemisphere catalogue of galaxies.
UT1	Coordinated Universal Time One: A more modern version of GMT. Effectively, there is currently little difference between UT1 and GMT.
WCS	World Coordinate System: The standard translation format for plate location to sky location translation.
WDS	Washington Double Star Catalogue.
WISE	Wide-field Infrared Survey Explorer: An infrared spaceborne telescope.
Z	Cosmological redshift: The ratio between the wavelengths of a spectral line.

Chapter 1
Introduction

Abstract During the course of your astrophysics degree, your will be undertaking observations of the skies. This book is designed to guide you through observational astrophysics through a series of practicals. You will learn important techniques in photometry, spectrography, astrometry and general imaging; as well as being navigated through some of astronomy's more arcane features.

1.1 The Reader

Most undergraduate astrophysics degrees have a practical component. This practical component would usually be undertaken using a small telescope, on campus if you are lucky, or otherwise at a large off-site facility. Remote observations are also sometimes used, these being done via an Internet link between a local computer and a robotic telescope. The amount of time you might spend at an instrument can be surprisingly short. Clearly this is not an ideal situation as observational astronomy is a complicated subject, and one that has developed over many centuries.

If you do not feel ready for the task that is in front of you, you are not alone. Scientific observation is a complex and somewhat daunting task, particularly when you find yourself alone in the dark late at night. This book was written to help new undergraduates studying astrophysics. It is intended to serve as a primer prior to observing, and also as a handy reference during the observing process itself. The book can also be referred to later when you are busy processing the images you took. However, a word of caution – this book will not act as a substitute for thorough note-taking. It is always a good idea to document what you are doing while you do it and while it is fresh in your mind. Also, with regards to the sections on data reduction and image analysis, the volume you are holding is not specific to any single observatory. Rather it is intended as a general reference. Nor is it intended to be a complete work for any of the areas of astronomy covered in the following chapters, all of which could, in themselves, be a complete book.

While this book is aimed at undergraduates, I hope that it should also be useful for postgraduates, particularly if you do not have much direct experience of observational astronomy, or if you are out of practice. Hopefully, it should also be

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useful for the amateur astronomer who wishes to do some scientific work with their home instruments. One exciting aspect of astronomy is that it is a modern science where the amateur can do highly-useful research.

1.2 Required Equipment and Software

For the purposes of this book, it's assumed that you have access to a permanently-mounted and fixed telescope. We also assumed that this is a Schmidt-Cassegrain type telescope (we'll discuss specifically what this is later on). The most common manufacturers of small telescopes are Meade and Celestron which are the most commonly-used by universities. (The make of the telescope only affects a small portion of this book, so don't worry if yours is different!)

We're also going to assume that the telescope has a mount that is capable of sidereal tracking, and also that it has an astronomical digital camera. The camera should be able to provide guidance information to the mount. Also, you will need science-grade astronomical filters (these would ideally be contained in the filter wheel that's part of the camera).

Your telescope may be controlled either by a computer or a hand controller. The camera will always be computer-controlled. There is a range of different software applications that exist for controlling camera and telescope; the most popular are ccdSoft and MaxImDL. These products are broadly similar, although, for the purposes of this book, we will be using MaxImDL. In addition, you may also find that your telescope had a powered focuser. This tool isn't essential to have, but it is a useful add-on.

Throughout this book, you will be asked to manipulate the images you take with your camera/telescope combination. Both ccdSoft and MaxImDL are capable packages, but both have associated licensing fees. There are also freeware image manipulation applications available for Windows, OSX and Linux. In this book, we will be using SAO DS9, Aperture Photometer Tool (APT), AstroImageJ and SALSAJ. Some universities may require you to use some professional-grade applications. A common example of such a tool is IRAF is versatile and fast, but it's only designed for Linux (this also extends to OSX in practise). IRAF requires the use of a command-line interface and it is notorious for its steep learning curve, so you may find that to be something of a speedbump when you start out. However, one of IRAF's merits is that it can be scripted to run a series of linked commands. This can speed many image analysis processes up a lot. Most of the image manipulation that this book will cover can be done in IRAF if you wish.

This book also covers tasks which can be done using commands in the Python programming language. Python is becoming increasingly popular in professional astrophysics, and I can't recommend it highly enough as a first language. I also highly recommend learning a programming language as it will help you in your course and also for getting a job post-degree. Python is open source (i.e., it's free), multi-platform, well-supported and easy to check for errors. It has many powerful add-in packages called modules, which include those for performing astronomical calculations, astronomical image manipulation, statistics and even a package to control IRAF viaPython (PYRAF). A number of Python distributions and versions are currently in use. The author's preference is for the Anaconda Python 2.7.6 distribution as it is easy to install and comes with most of the tools you use. However, other are available and might already be installed on your computer.

This book also cover spectrographs and radio observing. Most universities running astrophysics degrees have telescopes with cameras, but having a mounted spectrograph is rarer, and access to a radio telescope is even rarer. However, please do feel free to read these chapters even if you don't currently have access to these instruments. These are after all important tools in professional astronomy, and you never know when this knowledge might prove useful.

1.3 How to Use This Book

This book doesn't have questions for the reader to attempt, although it does provide worked examples for the trickier mathematics. Instead, there are practical activities for you to try. The idea is that you will gain the skills and the confidence you need by actually having a go at these things and seeing how they work in practise.

The initial chapters cover the underpinnings for the practical sessions introduced later. In most cases, it's possible to do the exercises from previously-acquired data. (This sometimes referred to as "canned data".) This does, however, provide a different experience from actually going out and doing the observations. (And don't underestimate how different sitting indoors in front of a computer in the mid-afternoon feels to actually being out there on a cold winter night.)

Unlike many books that you may encounter during your studies, this one is not limited to a single course or module. Rather, it's intended to be a companion that can guide you through all the stages of your course. Topics we will cover include general observing, solar system astronomy, solar physics, planetary science, stellar astrophysics, extragalactic astrophysics and cosmology. We will also touch on broader physical topics such as quantum mechanics (where relevant to the matter at hand).

The practical exercises are laid out in the order that I teach them in my own classes. Your course may have a different structure; if so, please feel free to skip back and forward as you need to. However, keep in mind that the initial chapters and practicals are vital as they cover essential observational skills. Even if you are not taught them, you should make yourself familiar with the materials and techniques therein. Whilst reading this book, you will likely also notice that some ideas are repeated between chapters. This is both to reinforce key concepts and also to allow the reader to skip forward to a desired section without the need to read every single item in between.

Your course likely requires a final-year project. Hopefully by the end of this book you will have the skills and the confidence needed to take this on and to succeed. This book will be able to help you when it comes to deciding on a project and its structure.

Astronomy is arguably the oldest of the sciences – the positions of the stars, the Moon and planets have been used for thousands of years to determine when to plant crops when to bring in harvests and for navigating the seas. The Mayan priesthood partly based their power on their ability to reliably predict eclipses, and many ancient monuments across the world feature astronomical alignments as part of their design. Astronomy's pre-history segues into what we would now consider astrology, just as chemistry has its origins in alchemy. Astronomy was only clearly established as an actual, recognisable science as recently as the Seventeenth Century. (Sir Isaac Newton's formulation of the theory of gravity was an essential achievement, as it provided a systematic theoretical foundation to what had previously been something of a disorganised grab-bag of facts, half-truths and conjecture. Because of this extensive pre-history, modern astronomy contains many odd and quirky traditions (as an example, the sexagesimal coordinate system ultimately originated with the Babylonians). You may well find some of these rather disorienting at first. However, you will get used to these quirks in time.

Throughout this book, we have preferentially used the SI unit system. There are some deviations from this rule as astronomy has plenty of non-SI measures, such as for instance the parsec, the Jansky and the Astronomical Unit. Where there is a suitable SI unit, though, we will be sticking to that, so you won't be bedevilled by ångströms or ergs. You will find Joules and nanometres, though. If you need to use a non-SI unit, please remember to make the appropriate conversion.

Before we move on, let's briefly define a couple of these unavoidable non-SI units. The Astronomical Unit (AU) is the median separation between the Earth and the Sun. It's equal to 149,597,870,700 m. The AU is most often used when dealing with distances inside the Solar System. The Parsec (pc) is the distance at which an object would show a parallax displacement of one second of arc. It equates to $3.08567758 \times 10^{16}$ m. Kiloparsecs (Kpc; 1 Kpc = 1,000 pc) and Megaparsecs (Mpc; 1 Mpc = 1,000,000 pc) are also often used in astronomy. For the most distant observable objects, the distance is expressed in terms of the degree of observable (z). Higher-z objects are further away, but the scale isn't linear and also, the actual distance will depend on the cosmological model being used. As a point of reference, objects at $z = 1$ are roughly halfway to the edge of the observable universe, whereas the Cosmic Microwave Background, is at $z = 2{,}000$. (The CMB effectively constitutes the "edge" of the observable universe, for all practical purposes.)

1.4 Introduction: Best Practice

The observatory that you will be using is a laboratory, just like any other that you may have worked in. As such you should treat it like a lab – read and understand the safety guidelines and the operating instructions before you begin observing.

It is very likely that there will be specific documentation relating to the use of the observatory. You need to make yourself aware of what is in these documents, and if you have any questions, it's a good idea to ask a member of the observatory staff.

It's a good idea to prepare properly for a night's observing. Before you set off you need to think about what you're planning to do and how you will go about achieving it. First of all, if your observation is part of a course, then make sure that you've read the practical notes and that you understand them. You should have a list of the observational targets that you want to work on and their co-ordinates in the sky. You should also know what exposure times and filters that you need. You should also note the position of the Moon in the sky. This can become important, especially if it's at or near full phase.

If it's possible to find out in advance, it's a good idea to know what instrument and camera you will be working on. Not all cameras or instruments are equal, and some specific setups may have implications for your work. It's also a good idea to have a list of backup objects, in case your primary targets are obscured. Never assume that there won't be a single solitary cloud parked in the worst possible bit of the sky. Telescope time is expensive, and the weather is not dependable.

It's also important to have some means to get your images off of the camera control computer. Examples of such means are memory sticks or cloud services. Different observatories have different standard methods, so check with the staff.

If you do as much preparation as you can in advance, you will have a rewarding and reasonably stress-free observing session.

The most important tool you will have with you whilst observing is your lab book. You need to record everything that happens during the course of your observations. Include the telescope name and its optical characteristics (we'll return to these details later on). Here is a list of the things you need to record:

- The time and date of each observation;
- The instrument you're using on the telescope;
- The exposure time for each observation;
- The filter used for each observation;
- The ID of the target that's being observed;
- Its location on the sky.

In addition, you should record the weather conditions, the sky brightness (if known) and the seeing of the observation (we'll show you how to work this out later on). These details are essential if you later wish to make a quantitative assessment of your data quality. If you are observing as part of a group, don't assume that someone else is making all the notes. They probably are not. Always make your own notes and make them as thorough and detailed as possible. Remember, precision is a key aspect of experimental science. If at all possible, always record the uncertainties on the data that you're taking. You should ask yourself the question, "Could another scientist reproduce my methods based on what I have written down?" The answer should always be yes.

Given the nocturnal nature of astronomy it can be extremely cold when you are observing so ensure that you dress appropriately for the weather, including footwear. It is easy to remove a jacket if you are too warm, but you cannot put one on that you have not brought. Telescope domes are unheated, so unless you are lucky enough to have a control room, you are, as likely as not going to exposed to the elements. You are also going to be stationary for considerable amounts of time. Consequently, hypothermia can become a problem. If you feel too cold, go somewhere warm for a while.

You will be observing in a dark environment for all but solar and radio observations. Initially, it may appear extremely dark, however after a while you will become dark adapted, and it will become easier to see. It takes about 20 min to become fully dark adapted , but only a few seconds of exposure to bright light to lose it. Remember this when entering lit buildings and when using a torch or flashlight. If you flash it in another observer's eyes, they will lose their dark adoption or stray light may fall into the instrument ruining the observation. Exposure to red light reduces the loss of dark adoption, although it still has an impact. If your observatory has red lighting, try to use it. Likewise, if you have access to red torches, please use these.

Chapter 2
The Nature of Light

Abstract Our understanding of the Universe at large comes almost exclusively from out understanding of light and how quantum mechanics explains the physical features we see in the spectra astronomical objects. Astronomers measure the brightness of astronomical objects in terms of flux and magnitude. The use of coloured filters enables us to examine specific characteristics of the object being observed.

2.1 Introduction

The physics of light are central to our understanding of astrophysics. For the vast bulk of history, astronomy was confined to the so-called visible spectrum, that wavelength region that can be seen by the human eye. In the year 1800, William Herschel – famous for the discovery of Uranus, the first new planetary discovery since prehistory attempted to measure the temperatures of individual colours in the rainbow. In fact, his efforts led to an inadvertent discovery. Herschel's thermometer carried on detecting warmth even after it was moved outside of the visible spectrum. Herschel had accidentally discovered the existence of infrared light.

Infrared astronomy began to develop as a quantitative science in the mid-1800s (the first actual object detected in the IR was a cow, as it happens).

Radio astronomy began as a discipline in the 1940s, having been made possible by advances in electronics and electrical technology. Subsequent to this, ultraviolet, X-ray, millimetre and submillimetre observations also became possible. Many of these depend on space-based or balloon-borne instruments, due to the opacity of the Earth's atmosphere at many wavelengths. (This is also why observatories tend to be built at high altitude – the more of the air you're above, the better. Water vapour, in particular, can bedevil infrared astronomy, due to its deep absorption features, and mid-infrared astronomy is effectively impossible from sea level.)

Classical physics presents light as a travelling wave consisting of an oscillating electric field with a perpendicular oscillating magnetic field. Both the electric and magnetic fields are at right angles to the direction of travel. The energy of the wave is inversely proportional to its wavelength, with longer waves having lower energies and shorter waves having higher energies. (In terms of visible light, red light has a longer wavelength than blue light, and so carries less energy.)

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The relationship between frequency and wavelength is given by Eq. 2.1, where c is the speed of light ν is frequency and λ is the wavelength. Hence as wavelength increase frequency decreases.

$$c = \nu\lambda \tag{2.1}$$

Conversely in quantum mechanics, light can be considered to be a particle, the photon, with a energy given by Eq. 2.2 where the energy is E, h is Planck's constant, frequency is ν and wavelength is λ.

$$E = h\nu = \frac{h\lambda}{c} \tag{2.2}$$

It is important to realise that light is not either a wave or a particle; rather it is both simultaneously and it is only the method of observation that determines if it has the features of a wave or a particle. Hence, when, for example, looking at lights interaction with optics it behaves wave-like, but within our cameras it behaves like a particle.

Within the research community, there is a tendency for different groups to have their own idiosyncratic means to describe the light they work with. Those who work at short wavelengths (such as X-ray astronomers) tend to characterise their photons by their energies. (This approach is also used with cosmic rays). Observers in the UV to IR range tend to refer to their light by wavelength. Lastly, those in the millimetre and radio regimes tend to describe things in terms of frequency.

The electromagnetic spectrum itself is continuous. For human convenience, we divide it up into a number of bands, although the exact cut-off locations for the bands are disputed. Table 2.1 shows these various bands with their typical energies, wavelengths and corresponding frequencies.

Table 2.1 Table showing the various bands of the electromagnetic spectrum and their approximate wavelength, frequency and energy

Name	λ nm	ν Hz	E eV
Radio	$>10^8$	3×10^8	10^{-8}
Microwave	$10^8 - 10^5$	$3 \times 10^9 - 10^{12}$	$10^{-5} - 0.01$
IR	$10^5 - 700$	$3 \times 10^12 - 4.3 \times 10^{14}$	0.01−2
Visible	7,000−400	$4 \times 10^{14} - 7.5 \times 10^{15}$	2−3
UV	400−1	$7.5 \times 10^{14} - 3 \times 10^{17}$	$3 - 10^3$
X-Ray	1−0.01	$3 \times 10^{17} - 3 \times 10^{19}$	$10^3 - 10^5$
Gamma ray	<0.01	$>3 \times 10^{17}$	$>10^5$

2.2 The Quantum Nature of Light

Photons are emitted when there is a change in the energy state of a system. For example, the model often used when introducing the atom to students is the idea of the electron orbiting the nucleus. Although this is strictly not true in all situations, it is a good enough model for most purposes. Electrons that are orbiting close to the nucleus have less energy than those orbiting further away. Hence, when an election gains energy it moves to a higher orbit and when an electron emits a photon is moves to a lower one.

In the late 1800s, Heinrich Hertz studied the photoelectric effect. He noticed that certain materials release an electrical current when exposed to light – but the release only occurred when they were exposed to certain wavelengths. If the wavelength was wrong, it didn't matter how bright the light was. No current would be observed. This was something of a puzzle – how did the materials 'know' that they were being exposed to the "wrong" sort of light? Why would they seem to care about anything except luminous intensity?

It was Albert Einstein who suggested a solution (and it was actually this solution that earnt him his Nobel Prize). He suggested that light is quantised, that is that its energy is delivered in discrete packets. These packets are, of course, the particles that we call photons. Also, Einstein theorised that instead of there being a uniform continuum of energy levels that electrons can exist in, there are actually just a few discrete ones – that is, electron energy levels are quantised too. For an electron to move between these fixed levels, an incoming or outgoing photon has to have just the right, energy. If it's wrong, then the photon can't absorb it and thus will not respond. This is why the materials in Hertz's experiments only reacted to certain wavelengths of light.

As to why electrons have this behaviour, imagine them for a moment not as particles but rather waves that are wrapped around an atomic nucleus. For a wave to wrap neatly, its orbital circumference has to correspond to a multiple of its wavelength. If the circumference doesn't follow this pattern, then some of the wave's peaks and troughs will overlap each other, and the wave will interfere with itself. Applying this to an electron, we can see that a situation would arise where an the electron could interfere itself out of existence, thus strongly-violating Conservation of Energy. This has the effect of forbidding electrons from occupying such orbits.

These insights, along with others, led to the emergence of quantum mechanics. It's worth noting that although quantum mechanics has many weird aspects, it is nonetheless one of the most successful and important scientific theories ever developed. It underlies the multi-billion dollar electronics and computing industries and it has practical applications throughout the whole of industry and commerce. It even has implications for seemingly unrelated fields such as biology and medicine, through things such as imaging machines to analyses of protein folding. Whilst quantum theory is certainly weird, it is a necessary weirdness.

The key insight of quantum theory is that the quantisation of energy in a bound system applies to every change between two states. Hence, not only is the energy of electrons quantised, so are the spins, the momentums and also molecular vibrations and rotations. There are some conditions where this rule breaks down, of course. The emission caused by the deceleration of electrons in a magnetic field is not quantised (note that we're considering free electrons here, moving outside of an atomic shell) and the emission caused by capture of electrons by an ion is also non-quantised. (Note again that the electron starts off free rather than bound – this is what makes all the difference.) The emission of photons due to the capture of electrons by atoms is called free-bound emission.

When the idea is applied to processes inside our Sun, quantisation leads to an important realisation.The Sun's core temperature is extremely high, around 15 million K. As a result the material inside it is in a highly ionised state – the Sun is made of plasma. This means most emission is of the free-bound variety, so the Sun should exhibit a continuous emission spectrum.

It doesn't. The Sun actually has some bright emission lines such as the Balmer and Lyman series.

What is going on? The answer lies in the fact that whilst the emission of photons has no dependency on direction, the absorption of photons does have a directional bias. As photons near the star's surface, there is more star below them than above, so by this point when a photon is absorbed, that photon is most likely from somewhere deeper inside the star. This causes dark features known as absorption lines to appear in the star's spectrum.

But also, given that a star's surface is much cooler and less dense than its core, it is possible for some bound absorption to occur at or near the surface. These will show up as bright emission lines, such as the aforementioned Balmer and Lyman features. The Balmer lines are caused by the transition of a hydrogen electron down to the level 2 energy state. The first Balmer line called the H α line, is seen at $\lambda = 656.3\,\mathrm{nm}$ (in the red area of the visual spectrum) and is caused by a 3-to-2 transition. The second Balmer line, Hβ represents the 4-to-2 transition, and so on through the other transitions.

The Lyman series covers the transitions down to the n = 1 energy level. The first, the 2-to-1 transition, is called the Lyman α line.

When seen under laboratory conditions, an emission line will be narrow. What broadening there is will simply be down to the optics used to observe it, and some due to the Uncertainty Principle. However, actual stars are less tidy than well-organised laboratories. Consequently, the thermal motion of the plasma within the star will spread the lines out due to the Doppler effect. In addition the bulk motion of the plasma and also the motion of the target with respect to the observer (rotation, radial velocity, etc.) will shift the line slightly toward either the blue or red ends of the spectrum. However, these effects don't broaden the lines.

The quantisation of light is the underpinning physics for the science of spectrography. In Chaps. 13 and 14 we will go into this subject in much more depth.

2.3 Measuring Light

You will often come across the terms **Luminosity** and **Brightness**. The luminosity of an object is the total amount of energy it is emitted across a specific wavelength range. The brightness is a related but a subtly different concept; the brightness of an object is the amount of energy it emits through a specific wavelength range and over a solid angle. Brightness is also dependent on the distance from the emitting object, and thus will decline with the square of said distance.

The **absolute magnitude** of an object is its brightness as it would be at a set distance. The point of this is that it allows a quick and valid comparison with the absolute magnitudes of other objects. For solar system objects, the reference distance is 1 AU. For non-solar system objects, absolute magnitudes are the magnitude the object would have if it were located at 10 pc.

In general you should use the term "apparent brightness", so as to avoid confusion.

You may also come across the term bolometric luminosity of brightness. In this case, this refers to the luminosity or brightness across the entire spectrum.

Stellar spectra broadly approximate to **black bodies**. That is, they can be modelled as perfect absorbers and emitters and they reflect next to no light. (A little-known fact about the Sun is that in terms of its reflectivity, it's actually about as black as coal.) This means that their output follows the pattern given by Planck's Law (See Eq. 2.3) where $B_\lambda(T)$ is the total radiation at wavelength λ and temperature T, h is the Planck's constant, k is the Boltzmann constant and c as usual is the speed of light in a vacuum.

$$B_\lambda(T) = \frac{2hc^2}{\lambda^5} \frac{1}{e^{\frac{hc}{\lambda kT}} - 1} \tag{2.3}$$

In general astronomical bodies emit their light **isotropically**. This means they emit equally in all directions. There is a small handful of exceptions, such as pulsars and quasars (pulsars emit their radiation in beams instead). However, the majority of objects that you encounter will be isotropic radiators. How do we analyse isotropic emitters? Well, we can start by considering a sphere of radius r which surrounds an isotropic emitter. The amount of energy passing through any given area of the sphere will decline as the sphere gets larger; remember, energy is conserved, so as the sphere grows, the same amount of energy has to be spread over a larger and larger area. This means that the flux is directly proportional to the emitter's luminosity and inversely proportional to the square of the distance between emitter and observer. Hence, flux can be defined as Eq. 2.4.

$$F = \frac{L}{4\pi r^2} \tag{2.4}$$

This implies that the total amount of flux that a telescope's aperture can gather is dependent on the diameter of the telescope. A smaller one will gather less, a bigger one will gather more. If it helps, think of a telescope as acting for photons like a bucket does for rainfall.

The **instrumental flux** is specific to a given instrument and its set-up, as it depends on exposure times, aperture sizes and the optics being used. In order to turn an instrumental flux back into a physical one, corrections have to be made for all of these factors and also the relevant atmospheric parameters under which the observation occurred. We will return to this in the chapter on photometry.

2.4 The Magnitude Scale

Instruments typically return their raw observing results in the form of counts per second – that is, the total number of photon strikes on the receiving area divided by the time taken for the observation. (There are some exceptions – radio astronomy uses a non-SI unit called the Jansky, which we will discuss later.) However, in the vast majority of cases you will be using so-called magnitudes. The magnitude scale is a logarithmic one that has been in use since antiquity.

In the Nineteenth Century, the magnitude system was quantified and put onto a rigorous, empirical footing. It was discovered that the scale is actually logarithmic, and the scale was calibrated so that a the difference of five magnitudes is equivalent to a difference of one hundred in the intensity of flux. Perhaps surprisingly, it turned out that for most stars, Hipparchus's assessments were actually roughly right.

In fact, we now know that Hipparchus's system was a bit off. The brightest star in the night sky, Sirius, has a magnitude of -1.46 and 15 other stars are brighter than magnitude 1. Also, the modern system has been extended to include the Sun and the planets. Venus can attain a magnitude of -4.5, the full Moon is around magnitude -13 and the Sun is around -26.7. Given that Hipparchus was working without telescopes or modern equipment, and, in fact, lived before the concept of negative numbers had been developed; we can probably forgive him for missing some details.

As to the logarithmic nature of the system, the most likely explanation for this is that the human eye scales its perceived response to incident light logarithmically rather than linearly. Certainly the eye can perceive enormously different levels of illumination without returning a sense of difference as great as that which quantitative methods show us are there. However, the whole relationship between physical stimulus and subjective perception remains one of active research in human physiology and psychology.

When the magnitude system was quantified, the British astronomer Norman Pogson introduced the equation that now bears his name. This equation relates the observed magnitudes of two objects to the fluxes they are emitting. It is shown below (Eq. 2.5)...

$$m_1 - m2 = -2.5log_{10}(F_1/F_2) \quad (2.5)$$

where m_1 and m_2 are magnitudes of two celestial bodies and F_1 and F_2 are their corresponding fluxes. Hence, the flux ratio of two stars on of magnitude 1 and one of magnitude 6 is 100/1.

The magnitude of an object is dependent not only on the physical characteristics of that object but also on its distance from the observer, the amount of material between the observer, the detector used and the transmission characteristics of the instrument, which is often determined by a filter. Filters are discussed at length in Sect. 2.5 but you should be aware that magnitudes should always be quoted as being in a particular filter band if they are not then it is assumed they are unfiltered, visual magnitude. Hence, a star or galaxy may be magnitude 10 in R band but magnitude 12 in B band. The magnitude of extended and therefore resolved objects is the integrated magnitude, i.e. the combined light over the whole surface of the object and is known as a **surface brightness**. You should, therefore, allow more exposure time for extended objects than you would for stellar objects of the same listed magnitude.

Most magnitude quotes are **apparent magnitudes**, the brightness as seen from the Earth. An import characteristic of a star is its **absolute magnitude**, The brightness of a star at a distance of 10 pc and hence relates directly to luminosity. As brightness drops of with the square of the distance, we can modify Eq. 2.5 to determine the absolute magnitude from apparent magnitude and distance.

$$M = m + 5 - 5log_{10}(d/pc) \quad (2.6)$$

where M is absolute magnitude, m is apparent magnitude and d is distance in parsecs.

A major challenge in galactic astronomy is the problem with interstellar reddening. Gas and, in particular, the dust between the object being observed and the observer, preferentially scatters blue light over red light. Hence, an object with more scattering appears redder. It would not be much of a problem if the amount of material were distributed uniformly throughout the galaxy (in fact it would be a major advantage if it were). However, it is not, so consequently we need to modify Eq. 2.7 to adjust for the reddening ...

$$M = m + 5 - 5log_{10}(d/pc) - A \quad (2.7)$$

where A is the reddening in magnitudes.

Two photometric calibration standards are in common use, Vega and AB. The Vega system assumes that the star Vega, α Lyra, is magnitude zero at all wave bands, this has the advantage that it simplified Pogson's equation to ...

$$m2 = -2.5log_{10}(F_1/F_2) \tag{2.8}$$

so that the magnitude of a star in any filter is just the log of the ratio of its flux over that of Vega (multiplied by −2.5).

The AB system, unlike the Vega system, is based on calibrated flux measurements. For completeness the definition of AB magnitude is...

$$m_{AB} = -\frac{5}{8}\log_{10}\left(\frac{f_v}{Jy}\right) + 8.9 \tag{2.9}$$

where f_v is the measured spectral flux density, the rate at which energy transfer through a surface as measured per second per unit area. In practice, AB and Vega are almost identically at least for objects with reasonably flat spectral energy distributions. For the purposes of the practicals in this book, we will always use Vega magnitudes. However, you should always note in your lab book, which system you are referring to.

Keep in mind that the flux you measure is the instrumental flux. This means when you apply Pogson's formula or the various derivatives to that flux, what you will get is an instrumental magnitude. If you wish to determine an actual physical magnitude, then you will need to undertake a much lengthier process, one that accounts for the optics and the specific filter set that you are using.

Another subtlety concerns extended objects such as galaxies and nebulae. The magnitude that is usually quoted for these is the integrated magnitude, this being the total amount of light emitted by the object across the band in question. However, this causes an observational difficulty. If you set your exposure time for such an object based on its integrated magnitude, it would be underexposed as the light is spread out over more pixels than a star. A workaround for this difficulty is to use the surface brightness, which is the magnitude per square arcsecond. (Effectively this is a magnitude density, if that makes sense.) The surface brightness has some issues of its own in that it too has inbuilt assumptions – it treats objects as having even brightness when, in fact, most of them will have concentrated brightness in one area and more diffuse brightness elsewhere. It's best to keep in mind what sort of object you're observing and to try to adjust your practise accordingly.

Later in this book we will discuss in-depth the concept of **photometry**, the measuring of the amount of light received from an astronomical object over a given time with a given aperture and converting that into magnitudes. Within photometry there is an important concept, that of the **Zero Point**, where the zero point is the magnitude of an object that, for the specified instrument set up, will produce one count per second. More typically it is expressed as a flux or count so that the magnitude of an object is...

$$m = -2.5\log_{10}\left(\frac{F}{F_{zp}}\right) \tag{2.10}$$

or if the zero point is in magnitudes...

$$m = -2.5\log_{10}(F) - m_{zp} \tag{2.11}$$

2.5 Filters

Optical astronomical cameras are sensitive to radiation from the Ultraviolet to the Infra-red and are unable to distinguish between, for example, a red photon and a blue one. Domestic digital cameras produce colour images so how is this done? If we take a domestic camera apart and look at the light-sensitive chip, the **Charged Coupled Device** or **CCD**, you will see a coloured grid over the surface of the CCD made of a series of green, red and blue squares. Looking closely, you will see twice as many green squares as red and blue. Each coloured square covers a pixel and only lets the light of the colour corresponding to the colour of the square pass through to the pixel. The process of only letting part of the spectrum to pass through is known as **filtering** and the optical device that does this is a **filter**. The camera's firmware joins together four pixels (one blue one red and two green) to form one colour pixel. The reason there is twice as much green is due to the need to make the picture match what you perceive. The human eye is much more sensitive to green than red and blue, unlike the CCD, so twice as many green pixels are used to allow for this.

Looking at the CCD of most astronomical camera, we would see no such grid (some exceptions exist, but these cameras are not considered research quality and are aimed at amateur astronomers). This is because we wish to be able to change what range of wavelengths we want to fall onto out CCD and need to control precisely the characteristics of the filter in use. Hence, rather than filter individual pixels, we filter the whole array. Typically multiple filters are held in a **filter wheel** mounted between the camera and the telescope. Issuing a computer command causes the wheel to rotate so that the correct filter is in the optical pathway. Filters come in a range of sizes and can be either circular or square mounted or unmounted.

Filters are divided into two classes, **broadband** which are transparent over a wide range of wavelengths and **narrow band** that only lets light from a specific spectral line to pass, allowing for a small degree of movement within the line due to, for example, gravitational reddening. Each filter type has a set standard so that observations at different observatories using the same filters will produce the same result. Typically broadband filters will come in a set with each member in a set being assigned a letter that is known as the **Band**. Hence V band (V for visual), covers a region of the spectra from 511 to 595 nm and R band covers 520 to 796 nm. However, you should aware that filter names are an indication of the range they cover and not their profile with that region. The exact profile is dependent on the photometric system being used. Also, filter names are case sensitive, so a V-band filter is not the same as a v-band filter. Some filters names also have a prime symbol associated so a V-band filter is not the same as v́ and v is not the same as v́, although, as they should cover the same region of the spectra they will broadly be similar but you cannot for the purposes of science use, for example, an R filter when an r filter is needed, without requiring a complicated filter transformation calculation.

A filter set in wide use in the astronomy community is the Johnson-Morgan UVB and its extended form the Johnson-Cousin set of filters UBRVI (and the very similar Bessel set) which were designed so that CCD's colour response was similar to photographic film. The Johnson-Cousin set consists of a blue (B) filter, a red (R) filter, a green (V) filter, an infrared (I) filter and ultraviolet (U) filter. As almost all imaging is now done with CCDs, the original purpose of the Johnson Cousin set has been eroded and its drawbacks, for example, some filters overlap, is making it less attractive for astronomy. However, there is a very large body of work that uses *UBVRI*. Observatories are now moving to Sloan filters that are known as ú, ǵ, ŕ,í. Sloan filters are an improvement over UBRVI as they do not overlap and have flatter transmission curves. However, you should be aware that there are multiple "standards", 200 plus at the last count, with many telescopes, both ground-based and space-based using variations of, for example, the Johnson-Cousin set. Within this book this difference should not be a problem as you should be using the same filter set all the time, it may cause problems if you are using data from elsewhere. A number of good websites show the transmission curves for individual filters (Fig. 2.1), and it is possible to convert from one system to another, especially if the spectral class of the target is known. However, this can be time-consuming and may introduce additional uncertainties into your results.

You might also encounter RGB filters, which shouldn't be confused with Johnson-Cousin R and B filters, as RGB are imaging filters and are non-photometric, in many ways they are the same filters are we see on colour CCD. For science purposes, they should be avoided as there is no RGB standard, and your results will be difficult to reproduce. The RGB sets are the imaging filters, often used by amateur astronomers to produce colour images although similar role can be undertaken using the Johnson-Cousin RVB filters. However, as you might expect, an RGB filter set is considerably less expensive to purchase than an RVB set.

Narrow band filters tend to have a transmission window only a few nanometers wide and in some cases, such as a solar H-α, which is much narrower than an astronomical H-α maybe less than a tenth of a nanometre wide.[1] Most narrow band filters are centred on a forbidden line, which should be denoted by putting the emission line is square brackets, hence OIII should be noted as [OIII] and SII and [SII]. However, H-α is not a forbidden line, although it is a narrow band filter. As was mentioned above, in astrophysics forbidden lines are emission lines that cannot be produced in the laboratory. The mechanism need to produce the excited state has a low probability of occurrence and tends to be collisionally suppressed in all but the very lowest densities, which although common in deep space, are all but impossible to achieve in a terrestrial setting. When the [OIII] line was first detected in planetary nebulae, it was thought to be an emission line from a new element, Nebulium. Forbidden line emission is an important component of cooling of low-density astronomical environments such as planetary nebulae; hence emission line filters are an important tool in observing such objects.

Filters form part of the optical pathway so changing the filter will change the optical characteristics of that pathway. To avoid refocusing every time an image is taken a set of filters should have the same refractive index, so that within a set the

[1]Only use specially designed solar H-α filters to observe the Sun, standard H-α filters are **not** safe.

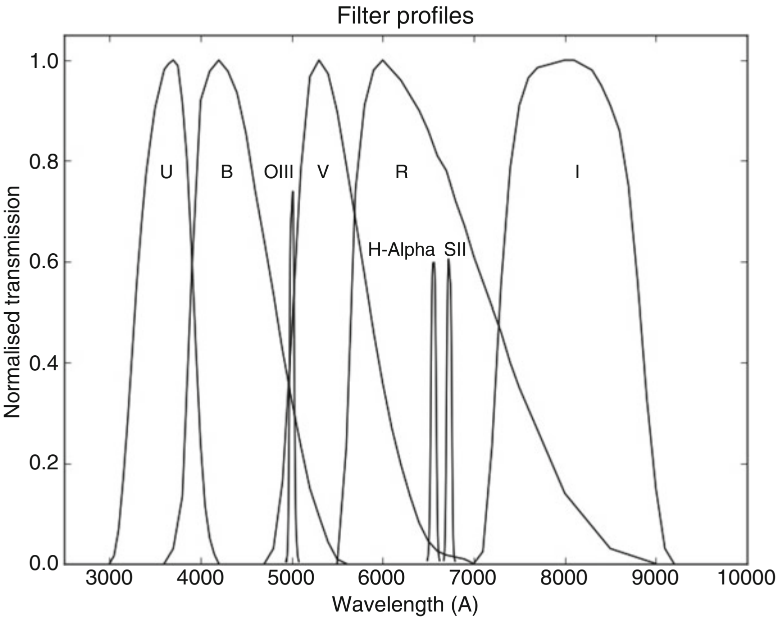

Fig. 2.1 Plot of the Johnson-Cousin U,B,V,R and I filter profiles along with the ESO La Silla [OIII], Hydrogen-α and [SII] profiles

same glass is used but it is differently doped to achieve the correct transmission characteristics. A filter set may include a clear, a transparent filter with the same characteristics as the coloured filters. If a blank space on the filter wheel was used instead of a clear the telescope would have to refocus when changing to the clear and then again when we move away from the clear hence the need to have a physical clear filter.

Other filters are manufactured that may be encountered, which are mostly used for visual observations, but may also be used in conjunction with a CCD. **Light Pollution** filters are designed to suppress emission lines associated with artificial lighting especially the characteristic sodium discharge lights. **Polarisation filters**

Johnson			Sloan		
Filter	Centre nm	Width nm	Filter	Centre nm	Width nm
U	365	70	ú	353	63
B	440	100	ǵ	486	153
V	550	90	ŕ	626	134
R	700	220	í	767	133
I	900	240	Other		
J	1,250	240	Y	1,020	120
H	1,650	400	L	3,450	724
K	2,200	600	M	4,750	460

may be used to examine the polarity characteristics of a light source although they are more commonly used in pairs to reduce the amount of light received from a bright object, such as the Moon. This light dimming can also be achieved by the used of a **Neutral Density Filter** or NDF, which uniformly reduces the amount of light received. Other filters exist to aid in the visual observation of emission nebula (Nebula filters) and to enhance planetary features.

2.6 Colour

For humans the perception of colour is a mixture of inputs from the three forms of cones in the retina as well as some complex image processing in the visual cortex of the brain which is highly sensitive to the language and culture of the individual. This means that in some cases two people, with perfect colour vision, will not be able to agree on a colour. In astronomy, colour has a very specific and very important meaning. It is the ratio of the observed fluxes in two filters (or magnitude A minus Magnitude B).

As discussed earlier, apparent magnitudes are dependent on distance from the observer but colour is distance independent but dependent on spectral class. Hence plotting a colour say V-B against a magnitude say B to produce a colour-magnitude diagram (CMD) provides important information about stars the their mass and their evolution, In fact this very plot was undertaken by Ejnar Hertzsprung and Henry Norris Russell in 1910. The **Hertzsprung-Russell (HR) diagram** is one of the most important plots in astrophysics; it enables astrophysics to identify the key stage of stellar evolution, the **main sequence**, a dense band of stars running from top left to bottom right during which stars are undergoing core hydrogen fusion. This observation coupled with the observation that as stars evolve they turn off the main sequence enables astronomers to age star clusters using the colour-magnitude diagram by looking for the main sequence turn off in the HR diagram.

Figure 2.2 shows a B-V against B colour-magnitude diagram of the open cluster M38. This plot contains a large number of stars that are either foreground or background objects and are, therefore, not part of the cluster. Likewise due to the large distance to M38 interstellar dust has preferentially scattered blue light over red light (a process known as **reddening**) resulting in the B-V colour increasing.[2]

[2]Remember low magnitudes are increasingly brighter so if there is less light at the blue end the blue magnitude increases in value and hence so does the B-V colour.

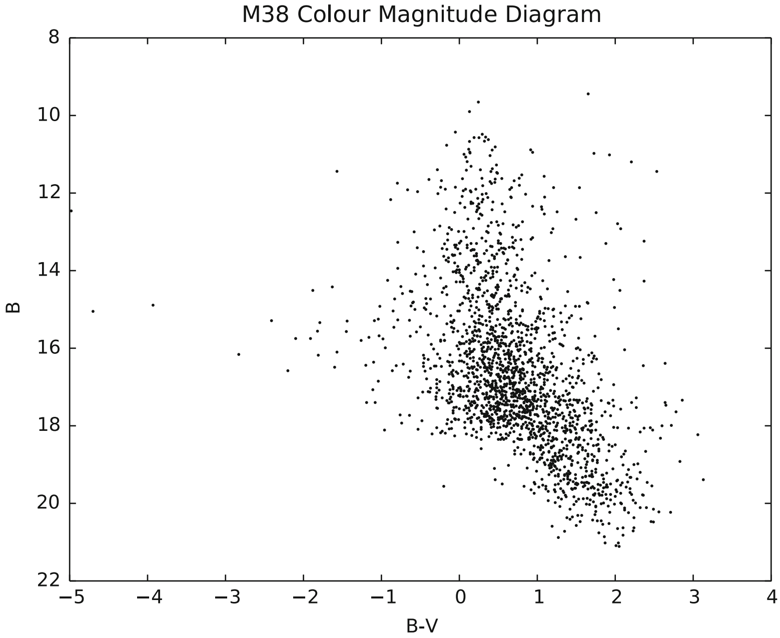

Fig. 2.2 A B-V to B colour magnitude of the Open cluster M38 using the USNO Nomad database

2.7 Polarisation

As stated previously light is an electromagnetic wave consisting of an electric field perpendicular to a magnetic field both of which are perpendicular to the direction of travel. In general, the orientation of the electric field is random. However, certain physical processes can cause the electric field to be inclined at a preferential angle to the direction of movement if this is the case the light is **linearly polarised**. Additional the electric and magnetic field may rotate with the axis of rotation being the direction of travel. Hence, in this case, the light appears to draw out a helix. We can characterise a wave as being with left-hand or right-handed depending on the direction rotation from the position of the observer. Again in general the distribution of the fields' handedness is random but physical processes can cause light to exhibit preferential left or right-hand rotation. In this case, it is **circularly polarised**.

Polarisation plays an important part in astronomy, with linear polarisation playing an important part in the study of interstellar dust and planetary atmospheres, whilst circular polarisation is a key investigative tool in understanding the nature of very large scale magnetic field especially in radio astronomy. In general, besides observations of the Sun, work with polarisation is beyond the capability of the small telescopes as the degree of polarisation from a source can be very small.

Chapter 3
The Telescope

Abstract The telescope is the principle tool of the astronomer. Therefore, it is important to understanding the type of telescope, the process by which the telescope is mounted, its limitations and the accessories available to the astronomer. Furthermore, on occasions telescopes can suffer from optical problems we show how to identify these problems and, if possible, how to resolve them.

3.1 Telescopes

Before starting this chapter, we need to issue you a small but necessary warning. Under no circumstance look at the Sun with any optical instrument or even with the naked eye unless it is by projection or with a special solar filter. If you use a filter, I strongly suggest you put a camera on to check the filter is not damaged before you put your eye to it. You can easily replace a camera, replacing your eye is not so easy. To illustrate the damage, some years ago the author was using a solar telescope mounted on a much larger telescope. Obviously I put the lens cap on the main telescope as pointing it at the Sun without the lens cap would destroy the very costly astronomical camera mounted on it. Unfortunately, I forgot about the small 5 cm finder scope mounted next to (and hidden by) the solar telescope. In the 2 s between pointing towards the Sun and realising that the lens cap was off and putting it on, the steel cross hairs in the finder scope melted. Fortunately, finder scopes are cheap but it was a lesson learned, so take heed, always double check.

A telescope has two functions, to gather light and to bring that light to a focus. Two ways exist that we can do this; we can refract the light using lenses or we can reflect the light using mirrors. You will be mostly using a derivative of a reflecting telescope, although your finder telescope and guide telescope, if you have one, will normally be a refracting telescope. Likewise if you use a solar telescope this is likely to be a refracting telescope.

Independent of the type of telescope it will have a number of important optical characteristics that you should be aware of:-

- The **aperture** size is the diameter of the telescope and is a measure of the light-gathering power of a telescope. Several techniques in telescope design may mean the optical aperture is in effect smaller than the physical aperture.

M. Gallaway, *An Introduction to Observational Astrophysics*, Undergraduate Lecture Notes in Physics, DOI 10.1007/978-3-319-23377-2_3

- The **focal length**, (f), is the distance between their first optical component and the first point of focus.
- The **Field of View** (FOV) the amount of sky visible through the instrument, measured in degrees or arcminutes is the amount of sky visible through the instrument, measured in degrees or arcminutes. The angular field of view of a telescope (in radians) is the aperture of the eyepiece (the field stop) or the diameter of the detector f_o divided by the focal length f.

$$fov = \frac{f_o}{f} \times 206{,}265 \tag{3.1}$$

- **Magnification** is the ratio of the angular size when using the telescope and when not using it. For unresolved objects such as stars, the magnification has little of no effect. For smaller, consumer telescopes, high magnifications are often quote. In practice high magnifications are rarely used in astronomy and generally cause more problems than they are worth. If you are using an eyepiece the magnification is simply the ratio of the focal length of the telescope, f and the focal length of the eyepiece, f_e; if however you have an astronomical camera at the focus then the nearest approximation to magnification is the image or plate scale. This is just the focal length of the telescope, f over the diagonal diameter of the light sensitive chip in the camera (the CCD).

$$Mag = \frac{f_e}{f} \quad Plate\ Scale = \frac{f}{f_o} \tag{3.2}$$

- Another often quoted feature of telescopes is the f number. This is simply the focal length f over the optical aperture size, A. In general telescope with low f numbers are called *fast* because they gather light quickly due to their large field of view. However, the downside is that they have reduced resolution. Additionally, optics for fast telescopes tend to be very expensive to manufacture when compared to telescopes with high f numbers.

$$F = \frac{f}{A} \tag{3.3}$$

- The **angular resolution** of the telescope (α_c) is the telescope's ability to separate two objects and is given by the average wavelength of the light being observed divided by the aperture of the telescope D_0. This is the diffraction limit caused by the formation of the Airy disc, which is produced by the diffraction of light within the optics.

$$\alpha_c'' = \frac{1.22\lambda}{D_o} \times 206{,}256 \tag{3.4}$$

Although when trying to resolve two stars the Rayleigh criterion Eq. 3.5, which gives the resolution in radians, is more representative, although the simpler Dawes' limit is also widely used (D is in millimetres).

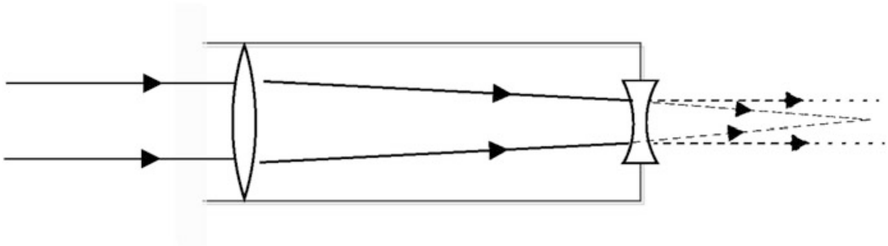

Fig. 3.1 Diagram showing the optical pathway of a Galilean telescope

$$\alpha = sin^{-1}\frac{1.22\lambda}{D} \tag{3.5}$$

$$\alpha \cong \frac{115.82}{D} \tag{3.6}$$

In reality, most telescopes are not diffraction limited but are limited by atmospheric turbulence, **seeing limited**.

The first refracting telescope were made in the early sixteen hundreds in the Netherlands by spectacle makers but it was not until Galileo Galilee designed and constructed his own a telescope only a few years after their first appearance, that the telescope became used for astronomy. Galileo's telescope used a convergent objective lens and a divergent eyepiece which, and unlike modern instruments produced an upright image.[1] Galileo's design was rapidly improved upon by Kepler who change the divergent eyepiece to a convexed one and inverted the image (Fig. 3.1).

The Keplerian telescope allowed the use of increased magnification but failed to overcome two of the refracting telescopes principle problems. The first of which is an engineering problem, as the objective lens becomes increasing large its mass significantly increases. Not only does this make the telescope unwieldy to use but the lens becomes difficult to mount due to the thinness of the glass at the edge and tends to sag under gravity, resulting in changing optical characteristics with elevation. Additionally, as the glass volume increases, it becomes increasingly more difficult to keep the lens free of imperfections (Fig. 3.2).

You will recall that a lens's refractive index is wavelength specific. This being the case a lens will bring each colour to a differing focus resulting in coloured rings round objects, **chromatic aberration**. This can be overcome in part by using much longer focal lengths. However it was not until the introduction of **Achromatic**, and later **Apochromatic** lenses that the problem was resolved. Achromatic lenses uses two types of glass with differing refractive indexes bonded together to form a lens with the same refractive indexes for red and blue light. Apochromatic lenses use

[1] Most modern telescopes invert the image.

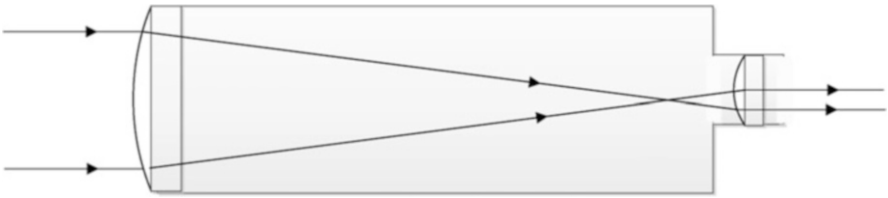

Fig. 3.2 Diagram showing the optical pathway of a Keplerian telescope

the same concept but have three components for red, blue and green. The problem with mass has never really been resolved and by the start of the twentieth Century refractors reach (and possible even exceeded) their maximum operational size. Over the years, additional optical coating have been applied to astronomical lenses and although this has improved them it has not overcome their physical limitations. Obviously telescopes are exposed to the elements, although hopefully not very extreme weather, but over time lens become pitted, coating become worn and the optical bonding compound used to bond achromatic and apochromatic lenses loses its optical transparency; at which point the lens will either need maintenance or replacing.

The concept of using mirrors to magnify an image had been known for at least five hundred years before their successful application to astronomy by Newton. The use of mirrors has a number of advantages over the use of lenses; only the surface of the mirror needs to be optically smooth reducing the need to produce optically perfect glass, mirror based telescopes do not suffer from chromatic aberration and their position at the base of the telescope makes them easier to operate. However, they are not entirely without design problems. Typically a secondary mirror is needed near the primary focus which causes an obstruction and reduces the efficiency of the telescope. Also, light is diffracted around the secondary supports and the secondary itself which causes the spikes and halos seen around some stars in astronomical images, although admittedly in some cases these might be added digitally for effect. There is also a tendency for light incident to the edge of the mirror to be brought to a slightly different focus to those at the centre, as a result objects become smeared out, a problem known as **coma**. An additional problem is the alignment of the primary and the secondary. If this alignment is not within a very small tolerance (Collimated) the light arrives at the focus at a slight angle to the focal plane and consequently point sources become elongated.[2] **Collimation** problems normally worsen over time and but are easy to address. Hence re-collimation is part of the normal maintenance procedure for a telescope.

As with lenses, mirrors become increasing massive as their diameters increase, sagging under gravity is prevented as a mirror can be supported along its entire rear surface area. However, it appears that the engineer limit for producing astronomical

[2]They are often described as looking comet like.

grade mirrors has been reached around 8 m. A number of 10 m class telescopes exist beyond this limit but these have segmented mirrors, that are manufactured and individual components then placed together, a technique not itself without problems. The planned 30 m+ class telescopes currently under construction, including the appropriately named 39 m European Extremely Large Telescope (E-ELT), all have segmented mirrors.

A wide range of designs for reflecting telescopes exist, most of which differ only in the path the light travels. The simplest and perhaps the least useful is the Prime Focus telescope. This has no secondary mirror rather the detector is placed at the prime focus. Although of little use for observing with the eye, the introduction of the camera to astronomy and, in particular, the digital camera has made this design more practicable, in fact many parabolic dish radio telescopes are of this design.

Newton's original design is still in wide use around the world, particular with amateur astronomers due to its low cost and ease of operations (Fig. 3.3). The Newtonian uses a flat mirror set at 45° to the primary mirror and just inside the focal point so that the point of focus is outside the tube, often close to the top. This allows for the easy mounting of a focuser and is normally a comfortable position of the observer. However, when mounting an instrument the focal point of a Newtonian is problematic as the additional mass of the instrument so far up the tube makes balancing difficult, although not impossible and for very large telescopes the focal point become increasingly inaccessible. Consequently, most professional instruments have move to a Cassegrain like design, although a small number of operational instruments, retain the ability to switch between Newtonian and Cassegrain operation.

The Cassegrain telescope uses a hyperbolic secondary mirror placed parallel to the parabolic primary which reflects the light down the tube and through a hole in the centre of the primary with the focal plane laying just outside of the tube at its rear. Hence, Cassegrain telescopes are viewed through the bottom, very much like a traditional refractor which allows the positioning of the instrument in a more convenient location, it also means that Cassegrain tend to be physically shorter than a Newtonian of the same focal length. By modifying the Cassegrain by replacing the

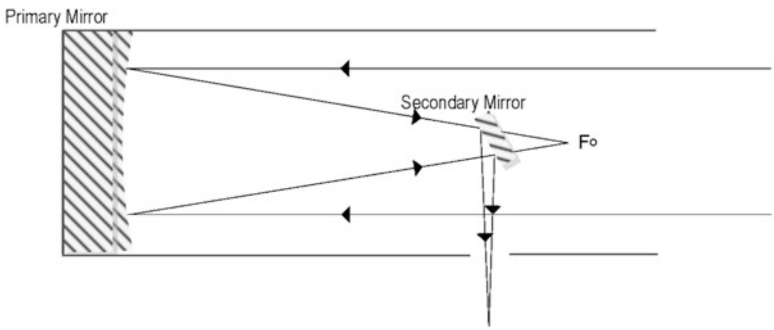

Fig. 3.3 Diagram showing the optical pathway of a Newtonian telescope

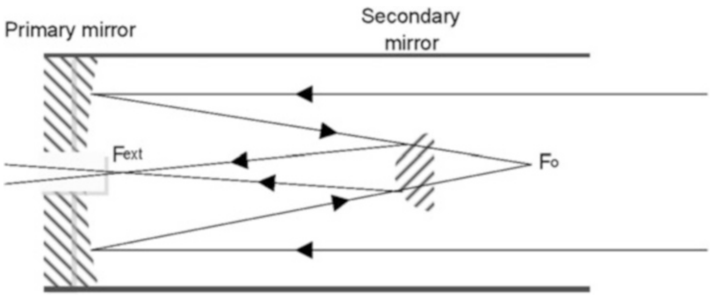

Fig. 3.4 Diagram showing the optical pathway of a Cassegrain telescope

primary with a hyperbolic mirror, suitable matched to the secondary, George Willis Ritchey and Henri Chreťien introduced a near coma-free and spherical aberration free telescope design that now bears their name and perhaps, more importantly, is the most common professional telescope designs now in use, including the four 8.2 m Very Large Telescopes in Chile, the twin 10 m Keck telescopes in Hawaii, the Hubble Space Telescope and the Spitzer Space Telescope (Fig. 3.4).

The telescopes you will be using as an undergraduate may well be neither a reflector nor a refractor but rather a hybrid of the two, a **Catadioptric** telescope. Catadioptric telescopes use a correcting lens (known as the correcting plate) in front of the primary mirror in order to improve optical imperfections and to reduce the physical focal length of the telescope whilst retaining a large actual focal length. This also allows the use of a spherical primary and secondary. Spherical mirrors are much easier, and as a consequences much cheaper, to manufacture. Most Catadioptric telescopes have a full diameter correcting plate and are known as **Schmidt Cassegrain** telescopes or SCTs. The positioning of a full width correcting plate at the front of the instrument allowed the introduction of a further simplification, the **Maksutov-Cassegrain** telescope, where the secondary is actually the curved mirrored surface of the correcting plates. This design allows for low cost construction and maintenance as the secondary is securely fixed and the tube sealed, reducing wear on the optical surfaces. Given the reality low purchase cost of these designs, their intrinsically low maintenance and the fact they are easy to handle due to their compact design means they are very popular with educational facilities and the more serious amateurs (Fig. 3.5).

An important point about the condition of optics. If you think a telescope has dirty optics or is suffering from condensation, NEVER, wipe them clean. In fact never clear them yourself at all. If you wipe them, small dust particles will scratch the optical surfaces. Your observatory staff will know how to clean the optics and will have specialist tools and chemicals to do so, hence it's best to leave it to them. If you are using a SCT and you are condensating you might find that slowly warming the correcting plate up with a hair dryer may solve the problem, but do not hold the hair dryer too close. If this doesn't work or is not effective the humidity is going to be too high for good observations, in which case it is normally worth shutting the telescope down until conditions improve.

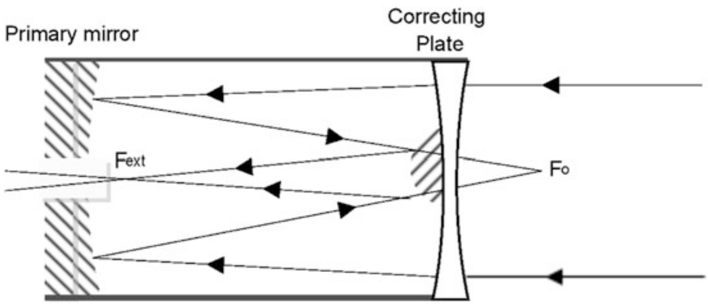

Fig. 3.5 Diagram showing the optical pathway of a Schmidt–Cassegrain telescope

As with refracting telescopes, reflecting telescopes suffer from optical degradation due to weather (in fact due to just being exposed to the air), over time the thin silver or aluminium coating on the mirror degrades and will eventually have to be replaced – re-silvering. This is a relatively easy process (depending on the size and putting the mirror back is not as easy as taking it out.) but requires a vacuum chamber, although many large professional observatories have silvering machinery on site.

3.2 Mounts

Just as important as the optics is the structure the telescope is fixed to which allows it to rotate. If the structure is not stable enough, finding an object and keeping it in the field of view becomes extremely difficult and is perhaps the biggest problem with low cost, toy telescopes, which are really only good for the Moon and the brighter planets. A wide array of mounts are currently use from, Dobsonian (which tends to be shortened to just Dob), German Equatorials, GOTO mounts, English Equatorials, Yoke mounts etc., although effectively only two designs are in use.

Initially, the first mounts were of an Alt Az design wherein the telescope is moved in a vertical (Alt) axis and in a horizontal (Az) axis. This design is cost effective and easy to use however it does have a major draw back. Astronomical objects move across the sky in an arc, caused by the Earth's rotation and it's inclination to the plane of the solar system (more on this later). The effect of this is that in order to track an object with an Alt-Az mount it needs to drive both axes and in a non-linear fashion, if it does not the object soon drifts out of the field of view. Such tracking is challenging to do by hand and is extremely difficult to do mechanically without the use of a computer. A camera fixed to a telescope using an Alt-Az mount rotates as the mount tracks which can cause problems with longer exposures. There is a single solution, however, the equatorial mount. If you tilt the mount so that the centre of rotation of the horizontal axis now points to the celestial pole and the vertical axis is perpendicular to the inclined horizontal axis, the telescope naturally follows the

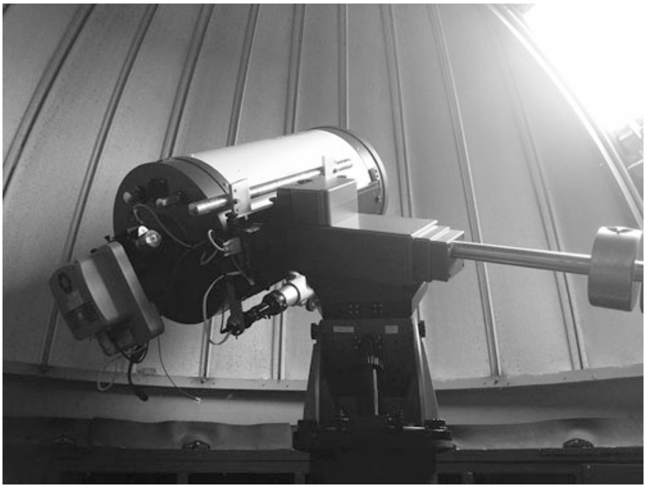

Fig. 3.6 Image of a telescope mounted on a German Equatorial mount (Image courtesy University of Hertfordshire)

same arc as stars and the object do not drift in the other axis at all. This orientation is known as polar alignment and means that only we need to drive only a single axis, which can be done simply with clockwork. Multiple designs of equatorial mounts exist, the German Equatorial (Fig. 3.6) being one of the more popular.

Equatorial mounts tend to be heavy due the amount of counterweighting needed and need a larger dome to house them than Alt Az telescopes. They also suffer from what is known as **Meridian Flip**. Meridian Flip occurs, not unsurprisingly, when the telescope transits the meridian. In order to stop the telescope hitting the mount the telescope must be re-orientated or flipped. This means that if you are performing a series of observations of an object, there will be a break as the object crosses the median, the very best place to image.

With the introduction of small, cheap, powerful computers, and field rotators, that rotate the camera so that it keeps the same position relative to the field the limitations of the Alt-Az mount have now been overcome. Consequently almost all modern large telescopes are of the Alt Az design, although equatorial mounts are still common.

Perhaps one of the least common telescope mounts that you are still likely to encounter is the transit telescope. Transit telescopes point only at the meridian (the line that runs north to south) and therefore only move in elevation. Transit telescopes, which tend to have very long focal lengths, were used to determine when an object crossed the local meridian and its elevation, which enables the observer to either determine the local time (and by doing do their position on the Earth) or determining the location of the object in the sky. Transit telescopes tend to be found in, older observatories and new constructions are rare. The Monument to the Great Fire of London, near to Pudding Lane, contains, somewhat bizarrely, a transit telescope.

One of the biggest changes in telescope mounts in recent decades is the introduction of GOTO telescopes and more recently robotic telescopes. GOTO telescopes feature powered mounts with encoders which feedback to a computer their position. If the telescope is portable it may feature a Global Positioning System (GPS) receiver to aid in set-up. Effectively a GOTO telescope always knows where it is pointing, although some degree of set-up is needed which normally consists of manually pointing the telescope at three bright stars so the computer can work out the translation from encoder position to actual sky position. Once a GOTO telescope is set-up for the night, finding a target is just a case of picking a target from a list held on the computer or telescope handset and pressing a button. Although you may occasionally need to re-align the telescope during the night as position errors do creep in. Gone are the days of finding a bright star first and then using a star map to hop between stars to find a faint target.

Robotic telescopes take the GOTO telescope concept further. With robotic telescopes the computer controls the dome, the telescope, the camera, the auto guider and the focuser. It is fed information about weather conditions by a weather station and is controlled by a set of instructions or plans, typically written in a markup language very similar to HTML, called RTML. Although more costly to implement and to maintain than a traditional telescope, robotic telescopes do not need an operator and can take advantage of changeable weather conditions. Examples of robotic telescopes are the Open University's Pirate instrument in Majorca and the University of Hertfordshire's Chris Kitchin and Robert Priddy telescopes which are both located in the UK.

3.3 Eyepieces and Additional Optics

Although you will be mostly observing objects using a camera or spectrograph there will be occasions where you will want to observe an object with the naked eye. In which case you will need an eyepiece. Eyepieces come in a number or optical designs, barrel widths, focal lengths and eye relief. The design of the eyepiece will affect the focal, length (and hence the magnification), field of view and the eye relief of the optical pathway. The barrel width is the focuser diameter the eyepiece is designed to fit, for historical reasons these tend to be sold in inches. Standard sizes are 1.25″, 1.5″ and 2.0″. Adapters are commercially available to allow a smaller eyepiece to fit a larger focuser.

Eye relief is the distance between the last optical component of an eyepiece at the point at which the naked eye observer can obtain the full viewing angle. Too small an eye relief will cause the user's eyelashes to touch the eyepiece which is both distracting and uncomfortable. Likewise too long an eye relief is also uncomfortable. 15 mm is about the ideal level of eye relief, especially for spectacle wearers. If you are a lone observer and wear spectacles, it is generally best to take the glasses off and adjust the focus to correct for the loss of the spectacles.

Fig. 3.7 An array of eyepieces, with both one and 1.5 inch barrels

The first eyepieces used by Galileo and Kepler were single pieces. The introduction of eyepieces made of multiple optical components were first introduced by Christiaan Huygens. Huygens eyepieces have now been largely replaced by more modern designed. However the Huygens design does not have any optical bonding between the optics. Optical bonding degrades over time and in particular when exposed to very bright light. Hence when observing the Sun via the projection method as discussed in Chap. 12 Huygens eyepieces hold a distinct advantage. The Ramsden eyepiece is no longer in wide use and like the Huygens it is a two piece lens but in this case the lenses are reversed with reduces chromatic aberration. The Kellner is a improvement on the Ramsden design in that it significantly reduces aberration, however in recent years the Kellner has been largely replaced in the marketplace by the Plössls and most mid-range amateur telescopes tend to come with at least one Plössls. However, short focal length Plössls' tend to have poor eye relief (Fig. 3.7).

You may wish to undertake an observation that is not achievable with the current optical set-up, requires the adjustment of the optical pathway for physical reasons, or to correct an optical defect. The **Diagonal** or **Star Diagonal** changes the orientation of the optical pathway by 90°.[3] This is most often done to accommodate and instrument or to make visual observations more comfortable. The change in pathway is performed by the use of either a mirror inclined at 45° or by the use of a prism. Mirror based diagonal have the additional advantaged that they right the inverted image, although the mirror, will deteriorate and therefore they have a more limited life than prism based diagonals. Modern prism diagonals now allow the transmission of more light than mirror based diagonal. They can introduce such as chromatic aberration but in general this does not occur when used for long focal length instruments such as the SCT mostly used by educationally facilities (Fig. 3.8).

On some occasions the object you wish to observe is resolvable but the image size is very small. When using a telescope visually you can generally use a shorter

[3] 45° diagonals are also available.

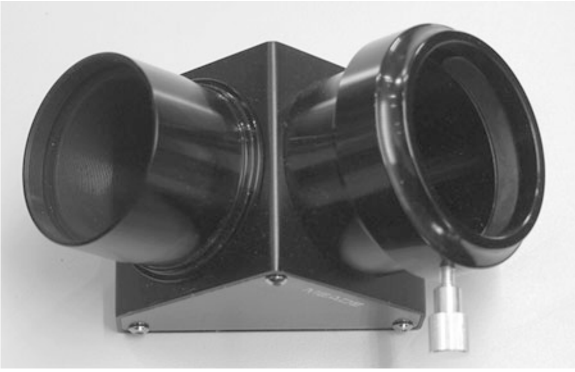

Fig. 3.8 A Star Diagonal with a two-inch barrel

focal length eyepiece, although this isn't always a comfortable choice and not a solution when using a camera. In order to overcome this problem, you may try to use a **Barlow**. A Barlow is an optical device that sits between the camera or eyepiece and the prime focus and effectively increasing the focal length of the telescope. However be aware that it might introduce aberration, increase exposure times and, as it extended the camera out further, balancing problems. A Barlow lens will typically double the focal length thereby doubling the image size.

Alternatively, you may have a target whose angular diameter is greater that of your telescope. In this case, you could take several images and montage them together. However, if this is not possible, you can use a **Focal Reducer**. This is somewhat like a Barlow and suffers from many of the same problems, but unlike a Barlow, the focal reducer, as the name suggests, reduced the effective focal length of the telescope. Remembering that $fov = \frac{f_o}{f} \times 206{,}265$ we can see that a focal reducer effectively increases your field of view.

The last item, in this in no way comprehensive list of items you may encounter between the camera/eye and the focuser, is the **flip mirror**. The flip mirror is, as its name suggest, a flat mirror which can be moved from a position where the mirror is parallel to the telescope to one where it is 45° with the telescope. When in the parallel position light travels straight through to the instrument or eyepiece in the normal mounting position. When flipped to the 45° position the light is diverted to another detected at a 90° angle to the primary detector. Hence you can have two instruments attached to the telescope at the same time. Typically this would be a spectrograph and eyepiece but not always.

3.4 Problems

A number of aberrations exist that appear in telescope optics. Some are caused by unavoidable parts of the design of the optics some by the incorrect calibration of the optics. When in focus a star, which is a point source, should have a Gaussian, or bell curve, like profile in both directions as the light from the star should be distributed circularly around the central pixel. If it is not, then more than likely it is a tracking error, however, it might be an indication of something more serious that needs addressing.

We have already discussed chromatic aberration, the appearance of coloured halo around sources, caused by refraction bringing differing wavelengths to differing focal points. As you will be using a reflector it is unlikely you will encounter this in all but cheap optical system, such as might be found in your telescope's finder scope.

In telescopes that use parabolic mirrors such as Newtonian an effect known as coma can sometimes be seen. Coma is a form of optical aberration which makes point like sources, such as stars, appear comet-like, in that they appear to have a tail, hence the name. Coma is caused by parabolic mirrors only bringing parallel rays to focus at the same point and may be more pronounced away from the centre of the field. Precise collimation of the optics may help reduce coma as may the introduction of additional optics. However in general coma is caused by the design of the optics. Telescopes using spherical mirrors, correcting optics, or both, such as Schmidt telescopes suffer from little or no coma.

Spherical aberration, which is caused by differing parts of the mirror bringing light to focus at differing points causing rings to appear around stars, is infamous for being the optical error that affected the Hubble Space Telescope (HST) shortly after launch. As this is purely a function of the optics it must be overcome by the design of the optical system. The easiest method is to use a parabolic mirror. Alternatively, if a parabolic mirror is not used, correcting optics, such as the corrector plate in a Schmidt can adjust the optical pathway so that there is no spherical aberration. This was the initial solution that saved the HST with the COSTAR optics package being placed in the optical pathway replacing the High Speed Photometer. Later, as the HST's cameras were replaced, correctors were build into the instruments and in 1993 COSTAR was removed and replaced by a new instrument.

The optical system of a telescope has two planes perpendicular to the light path. If the two planes are not correctly aligned there is a tendency for one plane to extend the image along one axis. Instead of getting a nice round star we get an ellipse or lenticular shaped star. This effect is known as **Astigmatism**. At first it may seem similar to coma however in astigmatism the light is spread equally along the axis whilst in coma it is not. Hence, for most science purposes astigmatism is preferable over coma. In some optical system, such as the Ritchey-Chrétien telescope a trade off is made between reduced coma and increased astigmatism. In general most telescope designs do not suffer from astigmatism and those that do can use additional optics to correct it, as happens in some Ritchey-Chrétien.

In order to produce the best optical image the optical components must be correctly aligned or **collimated** otherwise the image will be distorted. Over time and with use, most telescopes slowly lose collimation and have to be re-collimated. The easiest way to identify a telescope loosing collimation is to look at an out of focus star. This should be doughnut shaped with the dark disk making the hole in the doughnut; this hole being shadow of the secondary. The secondary should appear exactly in the center of the out of focused star, if it does not then the telescope is out of collimation. Newtonian telescopes, tend to lose collimation if they are not permanently mounted and are often moved. Fortunately they are straightforward easy to collimate with a laser collimator. Other designs such as the Schmitt Cassegrain are not. So if you suspect you have a collimation problem please see your observatory support staff.

The telescope you will be using is looking at the sky which is consider as a curved surface. Likewise, except for some systems that might have a flat secondary, our optical surfaces are also curved, however our imaging device is flat. This mismatch of a curved surface projected onto a flat surface leads to **field curvature**. This leads to the distortion of separation distances between objects within the field as well as non-uniform illumination of the imaging surface. This is a particular problem for fast telescopes that have a large field of view. Field curvature can be addressed with additional optics or in very rare cases, such as with the Kepler Space Telescope curving the detector surface. As we will see later in the chapter on CCD, certain processing techniques can also partly resolve field curvature.

Chapter 4
Time

Abstract The measurement of time was one of the first applications of astronomy and many of observational astronomy's underpinnings are related to time. Knowing the Julian date and local sidereal time are key in identifying what objects can be observed and when.

4.1 Introduction

Our modern lives are ruled by time; the tick tick tick of the clock, or more commonly now, the fast harmonic hum of a quartz crystal. We work to timetables and deadlines and the term working day often involves little actual daylight. However, for the vast part of human history and pre-history, when we were an entirely agrarian culture, time was just the changing of the seasons and the rising and setting of the Sun. When to eat, when to sleep, when to plant and when to harvest. Our ancestors noticed that the point on the horizon changed through the year, when the day was at its shortest the Sun rose at its most further point north and when the day was at its longest the Sun rose at its most further point south and twice a year, when the days were equal in length, the Sun between north and south. The first two events are called the Winter and Summer solstice and come from the Latin for Sun (sol) and standstill (sistere). The later are the spring and autumnal equinox from the Latin Equus (equal) and nox (night). Great significance was placed on these events and structures such as Stonehenge in Great Britain were erected to observe them. These structures became in effect, the first calendars.

Later, with the development of cities, military forces, organised religion and an elite ruling class the passage of time and its measurement became more important as you need to know when to pray, when to change the guard and when the tax collector was coming. Time was measured using Sundials, water clocks, sand timers, and marked candles and the year was often, but not always broken into 365 days with subdivisions, now known as months, based on the cycle of the Moon.

In 46BC the introduction of the Julian Calender by Julius Caesar as the official Roman Calendar resolved many of the problems of earlier systems of timekeeping and the span and influence of the Roman Empire spread the Julian Calender throughout what would become western civilisation.

M. Gallaway, *An Introduction to Observational Astrophysics*, Undergraduate Lecture Notes in Physics, DOI 10.1007/978-3-319-23377-2_4

By 1582, it was becoming clear that small discrepancies in the Julian Calender were causing the seasons to drift. This was resolved by the introduction of the Gregorian calendar, named after the Pope at the time, and a 10-day offset. The change encountered considerable resistance with some people believing they were losing 10 days of their lives! Surprisingly, it was not until the 1920s that the last country stopped using the Julian Calender although some non-Gregorian/Julian Calenders are still in use.

It is worth stopping for a brief while before we move from date to time to discuss BC and CE (also known as AD). The move to a CE dating system occurred before the introduction of the concept of Zero in the western world. The consequence of this is that 1 BC is followed by 1 CE. In general this should not cause you any problems unless you are using dates that pre-date the Christian Era. Likewise, you be aware that there is a 10-day difference between Julian and Gregorian dates. If you are using a pre-Gregorian date for some reason, you will need to convert.

By the fourteenth Century mechanical century clocks, used to mainly for religious reason and important cultural gatherings, were becoming increasingly common. The introduction of the pendulum by the Dutch astronomer and Christiaan Huygens and the hairspring by Robert Hooke, who coincidently also coined the word cell in relation to biology, led to clocks become more widespread, portable and affordable.

All clocks have inaccuracies and require resetting to a known time on occasions. Up until 1847 the vast majority of clocks were set at noon. However, the timing of noon is a function of local longitude hence noon in London occurs approximately 10 min before noon in Bristol, which lies 2°3” west of London. It was not until the coming of the railways, with the need for well-coordinated timetables over large areas, did the need for a centrally set time system become apparent. Ultimately this lead to the introduction of time zones and of **Greenwich Mean Time** (GMT) as the defacto zero point.

Typically zones are 1 h across and are measured from the time in Greenwich in the United Kingdom, **Greenwich Mean Time** (GMT) with each hour being 15° of latitude. Hence Western European Time is 1 h ahead of GMT and is GMT+0100 whilst Honduras is GMT-0600. Although since 1972 the standard has been **Coordinated Universal Time One** (UT1) the difference between GMT and UT1 is small enough (0.9 s) that for most purposes in astronomy it can be ignored. The introduction of daylight saving (British Summer Time in the UK) can somewhat complicates matters. Beware if you are not using GMT or a standard GMT time-zone related offset. Best practice in astronomy is not to use daylight saving. If you need to note down a time in your notes make sure that you use GMT or your local equivalent.

4.2 Solar Time

Johannes Kepler was a seventeenth-century astronomer and astrologer (at the time the two were indistinguishable) born in what is now modern Germany, who worked for some time with Tycho Brahe, one of the greatest observational astronomers of the time. Tycho was a colourful character who lost his nose in a duel over a mathematical problem (both sides turned out to be incorrect), had a pet elk, which died when it fell down a set of stairs whilst drunk, was one of the richest people in Europe and died in suspicious circumstances; potentially poisoned by his assistant over a woman. Claims that he died when his bladder burst when he refused to leave a banquet in order to relieve himself are likely to be untrue. Kepler used Tycho's excellent observations of the motion of the planets to formulate his three laws of planetary motion, although it wasn't until after Kepler's death that Newton's three laws and his law of Universal Gravitation lead to an understand of the origins of Kepler's laws.

Kepler stated that the planets travel in an ellipse around with the Sun at one of the foci, that a line drawn from the Sun to the planet will sweep out equal areas in equal time and that the square of the orbital period of the planet around the Sun is proportional to the cube of the length of its semi-major axis.

At this point, you may be wondering what this has to do with time. If we measured, over the course of a year, the time at which the Sun passes through the line is drawn from the northern cardinal point to the south cardinal point, the **Meridian** we find that this does not happen the same time every day. Some days it occurs before noon on our clock, sometimes after and on 4 days it will be close to the same time. The variation is between +15 and −15 min. The difference, between the time as measured by you, watch, **Mean Solar Time** and the time measured by the Sun, **Apparent Solar Time**, and the translational between the two is known as the **Equation of Time** (see Fig. 4.1).

There a two reasons for the difference between Mean Solar Time and Apparent Solar Time. Firstly let us consider Kepler's first and second laws. The first law tells us that are certain times during the year the distance between the Earth and the Sun decreases until it reaches closest approach, **perigee**, which occurs in January. It will then increase until it reaches its maximum distance **apogee** in June. From the second law, we know that the Earth must be going faster when it is near the Sun than when it is more distance. Hence, the Earth will travel a greater angular distance around the Sun in a day at apogee than at perigee and therefore during this time the Sun will move slightly further forward than the mean. Likewise, because the Earth is slightly closer to the Sun during perigee than apogee (and it is very small only 1.6 %) the apparent movement of the Sun caused by the Earth's orbit is larger.

Turning to the second cause. You will no doubt be aware that the Earth's axis is inclined by 23.5° to its orbit and that it is this that accounts for the seasons. The Sun tracks along the ecliptic every day with the peak of the curve increasing daily until

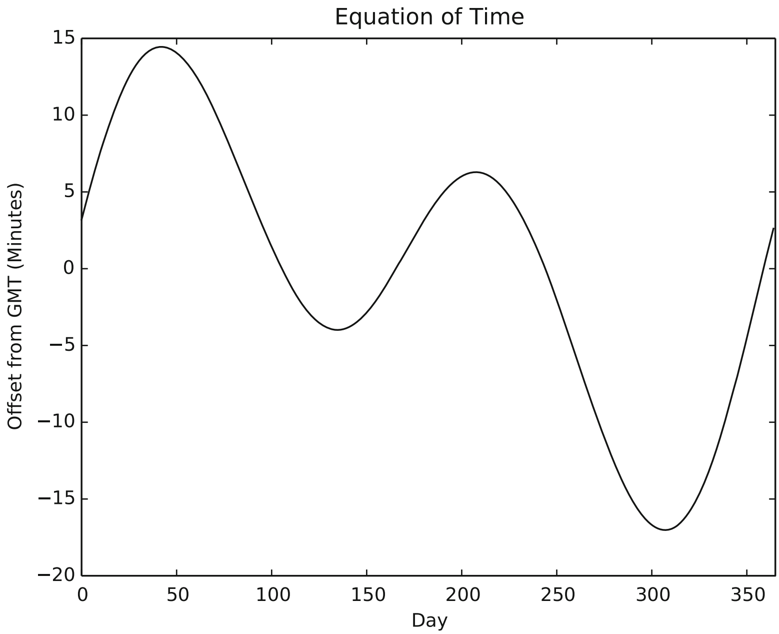

Fig. 4.1 Figure showing difference, over the course of a year, between 12:00 Mean Solar Time and the time the Sun cross the local meridian

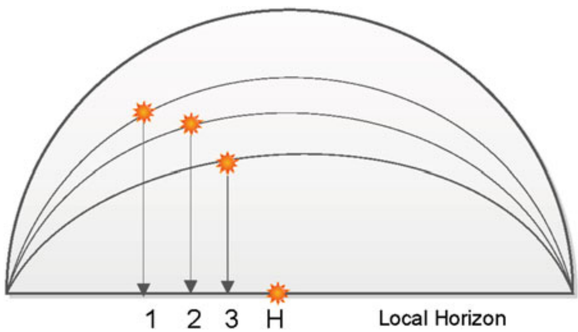

Fig. 4.2 Figure showing the effect of the inclination of the Earths orbit on the Equation of Time. Note the three arcs representing three different days, with the position of the Sun on each arc after a set time and that position projected onto the Horizon. As can been seen as we move towards the Summer Solstice angle at which the Sun climbs increasing and Apparent Solar time, as shown by the projection onto the horizon (*1*, *2*, and *3*) falls behind Mean Solar Time shown by *H* on the horizon

the summer solstice at which point it begins descending reaching its minimum at the winter solstice. Looking at Fig. 4.2 which projects the Suns position on the ecliptic onto the Equator. The Sun tracks across the sky at a set angular rate, governed by the rotation of the Earth. When it is near the summer equinox the height of the Sun at noon it at its highest. Hence the Sun has further to travel and, therefore, the position of the Sun, projected onto the horizon, is behind that of a Sun that would not have an inclined path.

It is the combination of these two effects, the eccentricity of the Earth's orbit and its inclination to the ecliptic that is responsible for the Equation of Time.

4.3 Julian Date

Astronomers do not use the standard Gregorian Calender. Instead a decimal, year independent, dating system known as the **Julian date** is used. The Julian date is the number of days since at noon on 1st January 4713 BC by the Julian Calender, as date chosen as it is the convergence of the Indiction, Metonic and Solar cycles an event that will not occur again until 3268 CE. Fractions of days are expressed as decimals hence 6 h is 0.25 of a day. The calculation from Gregorian date to JD is straight forward but long and laborious which makes it an ideal candidate for a small computing project. However, it can be calculated in a truncated form using Eq. 4.1 for dates after the 1st January 2000. For Eq. 4.1 Y is the current year, D is the number of days since the start of the current year[1] and L is the number of leap days since 2001. For fractions of a day take the number of fractional days since midday on the day you are working on and add it to the Julian date found using Eq. 4.1. Hence 6 am is −0.25 whist 6 pm is +0.25

The reader should be aware that a number of different Epochs are used for the Julian Date other than noon on 1st January 4713 BC, for example Truncated JD uses Midnight 24th May 1968 and Dublin JD uses Midday 31st December 1899 so be aware what system you need to use and ensure your use is consistent. If you are looking at a Julian date that pre-dates the introduction of the Gregorian Calendar I recommend using a long form equation rather than Eq. 4.1.

$$JD = 2{,}451{,}544.5 + 365 \times (Y - 2000) + D + L \tag{4.1}$$

4.4 Sidereal Time

A mean solar day, the time in between consecutive solar transits from East to West, is 23 h 56 min 4 s in duration. This includes the approximate one degree per day displacement of the Sun caused by the Earth's orbit (see Fig. 4.3). However, we use a 24-h clock in our day to days lives. Hence **Sidereal Time**, the time between

[1]Note that most spreadsheets find this very quickly.

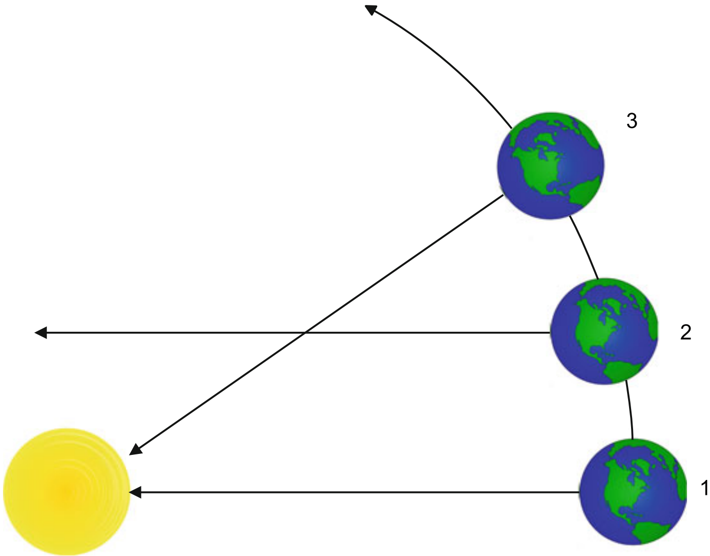

Fig. 4.3 Illustration showing the difference between solar and sidereal time. At position 1, the Sun is directly overhead. At position 2 the Earth has made one full rotation but due to its orbit around the Sun, the Sun is not now overhead. At position 3, the Earth has rotated for a further 4 min bring the Sun back to being overhead

consecutive stellar meridian crossings, migrates away from solar time at a rate of approximately 4 min per day, with the exact difference determined a number of additional the effects in addition to the difference between the lengths of the solar and sidereal day. Given that the **Local Sidereal Time** (LST) is the RA of the objects crossing the meridian at that moment we can see that an observer needs to know their local sidereal time when they intend to make their observations, in order to convert from RA and Dec used to locate the object on the celestial sphere to the AltAz used to point the instrument

In order to find the local sidereal time, we need first to find the current GMT. You should be aware of your local time zone adjusted for any daylight saving. Let us use an observer in Boston as an example. Boston is in UTC-5 hence when it is 14:00 is Boston the GMT is 19:00 as $GMT = LT - TimeZone$.

By finding the number of days since midday on the 1st January 2001 D, where partial days are expressed as fractions, we can apply Eqs. 4.2 and 4.3, to determine the Greenwich Sidereal time. Reversing the operation we undertook to find GMT gives us our Local Sidereal time.

$$t = 18.697374558 + 24.06570982441908 \times D \tag{4.2}$$

$$GMST = (t/24) - int(t/24) \tag{4.3}$$

For example, we wish to know what time Vega will cross the median from the view point of an Observer in Boston on the 7th July 2015. Boston is at 43°21’ 29”

N 71°3’ 49” W and Vega is located at $18^h\ 36^m\ 56^s\ +38°\ 47'\ 01”$. Hence, the transit will occur at approximately 18:37 LST. We need to find the GST for that LST that we do as follows:-

1. Convert the LST to decimal hours. So $18^h\ 36^m\ 56^s$ becomes 18.61555556.
2. Convert the longitude into decimal degrees. So 71°3’ 49” W becomes 71.06361111°W.
3. Covert the longitude to hours by dividing by 15 to give 4.737574074.
4. If the location is west of Greenwich add the longitude to the LST if it is East subtract it. If the result exceeds 24 subtract 24 if it is negative add 24. Hence our decimal GST is 23.35312963 or $23^h\ 21^m\ 11^s$.

Now we know the GST of the transit we need to find UT for $23^h\ 21^m\ 11^s$. GST on the 7th July 2015. For this we need the Julian Date at 0^h, which as found as follows:-

1. Set the year month and day as y,m and d respectively. Hence, 7th July 2015 becomes y=2015, m=7 d=7.
2. If $m \leq 2$ then subtract 1 from y and add 12 to m.
3. If the date is greater than 14th October 1582 find:
   $A = INT(y/100)$ $B = 2 + A + INT(A/4)$
   Hence $A = 20$ and $B = -13$
   If the date is earlier than 15th October 1582 $B = 0$.
4. If $y < 0$ $C{=}INT(365.25 \times y) - 0.75$ otherwise $C = INT(365.25 \times y)$. Hence $C = 735978$.
5. $D = INT(30.6001 \times (m + 1))$. So $D = 244$.
6. The Julian date is then $JD = B + C + D + d + 1{,}720{,}994.5$ Or for 7th July 2015 2,457,210.5

We can now find the UT for the GST we found above.

1. Calculate $S = JD - 2{,}451{,}545.0$ S = 5665.5
2. Calculate $T = S/36{,}525.0$. T = 0.155112936
3. Calculate $T_0 = 6.697374558 + (2{,}400.051336 \times T) + (0.000025862 \times T^2)$.
   $T_0 = 378.9763853$.
4. Reduce T_0 to the range 0–24 by using $T_0 = T_0 - INT(T_0/24) \times 24$.
   $T_0 = 18.97638528$
5. Convert GST to decimal that we already have. 23.35312963.
6. Subtract T_0 from the decimal GST. Adding or subtracting 24 as appropriate. Which gives 4.376744348.
7. Multiplying by 0.9972695663 gives the decimal UT, i.e., 4.364793938 or $4^h\ 21^m\ 53^s$ UT.

Boston’s local time is UTC-5 so the local UT is $23^h\ 21^m\ 53^s$. As it turns out, this is the transit for the previous evening. However, we can safely assume that the transit will be around the same time. Otherwise, we will have to perform the calculation.

Chapter 5
Spheres and Coordinates

Abstract In observational astronomy we are dealing with the Celestial Sphere, which is non-euclidian. Hence, astronomers use the non Cartesian Alt-Az and Equatorial coordinate system to point telescopes and identify objects. The non-euclidian nature of our observations means that we have to use spherical geometry when measuring the position of objects in the sky.

5.1 Introduction

In astronomy, the current position of stellar objects depends on its position in the sky about fixed points on the sky, the observers position on the Earth and the current time. All of these depend on spherical geometry.

Let us consider the sphere. Draw a line through the centre of the sphere intersecting the surface at two points such that the length of the line is the sphere's diameter, for rotating bodies such as the Earth, we make this the axis of rotation. In order to give us some orientation we shall call the top most point of interception the **North Pole** and the bottom the **South Pole**, although this is entirely arbitrary. We draw a circle intercepting both the North and South Poles so that this circle with the same circumference as the sphere. Such a circle is known as a **great circle** and if we used it to bisect the sphere it would do so through the centre. We shall call this great circle the **Meridian**. We now draw a second great circle that is perpendicular to the Meridian. This second great circle is the **Equator** although when used within a coordinate system it is more correctly known as the **Fundamental Plane**, see Fig. 5.1. Any circle on the surface of the sphere that does not bisect the centre of the sphere is know as a **small circle.**

One of the two points where the fundamental plane intercepts the meridian becomes the origin for the coordinate system using that fundamental plane. An object can be located on the surface of this sphere by using just its angular distance from the meridian line and its angular distance above or below the fundamental plane.

The most common such arrangement is the familiar longitude and latitude system for determining position on the Earth where the two great circles are the equator and the prime meridian, which is perpendicular to the equator, and runs through the poles and Greenwich in London. Hence, the origin is the point at which the

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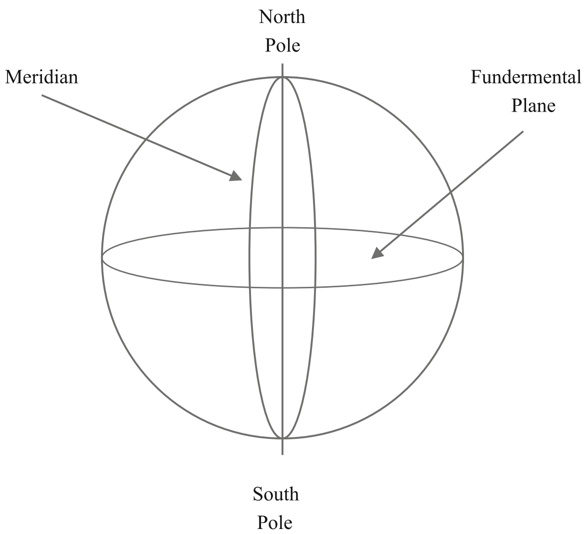

Fig. 5.1 Figure showing the position of the fundamental plane and the meridian in relation to the poles

prime meridian intersects the equator and is located in the Atlantic off the coast of Equatorial Guinea. Typically longitude and latitude are listed in degree minutes and seconds North or South of the Equator and East or West of the meridian. However, in some cases you may find that decimal degrees are used or that the cardinal points are replaced with a plus (for North or West) or a minus (for South and East). In some cases, for example working out Sunrise or Sunset times, you will need to know the observatory's altitude. Although typically altitude is taken from mean sea level, Global Positioning System (GPS) use true spherical coordinates in order to determine position, therefore the report altitude is based on the distance from the centre of the Earth with the Earth's radius taken as being 6,378.137 km (if you have a handheld GPS you may notice in the setting that it uses a coordinate system known as WGS84). Additional; the GPS system does not translate perfectly into geographical systems Although this is unlikely to cause any problems for an astronomical observer as the error can be less than 100 m, you should be aware of this if the telescope you are using uses a GPS receiver to determine time and position.

5.2 Altitude and Azimuth

The simplest coordinate system you will encounter in astronomy is the Altitude and Azimuth system, also know as Horizontal system, but more commonly known as AltAz. AltAz and is based on the two great circles formed by the observers horizon and the celestial meridian line. The celestial meridian line passes through the **zenith**, the point directly above the observer, and the **nadir**, the point directly underneath the observer intersecting the North and South cardinal points. For an observer in the Northern Hemisphere, the origin point is the point on the horizon due south of the observer. For observers in the southern hemisphere, it is the point on the horizon due north of the observer.

The AltAz system is determined by features local to the observer and is not fixed to the celestial sphere. Hence, astronomical objects transit through the AltAz grid during the course of the night as the Earth rotates. However, it is the AltAz grid that is used to point a telescope and the observer must translate between the coordinate system of the target and AltAz. Objects are located using its Azimuth (A) the objects angular position along the **Celestial Equator**, the Earth's equator projected onto the sky, measured Eastwards from, normally but not always, due North and its Altitude (α), the angular distance of the objects current position and the local horizon. Hence, for an observer in the Northern Hemisphere an object laying due south on the horizon has an Alt $= 0°$ and Az $= 0$ °, whilst the point laying due north on the horizon has Alt $= 0°$ and Az $= 180$ °. A star directly overhead, i.e., at the zenith, would have Alt $= +90°$ while it's Az could be anything as all azimuth angles converge at this point. For observers in the southern hemisphere it would be an object to the North on the horizon that would have an Alt $= 0°$ and Az $= 0°$ whilst the object at the zenith would have an Az of $-90°$.

As AltAz is a local system it is used not only to determine when and where to point but also to determine if the target is visible, if the target is in its optimal position in the sky for observing and if the location of the target is going to exceed the telescopes pointing limits. In many cases the instrument you will be using will have some limitation as to where it can point. This is typically the minimum altitude that can be observed, which is rarely as low as 0° as often there is some obstruction, for example, the edge of the dome or trees, which prevents this. Furthermore, with the increasing trend of bigger and bigger astronomical cameras in recent years, maximum altitude limits have also become more prevalent due to the inability of these larger cameras to pass safely under the telescope forks. Although somewhat rarer than altitude limits some sites also have azimuth limits due, for example, to physical obstructions, problems with prevailing winds or mount limitations. In many cases, you may find that a telescope has declination limits. Typically this is done when the telescope is **polar aligned** (when the Az axis of the telescope points at the pole), as most are. Declination limits are similar to altitude limits but are adjusted for the tilt in the telescopes vertical axis due to the polar alignment. Before you observe you should be aware of the telescopes pointing limits be they in altitude, declination or azimuth and the restraints it may put on your observations (Fig. 5.2).

The axis of the Earth's rotation is tilted by approximately 23.5° in reference to the plane of it's orbit around the Sun. Consequently the Sun the planets and the stars appear to follow an arc across the sky as the Earth rotates. Like all the planets in the solar system effective lie on the same plane, the Sun and the planets appear to be fixed on a line, the **Ecliptic**, inclined by 23.5° to the horizon. Originally this line was straddled by 12, **Zodical** although for reasons to be explained later, Ophiuchus now also straddles the Ecliptic. The Ecliptic intercepts the Celestial Equator at two points are the vernal equinox and autumnal equinox. The days at which the Sun is seen

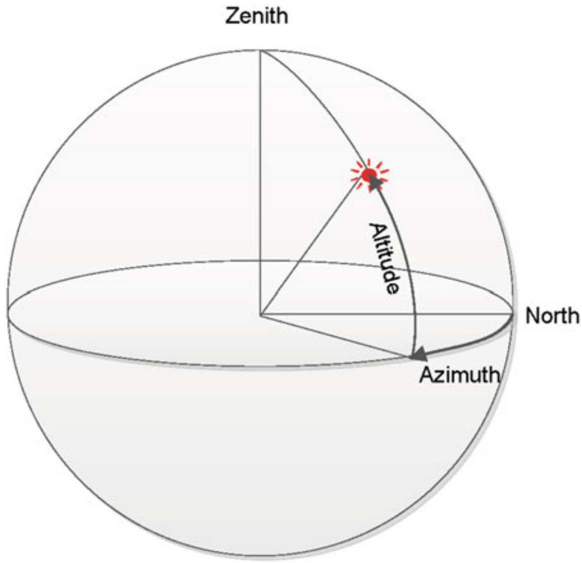

Fig. 5.2 Figure showing the alignment of the AltAz system. Altitude is measured in degrees from the horizon toward the zenith while Azimuth is measured anticlockwise from due north

rising at these points mark the start of astronomical spring and autumn respectively, in addition to being the point of time where day and night are of equal length, hence the name. The Solstices are the two points where the Sun rises the furthest North and South (and hence the longest and shortest days) and marks the start of astronomical summer and winter.

5.3 Equatorial

The Equatorial system is the most common coordinate system in astronomy. It uses the Earth's equator projected onto the celestial sphere, the celestial equator, as its fundamental plane, the line passing through the Vernal Equinox (also known as the First Point of Aries, which for reason that will become clear is no longer in Ares) and the North and South celestial poles to determine the origin.

The Equatorial system identifies an object by its Right Accession (*RA* or α) the angular distance Eastwards along the celestial equator between the object and the First Point of Aries, and it's declination (*Dec* or δ), the angular separation between the object and the celestial equator with an object laying on the celestial equator being 0°, objects North of the line have positive declinations and those south negative. Typically Right Accession is measured in Hours, Minutes and Seconds and Declination in Degrees, Minutes and Seconds, **Sexidecimal notation**. If this is the case, your should be aware that minutes and seconds in RA are *not* the same length as minutes seconds in Dec. If you performing positional measurements, you may wish to convert your RA and Dec to decimal as this resolves this problem.

You should be aware that on occasion decimal degrees maybe required or supplied. In general decimal degrees are just a decimal number while Sexidecimal notation has the individual components separated by either space, hms (for hours, minutes and seconds), or more commonly, a colon. Hence the star known as Betelgeuse or α Orion has an RA and Dec of 05 h 55 m 10 s +07 d 24 m 25 s or 05:55:10+07:24:25 or in decimal, 238.7917+7.406944

The Earth is precessing on its axis, like a slowing spinning top. Hence over very long periods points that seem fixed, for example, the celestial poles, appear to draw out a circle. It is for this reason[1] the First Point of Ares is no longer is Ares and also why there are now 13 zodical constellations. Over short periods, nutation does not pose a problem for astronomers. However, over longer periods stars will move away from their catalogue positions. In an attempt to resolve these astronomical catalogues have fixed dates for the positions they represent, this is known as the **Epoch**. The current Epoch is J2000 and uses positions determined from Midday 1st January 2000. The previous Epoch was B1950, which was in use prior 1984. Although you should only encounter J2000 coordinates, if you are using an older catalogue or academic paper you may come across B1950, or even older epoch systems. In such cases, you should endeavour to find the objects J2000 coordinates by using, for example, the SIMBAD astronomical catalogue. You may also encounter JNow when using astronomical software. This is the current object RA and Dec based on its J2000 positions, the current date and the precision rate. In general users should avoid using JNow as at the time of writing the difference between J2000 and JNow are minimal. However, as J2000 becomes increasing outdated JNow may become necessary (Fig. 5.3).

For an observer in the Northern Hemisphere stars with declinations greater than $\theta - 90°$ (where θ is the observers latitude) should be visible at some point during the year, although this depends on the quality of the local horizon. Likewise in the Southern Hemisphere stars with declinations less than $\theta + 90°$ will be visible.

This leads to an important observation. If not all stars are visible for an observer that implies that some are always visible. Such stars are known as **Circumpolar**.[2] Stars are circumpolar in the Northern hemisphere if $\theta + \delta \geq 90°$, likewise in the Southern Hemisphere if $\theta + \delta \leq -90°$. So for an observer in Greenwich where $\theta = 51.48°$ the star Vega $\delta = 38.8°$ is just circumpolar.

Another great circle is one that runs perpendicular to the equator and pass through the pole. Hence, every point along that line between the two poles and on the same side of the sphere have the same right accession. This line is known as a **hour circle** and the meridian is a special member of the set of hour circles, being the only one that pass through the origin. The angular distance between the hour circle on which an object current lies and meridian is the **hour angle** which is simply the Local Sidereal time minus the RA of the object.

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[1] As well as others that are beyond the scope of this book.

[2] Technically stars that are also never visible are also Circumpolar but this term is hardly ever used in this context.

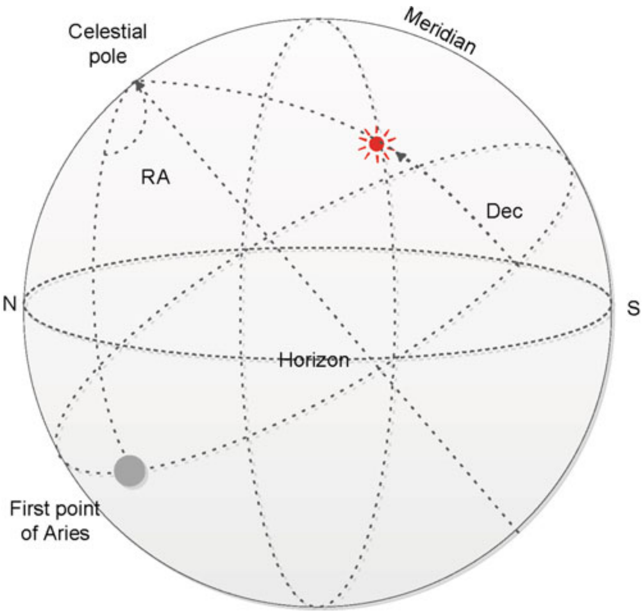


Fig. 5.3 Figure showing the alignment of the equatorial coordinate system

5.4 Galactic Coordinates

In 1958, in part to overcome the problem with the precession, the **International Astronomical Union (IAU),** the defacto governing body for astronomy, introduced the Galactic Coordinate system. In this system the origin is the galactic centre that is marked by the radio source (and supermassive black hole) Sagittarius A. The North Galactic Pole is designated as having a latitude (b) of $+90°$, the South Galactic Pole is designated as having a latitude of $-90°$ with longitude (l) measured in degrees right-handed along the galactic plane with the Galactic being $l = 0°$ and the anti-centre $l = 180°$. Hence in this system Betelgeuse is located at 199.7872–08.9586.

Galactic Coordinates are most commonly used by galactic astronomers, those whose interests line within the galactic plane, however for extra galactic uses the system is without merit and, in general, the Equatorial system is used in these cases, although a Super Galactic coordinate systems for extragalactic use does exist.

Because the system uses the galactic centre as its point if origin the rate of precession is very small as the orbital period of an object around the galactic centre is of the order of 200 million years. For high **proper motion** objects, the proper motion being the movement of stars against the more distant background, the rate of change is high as it is for Equatorial coordinates. However, these objects are normally extremely close and, therefore, are unlikely to lie on the galactic plane, where Galactic Coordinates are used.

5.5 Other Minor System

An earlier system, now mostly used by those interested in objects within the solar system, is **Elliptic Coordinates.** The fundamental plane for Elliptic Coordinates is the ecliptic, the arc travelled by the Sun and to some extent the planets and the origin is the First Point of Aries. Objects are located by their ecliptic latitude β and ecliptic longitude δ

The **Super Galactic system** is a coordinate system for extra galactic objects. It uses a flat, two-dimensional structure formed by the distribution of nearby galaxies know as the **Supergalactic plane**, as its fundamental plane and its inception point with the galactic plane as the origin. The system can be extended three-dimensional with galaxies on the plane having an SGZ value of zero. In deference to Galactic Coordinates super galactic latitude is noted as SGb and super galactic longitude is noted as SGb.

5.6 Spherical Geometry

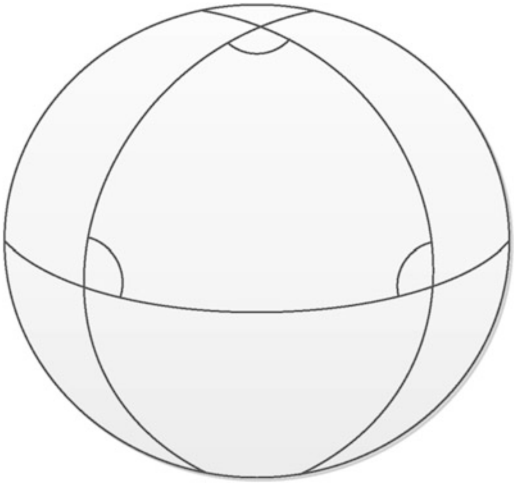
As you have seen astronomical coordinate system are based on spheres. However, the geometry of a sphere is non-euclidean and, therefore, the measurement of distances and angles within astronomical coordinates requires the application of spherical rather than Euclidean geometry.

Perhaps one of the more day to day applications of spherical geometry is the routes taken by aircraft. You may have noticed on intercontinental flights where the aircraft has a GPS location system in the entertainment system, that the flight is not a straight line between the point of departure and the destination but is, in fact, an arc. In spherical geometry a line is the shortest distance between any two points and, except in the case where those two points can be connected by a line passing through the centre of the sphere, that line will always be a segment of a great circle.

Likewise, we are told early in our education that the sum of the interior angles of a triangle adds up to 180°. Now let us consider a sphere and three great circles, one the equator and one the meridian and one perpendicular to both the meridian line and the equator that I shall call the offset. These three circles form a triangle with the vertices being located at the pole where the meridian and the offset meet, the interception of the Equator and the meridian and the interception of the Equator and the offset. As can be seen from Fig. 5.4, all these points have interior angles of 90° in violation of the previous fore mentioned rule. In fact, the interior angles of a triangle on the surface of a sphere always are greater than 180° and less than 540°, Fig. 5.4.

In many cases you will either be translating between coordinate system or measuring the distances between two points. In general given the amount of computation to convert between systems and the likelihood of error I strongly recommend either the Coco programme contained within the StarLink package or using the Python astropy.Coordinates package.

Fig. 5.4 Image showing the formation of a triangle on the surface of a sphere where the sum of the internal angles exceeds 180°



As to the measuring of the distance between two points, this will become increasingly important as we progress through this book. In general Eq. 5.1 is sufficiently precise for most of our uses. In this case $\phi_1, \phi_2, \lambda_1, \lambda_2, \Delta\lambda$ and $\Delta\phi$ are the latitude of points 1 and 2, the longitude of points 1 and 2, the difference between the longitudes of 1 and 2 and the difference between the latitudes of 1 and 2, respectively, with $\Delta\sigma$ being the central angle between points 1 and 2.

$$\Delta\sigma = arctan\left(\frac{\sqrt{(\cos\phi_2 \sin\Delta\lambda)^2 + (\cos\phi_1 \sin\phi_2 - \sin\phi_1 \cos\phi_2 \cos\Delta\lambda)^2}}{\sin\phi_1 \sin\phi_2 + \cos\phi_1 \cos\phi_2 \cos\Delta\lambda}\right) \tag{5.1}$$

In situations where you measure the differential offset, i.e., when the offset from the origins is not known or perhaps trusted, a different approach maybe taken. If, for example, we have two reference positions A and B and we wish to find the location of a point C in relations to points A and B we can form a triangle on the sphere with edges AB, AC and BC and internal angles A, B and C such a triangle is shown in Fig. 5.5. In this situation we can use Eqs. 5.2, 5.3 and 5.4 and their derivatives to find the location of the unknown point.

$$\cos a = \cos b \cos c + \sin b \sin c \cos A \tag{5.2}$$

$$\cos b = \cos c \cos a + \sin c \sin a \cos B \tag{5.3}$$

$$\cos c = \cos a \cos b + \sin a \sin b \cos C \tag{5.4}$$

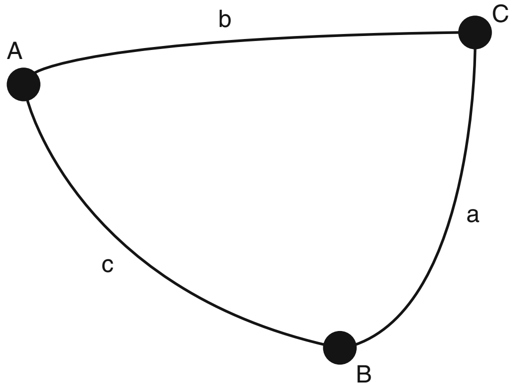


Fig. 5.5 The above diagram shows three points *A*, *B* and *C* on the surface of a sphere that are connected by the edges *a*, *b* and *c* to form a *cap*

as the distance between A, B and C tends to zero Eqs. 5.2, 5.3 and 5.4 tend towards the sine rule (Eq. 5.5). You may, in certain situation therefore resort to Eq. 5.5 but be aware that it introduces further errors.

$$\frac{\sin A}{\sin a} = \frac{\sin B}{\sin B} = \frac{\sin C}{\sin c} \tag{5.5}$$

Chapter 6
Catalogues and Constellations

Abstract Constellations are an important part of many of many cultures. Astronomers have adopted the Greek constellations to provide a regional map into which to located objects. Before the first telescope astronomers have been classified and catalogued objects, so that there know are in access of 200 such catalogues. These knowing, which catalogue appropriate to the task at hand is an important tool in observational astronomy.

6.1 Introduction

Astronomer do like to order things, to find them, to identify what they are and to put them into neat pigeon holes, perhaps not always successfully as the still ongoing problem with Pluto testifies. Clearly with 200 billion stars in the Milky Way and 200 Billion Galaxies, not to mention the billions of solar system objects, nebula, supernova remnants, X-ray sources, exoplanets, masers, radio sources etc, there is a lot to catalogue.

Consequently, there are a very large number of astronomical catalogues (in excess of 200) some of which have a very considerable overlap and some that have a small number of members. This chapter is here to help you to identify an object type from the catalogue and highlight the more important and usable astronomical catalogues.

6.2 Constellations, Asterisms and Bright Stars

Many cultures have used the heavens to tell stories and explain their cultural mythology. Stars were grouped to form images in a time when skies were darker, the night more intimating and imaginations perhaps more vivid, than in more modern times. These imaginary figures became the constellations. The stars within the constellations are just random associations with no actual physical connection between members of any one constellation; however they have an important role in both astronomy and the cultural identify of the civilisations that created them. Modern astronomers use the constellations as derived from Greek mythology with

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some additional constellations added at later dates that consist of the stars too far south to be seen by from the region occupied by Greek civilisation. It is also common, especially when speaking to non-astronomers to use an anglicised version of the Greek name hence The Great Bear is Ursa Major and the Swan is Cygnus.

The use of constellations may seem archaic but they simplify the naming scheme of many astronomical objects in that they act in many respects as the country or state does as part of a postal address. Although there is no need to be able to identify all 88 as constellations, defined by the International Astronomical Union (IAU), by sight, you should be aware of the brighter and more prominent ones visible from you site, for example Ursa Major for northern observers, Crux for southern observers and Orion and Leo for all observers. A full list of constellations can be found in Table 6.1 which also shows their official abbreviations and the hemisphere in which they are located.

Asterisms, like constellations, are collections of stars that often have no association except their location in a particular region of the celestial sphere. Technically constellations are a subset of asterism in that they are formally accepted were as most asterisms are just the brightest members of a constellation, a combination of constellations, or even a small segment of a constellation that form a very easily identified structure. Hence, astriums, in general, have much fewer members than constellations. Perhaps the most well know asterisms are the Plough (also known as the Big Dipper or the Sauce Pan) which is formed from the brightest seven members of the constellation Ursa Major and the Summer Triangle, which is a northern hemisphere feature formed from the stars Deneb, Vega and Altair, which are the brightest members of Cygnus, Lyra and Aquila respectively.

If you are unlucky enough to be observing from a light polluted site, it is probable that the asterism formed from bright stars will be more prominent than the constellations. Likewise, if you are setting up during twilight when faint stars are still not visible. Bright stars are very important in observational astronomy. They are used to set up the alignment of GOTO telescopes,[1] as a check of alignment, a focusing target and a first point of call when star hopping if you are not using a GOTO. It is unlikely that you will be able to identify the first visible star of the evening's observing (which might even be a planet) especially if you have been unable to observe for some time. However, as the sky darkens and more stars become visible the astriums will appear to form first allowing you to identify a few stars and get set-up before it is truly dark.

[1]GOTO telescopes have powered mounts and either a customer built interface or computer control that allows the users to rapidly move between targets. Typically the users have to point the telescope to three guide stars to set it up.

Table 6.1 The above table lists the 88 constellations as defined by the IAU. Col 1: Name Col 2: Abbreviation Col 3: Hemisphere located. Note that constellations listed as Northern may be visible in the southern hemisphere and Vice versa. Likewise, some constellations extended over the equator

Name	Abbreviation	Hemisphere	Name	Abbreviation	Hemisphere
Andromeda	And	N	Lacerta	Lac	N
Antlia	Ant	S	Leo	Leo	N
Apus	Aps	S	Leo Minor	LMi	N
Aquarius	Aqr	S	Lepus	Lep	S
Aquila	Aql	N	Libra	Lib	S
Ara	Ara	S	Lupus	Lup	S
Aries	Ari	N	Lynx	Lyn	N
Auriga	Aur	N	Lyra	Lyr	N
Boötes	Boo	N	Mensa	Men	S
Caelum	Cae	S	Microscopium	Mic	S
Camelopardalis	Cam	N	Monoceros	Mon	N
Cancer	Cnc	N	Musca	Mus	S
Canes Venatici	CVn	N	Norma	Nor	S
Canis Major	CMa	S	Octans	Oct	S
Canis Minor	CMi	N	Ophiuchus	Oph	S
Capricornus	Cap	S	Orion	Ori	N
Carina	Car	S	Pavo	Pav	S
Cassiopeia	Cas	N	Pegasus	Peg	N
Centaurus	Cen	S	Perseus	Per	N
Cepheus	Cep	N	Phoenix	Phe	S
Cetus	Cet	S	Pictor	Pic	S
Chamaeleon	Cha	S	Pisces	Psc	N
Circinus	Cir	S	Piscis Austrinus	PsA	S
Columba	Col	S	Puppis	Pup	S
Coma Berenices	Com	N	Pyxis	Pyx	S
Corona Australis	CrA	S	Reticulum	Ret	S
Corona Borealis	CrB	N	Sagitta	Sge	N
Corvus	Crv	S	Sagittarius	Sgr	S
Crater	Crt	S	Scorpius	Sco	S
Crux	Cru	S	Sculptor	Scl	S
Cygnus	Cyg	N	Scutum	Sct	S
Delphinus	Del	N	Serpens [5]	Ser	N
Dorado	Dor	S	Sextans	Sex	S
Draco	Dra	N	Taurus	Tau	N
Equuleus	Equ	N	Telescopium	Tel	S
Eridanus	Eri	S	Triangulum	Tri	N
Fornax	For	S	Triangulum Australe	TrA	S

(continued)

Table 6.1 continued

Name	Abbreviation	Hemisphere	Name	Abbreviation	Hemisphere
Gemini	Gem	N	Tucana	Tuc	S
Grus	Gru	S	Ursa Major	UMa	N
Hercules	Her	N	Ursa Minor	UMi	N
Horologium	Hor	S	Vela	Vel	S
Hydra	Hya	S	Virgo	Vir	S
Hydrus	Hyi	S	Volans	Vol	S
Indus	Ind	S	Vulpecula	Vul	N

6.3 Catalogues

Broadly catalogues can be divided into stellar and non-stellar catalogues, although a small number are mixed. Additional catalogues can be further split into the historical divide of visual, photographic and digital observing, although a number of photographic surveys have now been digitised, which is a boon for those who wish to undertake proper motion studies that require large gaps between observations.

You should be aware that many catalogues are geographically limited, in that they are taken with a single ground-based instrument that cannot see the entire sky. Likewise not all objecting being capable of being detected by the survey will be in the catalogue as some will always be missed. The number of detected objects compared to the number of expected detected objects from, what are now days computer models, is the **Completeness**. Most modern surveys have completeness levels in excess of 90 % but you should not be surprised if you find an object that is not in the catalogue you are using.

As stated above, the bright stars are extremely useful for anyone observing the night sky and the ability to identify bright stars will greatly enhance the amount of observing time you will get at most teaching observatories, especially when things go wrong as they often do. Unlike constellations the bright star common names come, mostly, from the Arabic, with some Greek and Anglicised names in use. Most telescope equipped with GOTO controllers use the common name for bright stars.

Bayer Catalogue

Most astronomers use Johann Bayer's 1603 catalogue of approximately 2,500 of the brightest stars. Bayer used Greek letters and the constellation name, now often abbreviated, to identify each star in his catalogue, with alpha being the brightest, beta the second brightest, etc. Hence, α UMa is the brightest star in the constellation of Ursa Major. Of course there are far more stars in any constellation than there are Greek letters (in fact there are approximately 10,000 stars visible with the naked eye) so this process only applies to the brightest but it is very useful none the less.

While we are on the subject of Bayer's catalogue, you should be aware that there are a number of errors regarding the ordering by brightness of some stars. For example, α and β Gem are the wrong way round. This error and others have never been corrected and their existence is a topic of discussion with the astronomy community as it is not clear if they just a mistake or if there has been an actually physical change in one or more of the stars. Whatever the reason you should attempt to learn the common and Bayer names of the brightest stars visible form the location you intend to observe from, A list some of these stars with their common names is presented in Table 6.2 along with their visual magnitudes.

Flamsteed Catalogue

Bayers catalogue was improved upon by the introduction of a new catalogue by the first Astronomer Royal, John Flamsteed in 1712. As with the Bayer catalogue, the Flamsteed catalogue is geographically limited and did not originally cover the entire sky, although the catalogue has been extended after the Flamsteed's death in order to cover the region too far south to be seen from Greenwich, Flamsteed used a similar naming method to that used by Bayer but used numbers as opposed to Greek letters, which significantly increased the number of names available and labelling in right accession order rather than magnitude Hence the star known as Betelgeuse is known as α *Ori* in the Bayer system and 68 Ori in Flamsteed's.

Bonner Durchmusterung Catalogue

The last great visual catalogue was that undertaken by the Bonn Observatory between 1799 and 1875 and extended in 1886 by Cordoba Observatory to cover the regions too far south to be observed from Bonn. The Bonner Durchmusterung survey, as it is now known contains approximately 325,000 down to magnitude 9.5. The catalogue is broken down into a series of 1° wide strips with stars numbered within that strip in order of their right ascension hence Betelgeuse is known as BD +07 1055 in the Bonner Durchmusterung catalogue.

Messier Catalogue

Charles Messier was an Eighteenth-century astronomer with a passion for comets. Unfortunately Eighteenth-century optics was not of the quality we have now and a very large number of non-comet objects looked like comets to Charles so he created a catalogue of 110 objects in order to avoid later confusion. This catalogue is the now famous Messier catalogue. Given that Messier was operating from France many

Table 6.2 Table is showing some of the brightest stars with their apparent magnitudes, Bayer identifier and their proper names

Magnitude	Bayer ID	Proper name	Magnitude Bayer ID	Proper name	
0.5	α Eri	Achernar	2	α Ari	Hamal
2.5	β Sco	Acrab	2.39	ϵ Boo	Izar
0.77	α Cru	Acrux	1.8	ϵ Sgr	Kaus Australis
1.51	ϵ CMa	Adara	2.08	β UMi	Kochab
0.6	β Cen	Agena,	1.96	δ Vel	Koo She
0.85	α Tau	Aldebaran	2.32	η Cen	Marfikent
2.44	α Cep	Alderamin	2.49	α Peg	Markab
2.01	γ Leo	Algieba	2.46	κ Vel	Markeb
2.12	β Per	Algol	2.28	α Lup	Men
1.9	γ Gem	Alhena	2.06	θ Cen	Menkent
1.76	ϵ UMa	Alioth	2.35	β UMa	Merak
2.15	γ And	Almach	1.68	β Car	Miaplacidus
1.74	α Gru	Alnair	1.3	β Cru	Mimosa
1.7	ϵ Ori	Alnilam	2.23	δ Ori	Mintaka
1.7	ζ Ori A	Alnitak	2.06	β And	Mirach
1.98	α Hya	Alphard	1.82	α Per	Mirfak
2.21	α CrB	Alphecca	1.98	β CMa	Mirzam
2.06	α And	Alpheratz	2.23	ζ UMa	Mizar
0.77	α Aql	Altair	2.17	γ Cen	Muhlifain
2.4	η CMa	Aludra	2.21	ζ Pup	Naos
2.37	α Phe	Ankaa	2.06	σ Sgr	Nunki
1.09	α Sco	Antares	1.91	α Pav	Peacock
−0.04	α Boo	Arcturus	2.43	γ UMa	Phecda
2.25	ι Car	Aspidiske	1.97	α UMi	Polaris
1.92	α TrA	Atria	1.15	β Gem	Pollux
1.86	ϵ Car	Avior	0.34	α CMi	Procyon
1.64	γ Ori	Bellatrix	2.1	α Oph	Rasalhague
1.85	η UMa	Benetnasch	1.35	α Leo	Regulus
0.42	α Ori	Betelgeuse	0.12	β Ori	Rigel
−0.72	α Car	Canopus	−0.27	α Cen	Rigil Kent
0.08	α Aur	Capella	2.43	η Oph	Sabik
2.27	β Cas	Caph	2.24	γ Cyg	Sadr
1.58	α Gem	Castor	2.05	κ Ori	Saiph
1.25	α Cyg	Deneb	1.86	θ Sco	Sargas
2.04	β Cet	Deneb	2.42	β Peg	Scheat
2.14	β Leo	Denebola	2.25	α Cas	Schedar
2.29	δ Sco	Dschubba	1.62	λ Sco	Shaula
1.79	α UMa	Dubhe	−1.46	α CMa	Sirius
1.68	β Tau	El Nath	1.04	α Vir	Spica
2.23	γ Dra	Eltanin	2.23	λ Vel	Suhail

2.4	ϵ Peg	Enif	1.78	γ Vel	Suhail
1.16	α PsA	Fomalhaut	2.39	γ Cas	Tsih
1.63	γ Cru	Gacrux	0.03	α Lyr	Vega
2.5	ϵ Cyg	Gienah	2.29	ϵ Sco	Wei
2.38	κ Sco	Girtab	1.84	δ CMa	Wezen

of the objects are not visible from the Southern Hemisphere. Also given that some of the objects, such as the M31 (The Andromeda Galaxy) and M45 (The Pleiades) are very well know, naked-eye astronomical objects, it is unlikely that all of the objects in the list would be mistaken for comets. However, the 110 objects in the Messier Catalogue are favourites among amateur astronomers for being bright, spread out over the sky and a mixture of **Deep Sky** astronomical object types, objects too faint to see with the naked-eye and outside the solar system.

Caldwell Catalogue

Complied in the 1990s by famed amateur astronomer and broadcaster, the Caldwell catalogue was compiled by Sir Patrick Moore (or more correctly Patrick Caldwell-Moore) the Caldwell catalogue is a list of deep sky objects omitted by the Messier Catalogue either due to error, deliberate omission or due to the object being to far south. The Caldwell Catalogue contains 109 objects including star clusters, planetary nebulae and galaxies.

United States Navy Observatory Catalogues

The United States Naval Observatory have produced a series of catalogues some of which are now considered obsolete. USNO A2.0 is the re digitisation of the plates from three previous catalogues. USNO-A2.0 contains in excess of 500 million objects with R and B magnitudes down to magnitude 20. USNO A2.0 has now been largely replaced by USNO-B1.0, which contains in excess of one billion objects down to magnitude 21. Both USNO-A2.0 and USNO-B1.0 are accessible via SAO DS9.

UCAC2 and its replacement UCAC3 are high precision (in regards to astrometry) catalogues to magnitude 16 in f band created from CCD images taken at Cerro Tololo Interamerican Observatory (CTIO) in Chile, the Naval Observatory Flagstaff Station (NOFS) in Arizona. The catalogue includes B, R and I magnitudes from the SuperCOSMOS survey. Both UCAC3 and UCAC2 are accessible via SA0 DS9.

The Naval Observatory Merged Astrometric Dataset (NOMAD) is a merged catalogue using the sources found within the Hipparcos, Tycho-2, UCAC2 USNO-B1.0 and 2MASS surveys. Hence NOMAD covers B, V, R, J, K and H bands down to approximately magnitude 20, depending on the source catalogue.

Hipparcos and Tycho

Hipparcos is a European Space Agency Telescope launched and operated between 1989 and 1993 with the goal of producing high precision proper motions and parallaxes to stars within 200 parsecs. Observations were made in B and V band with a limiting magnitude of 12.4 in V and resulted in an 118,218 source catalogue. Objects in Hipparcos are numbered and prefixed with HIP, hence α Ori is HIP 27989 in the Hipparcos catalogue.

Two further catalogues were derived from the Hipparcos observations. Tycho-1 contains approximately 1 million entries down to a V magnitude of 11.5 while Tycho-2, which supercedes Tycho-1, contains 2.5 million objects down to a similar limiting magnitude as Tycho-1. Tycho object numbers are prefixed with TYC and consist of a three numbers separated by a dash. The first number is the Hubble Guide Star Region; the second is the star number within that region and the last is the component identifier for multiple star systems. Hence α Ori is TYC 129-1873-1 in Tycho-2

The Hubble Guide Star Catalogue

The Hubble Guide Star Catalogue (GSC) was designed as the pointing catalogue for the Hubble Space Telescope and has been upgraded a number of times, with the current version at time of writing being GSC-2 which is intended to be the pointing catalogue for Hubble's successor the James Webb Space Telescope. The GSC contains over a billion objects with the majority of objects having J and f magnitudes down to around magnitude 21. As the GSC was designed for pointing, multiple star systems are excluded to avoid confusion within the pointing system. The GSC is a popular catalogue and is used in many astronomical packages. A digital version of the catalogue can be queried or downloaded from the Space Telescope Science Institute (STScI). The GSC divides the sky into regions and assigns a number to each source within the region. A third number (either a 1 or a 2) indicates which CD-ROM the star is located on and is now normally omitted. In the GSC format α Ori is GSC 00129-01873.

The Sloan Digital Sky Survey

The Sloan Digital Sky Survey (SDSS) is a very large, very deep survey being conducted by the Sloan Foundation 2.5-m optical telescope at the Apache Point Observatory. The Sloan is very fast for a large telescope being f/5 with 3 deg field of view. Sloan is equipped with a massive 120 Megapixels CCD and a 640 fibre integrated Fibre Spectrograph. The SDSS contains over 500 million stars with magnitudes in u, g, r, i, z (i.e. Sloan filters) down to magnitude 22. The catalogue also includes 1 million spectra of galaxies. SDSS can be queried via the SDSS website http://www.sdss.org/.

2MASS

2MASS is the 2 Microns All Sky Survey an astronomical survey of the whole sky in the near infra-red J,H and K_s (the s standing for short) utilising two 1.3 m telescopes located in Arizona and Chile. The survey contains 300 million stars or star like objects and a further 1,000,000 galaxies or nebulae like structures. Completed in 2003 the survey contains objects as faint as approximately magnitude 15.

Although 2MASS is an infra-red survey, it contains many optically red objects such as red dwarfs that are within the reach of many teaching instruments, missed by other surveys. Likewise, the catalogue has a high degree of completeness (i.e most of the objects within its sensitivity limits are detected) and hence most non-solar system objects you are likely to observer are going to have 2MASS numbers. 2MASS catalogue numbers are a representation of the items Right Accession and declination. Hence, α Orion is known as 2MASS J05551028+0724255 in 2MASS.

The NGC and IC Catalogues

The New General Catalogue of Nebulae and Clusters of Stars (NGC) is a catalogue of non-stellar sources undertaken by Danish-Irish astronomer John Louis Emil Dreyer in 1888. The catalogue contains a large number of galaxies that, at the time, were not fully understood, as well as star clusters and nebulae. In all the NGC contains 7,840 objects, including all of those in the Messier Catalogue. NGC objects are numbered and prefixed with NCG hence M31 is also known as NGC 224. The NGC catalogue was later supplemented by two further catalogues by Dreyer, known as the Index Catalogues (IC) which together adds a further 5,386 objects that are prefixed IC.

Uppsala General Catalogue

Uppsala General Catalogue (UGC) is a northern hemisphere catalogue of bright galaxies down to magnitude 14.5, which should be within the reach of most northern teaching observatories equipped with a CCD. The UGC contains 2,921 galaxies that are prefixed with UGC.

The Burnham, Aitken and Washington Double Star Catalogues

The Burnham, Aitken and Washington Double Star Catalogues are a generational series of double star catalogues initiated by Sherburne Wesley Burnham, who published his catalogue of 13,665 pairs of double stars. Burnham continued observing double Star and this further work was included in the Aitken double stars catalogue (ADS) which extends the Burnham catalogue to 17,180 double stars, although like the Burnham catalogue it is confined to mostly Northern latitudes. Both Aitken and Burnham have been superseded by The Washington Double Star Catalogue, which contains 115,769 pairs of double star positions along with the magnitude and spectral types. Washington Double Star Catalogue objects are noted by their J2000 position and prefixed with WDS.

Henry Draper Catalogue

The Henry Draper Catalogue (HD) is a comprehensive catalogue of 359083 (including various extensions) spectrally classified stars undertaking between 1886 and 1949 with a limiting magnitude of around nine, although in some cases it extended as faint as magnitude 11. The spectral classification used by Henry Draper in the catalogue is the basis for the standard OBAFGKM spectrographic classification scheme used by astronomers and astrophysicists worldwide. Objects are numbered and prefixed HD, HDC or HDEC depending on the catalogue version, although objects are number consecutively throughout all versions and supplements in order to avoid confusion. α Orion is known as HD 39801, in the Henry Draper Catalogue.

General Catalogue of Variable Stars

The General Catalogue of Variable Stars (GCVS) and its sister catalogue New Catalogue of Suspected Variable Stars is a catalogue of, not unsurprisingly, variable, or suspected variable, stars. The GCVS lists the stars position and the class of variable (which the suspected catalogue omits). Stars within the GCVS are prefixed with GCVS or in some cases, for example, the SIMBAD database are noted with a V*. The American Association of Variable Star Observers (AAVSO) also keeps variable star catalogue, the International Variable Star Index, which currently holds over 20 million variable stars many observed by amateurs.

6.4 Databases

Most of the catalogues listed above were originally supplied in printed form that makes searching and cross referencing somewhat tiresome. Later versions, such as the GSC and the SDSS are supplied on CD-ROM (or DVD) which are machine readable and allow for fast searches. However, the trend is now to have on-line access with catalogues accessible by a web-based SQL search tool. In some cases, multiple catalogues are held at a single site and maybe cross matched with relative ease.

SIMBAD

The most important online astronomical catalogue is **SIMBAD** (the Set of Identifications, Measurements, and Bibliography for Astronomical Data), an online catalogue formed from the concatenation of a large number of astronomical catalogues. At the time of writing SIMBAD contained over five million objects, all located outside of the solar system. SIMBAD is searchable by both name and location and interacts with the Aladin application used for creating, amongst other things, finder charts. It supports multiple coordinates systems, including J2000, B1950 and Galactic and provides bibliographical references to the objecting being searched for. SIMBAD can be accessed at http://simbad.u-strasbg.fr/simbad/ (Fig. 6.1).

Minor Planet Centre

For solar system objects (other than planets) the principle catalogue is the one held by the International Astronomical Union's Minor Planet Center (MPC) Like SIMBAD the MPC is a searchable database containing over 600,000 asteroids, comets, moons and other bodies whose positions are regularly updated in order to provide accurate ephemeris information (the data that allows the calculation of an objects position in the sky).

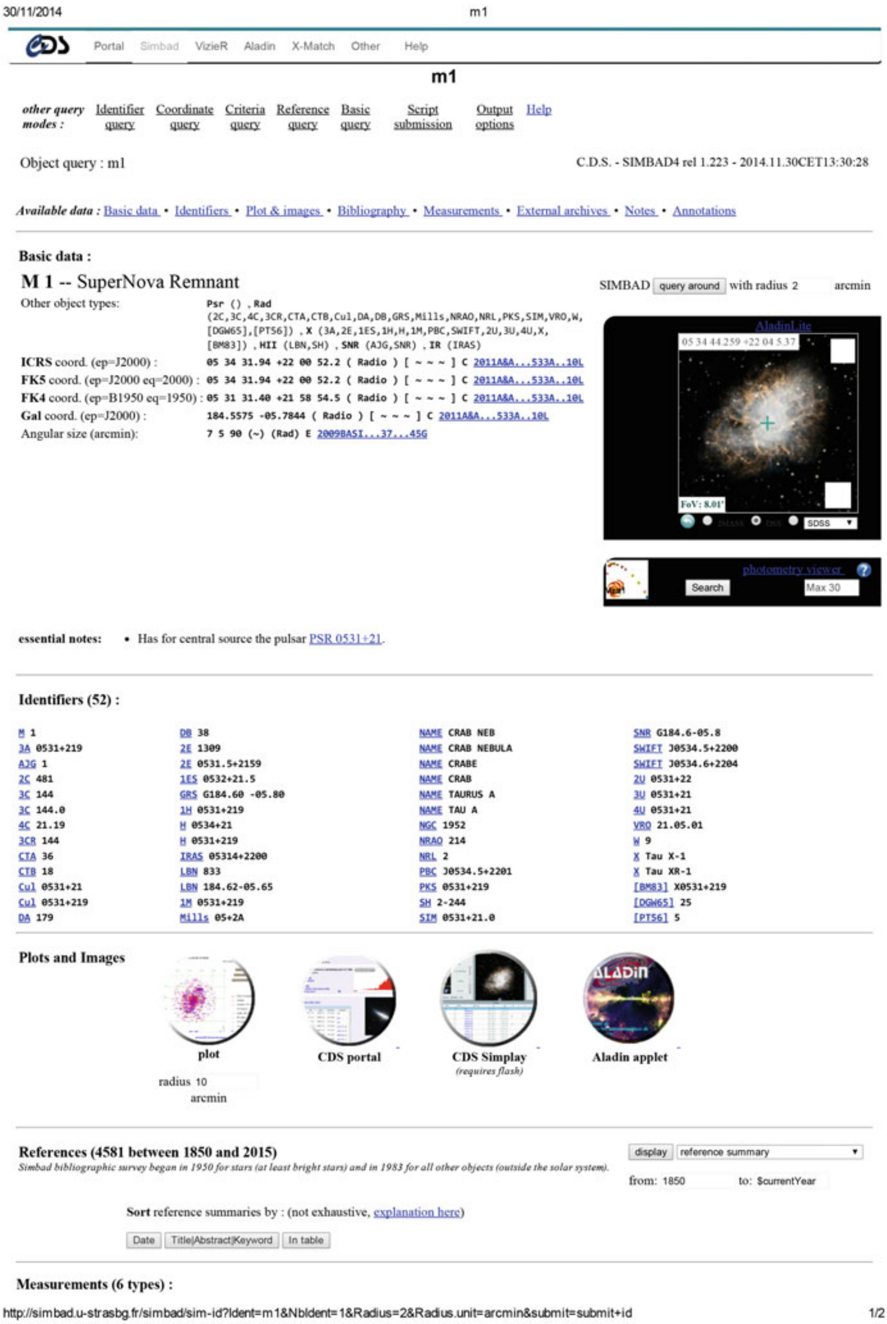

Fig. 6.1 Screen shot showing the results of an SIMBAD search for M1. The *top section* gives the objects position in multiple coordinate systems, provides its angular size, for resolved objects and its magnitude in multi bands (although for some reason not in this case). A zoomable image of the object to aid with the location is on the *top right-hand* side. Below is a list of all the identifiers associated with this object, 52 in this case, and links to that catalogue (Image supplied by NASA's Astrophysics Data System)

Infrared Science Archive

The Infra-red Science Archive (IRSA) is a multiple infrared and sub-millimetre source and image database jointly operated by NASA and the Infra-red processing centre (IPAC). It contains catalogues from surveys such as 2MASS, ISO, Planck, Spitzer, WISE and Herschel. IRSA can be accessed via the web interface located at http://irsa.ipac.caltech.edu/applications/Radar/.

VizieR

VizieR is an online web-based astronomical catalogue service hosted by Centre de données astronomiques de Strasbourg (CDS) who also host SIMBAD. VizieR currently holds 12,488 astronomical catalogues. Catalogues can be cross matched using the X-Match services also run by CDS.

Chapter 7
The Astronomical Detector

Abstract The Charge Couple Device is the foremost imaging device in observational astronomy. The CCD has high quantum efficiency and linearity when compared to film but suffers from a number of sources of noise. We illustrate how to reduce noise by backing and applying Flat, Bias and Dark frames. We show how to determine the linearity of the CCD, and how to recognise and deal with artifacts such as bad pixels, cosmic rays and saturation.

7.1 The Detector

Before the introduction of photographic film, the only usable astronomical detector was the human eye. Even through an astronomical telescope the vast majority of astronomical objects are not bright enough to be detectable with the human eye, especially in colour. In fact the **quantum efficiency (QE)**, the number percentage of photons that cause a physical response, is less than 5 % of optical observations. This figure is for the light-sensitive rods within the retina, the colour sensitive cones achieve much lower QE, and hence the colour is seen in a very small number of astronomical objects when viewed with the eye. The cones are concentrated in the centre of the retina and the rods round the periphery. A common technique use by amateur astronomers to see faint objects is to look at the target out of the corner of the eye in order to use the light-sensitive rods. Another drawback to observations using the eye is that the exposure time of the eye is around 1/30th of a second. Hence, this is often not sufficient for enough photons to hit the retina to cause a reaction.

The 1970s saw the introduction of the charge couple device (CDD) for professional astronomy, which is now a ubiquitous, component of almost all cameras, astronomical or terrestrial. The QE of modern astronomical CCD are now approaching 90 % and are close to being non-linear. The introduction of the CCD is one of the greatest driving factors in the major step forwards that have occurred in modern astronomy in the last few decades. The instrument you will be using will be an astronomical CCD, or contain a CCD (or the similar CMOS) in some form.

The CCD camera uses an array of light sensitive cells or **pixels**. You can consider each pixel to be a well with a series of balls (or in this case electrons) around the rim. As photons hit the well, they knock an electron into the well. At the end of

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the exposure, the number of electrons in each well is proportional to the number of photons that hit that pixel, the number of hits each pixel receiver is known as its **count**.

Looking at the literature supplied with an astronomical camera or a manufactures website, you will see a plethora of facts, functions and figures associated with a camera. For example backlit or frontlit, frame transfer rate, etc. In general this information will be at most interesting. However, there are a number of pieces of information you do need:

The **Array Size** is the size of the CCD in pixels. The Array Size is not the physical size as pixel sizes vary. In general bigger is better. However, very large CCD might not be fully illuminated by telescope so you instrument should match your telescope.

The **Pixel Size** is the physical size of each pixel, which in most cases are square, but no always. Typically, pixel sizes are in microns. You can use the array size and pixel size with the telescope's focal length to find the image scale as discussed in Sect. 3.1.

The **Quantum Efficiency** (QE), as mentioned earlier, is the measurement of how effective the chip is at capturing photons. The QE is wavelength specific and hence will change depending on the object being observed. For chips designed for optical observing the peak QE is around 500–600 nm, with a very sharp drop off in the blue end and a less sharp but signification drop off in QE at the red end. Hence, you may find that you need to exposure longer for some objects that others if their peak emission is outside the peak QE wavelength. Furthermore, QE is very sensitive to temperature, and you will not be able to achieve the expected sensitivity if the CCD is running hotter than specified, one of the reasons astronomical detectors are cooled. Additionally, the published QE for the camera is unlikely to achieved by every pixel in the array. This variation in QE can be a significant source of error and it can be dealt with in the most part by a process known as flat fielding that we will discuss at length later in this chapter.

The **ADU bit number** refers to the processing limitations of you camera. The ADU or Analogue to Digital Unit effectively counts the electrons trapped in each pixel and then converts it into a number. The number of bits limits the maximum number the ADU can report With the 8-bit camera's this limit is 256, for a 16-bit 65,536 and a 32-bit 4,294,967,296. Note that counts are dimensionless, non-negative and integer. The **Full Well Capacity** of the camera is the number of electrons each pixel can hold before it becomes unresponsive. A typical full well depth for a modern astronomical camera is 100,000–200,000 electrons. Having a higher Full Well Capacity means you can exposure for longer which can improve the quality of the data within the image. However, most astronomical cameras have an ADU that cannot count as high as the full well capacity and as a consequence the number of electrons in each pixel has to be scaled. The scaling figure is known as the **Gain**. You would be forgiven for assuming that gain for any camera is just the full well capacity divided by 2^{bit}, and in such a case the gain of a 16-bit 100,000 capacity CCD would be 1.5259. However, this is the lower gain limit, and typically camera have higher gains than this (Fig. 7.1).

Fig. 7.1 An SBIG Astronomical CCD camera. The CCD is clearly visible inside its window. The black tube on the side is full of dessicate to keep the air or inert gas, dry. The I-beamed shaped metal object in the centre of the image is the shutter

The camera you are using may be based on a CMOS (Complementary metal oxide semiconductor) rather than a CCD. CMOS chips have electronics attached to each pixel that allows for more rapid downloading of the image. Furthermore, a CMOS is less costly than a CCD to manufacture. However, CMOS cameras have lower QE when compared to CCD cameras, so although CMOS are found in domestic cameras, they are not yet suitable for astronomical, or very low light, imaging. None the less CMOS development is progressing rapidly and in time it may well be that CMOS chips overtake CCDs. However, some very sophisticated CMOS cameras for astronomy are used occasion when there needs a very high frame, such as when imaging asteroid occultations. However, these cameras tend to include a very powerful photomultiplier tube that improves sensitivity but significantly increases the cost.

7.2 Noise and Uncertainty in CCD

We now move on to some of the sources of noise associated directly with using a CCD. Noise in this context means anything that looks like part of the signal but does not come from directly from the source. Noise and uncertainness, in general, are covered in Chap. 11. This section largely deals with the noise sources directly linked to the CCD while Sect. 7.2 discusses how to remove it.

Not all pixels are created equal; there will always be a variation in the QE between individual pixels, which may change between exposures. This variation can be mostly address by the application of what is known as a flat field, which is discussed in detail in Practical 2. Knowing the flatness of your camera and the effectiveness of your flat field is essential when undertaking very sensitive photometry.

The electrons in each pixel are held in place by the application of a small charge. Intolerances in the makeup of the CCD and impurities in the silicon from which is constructed causes electrons to move between pixels a process known as **Charge Diffusion**. Given the nature of the causes of Charge Diffusion it will tend to be found in some parts of the CCD and not others. It is also to some extent wavelength depending. As with QE, Flat Fielding will deal in part with Charge Diffusion and for most applications that you will encounter it will not be a problem as other sources of noise may well be dominate.

To fully understand the operation of the CCD you will need to understand the processes that lead from an electron being liberated to an image being produced. For physical reasons, we cannot read all the pixels at the same time. Instead the camera firmware and hardware undertakes the following read process as illustrated by Fig. 7.2. You can consider the CCD to be a matrix of pixels with each pixel being identified by a row and column number. In general the only pixel that is read is the one at 0,0. Prior to being read and as an additional precaution a small number of additional electrons, are added to the read pixel. This addition, known as the **pedastal** is done to avoid a pixel going effectively negative (and potentially wrapping round, so that a zero becomes a 65,535 count) during later operations. The ADU then converts the number of electrons into a count by applying the Gain. This figure is then passed to an array. Once the 0,0 pixel is read and the are electrons removed from the pixel so that it now reads zero, the electrons in the next pixel up in the zero column are transferred into the pixel at position 0,0 and all the pixels in the column are moved down one. This process continues until the whole column is read. The camera then transfers the next column over to the now empty first column and all columns are moved over one. This cycle continues until all pixels are read. The consequence of this is that **read noise** increases as we move towards the read column. Some very large CCDs are multiple CCD joined and, therefore, have multiple read edges. In this case, you will see the level of noise increase towards the edges of the array, as opposed to one edge.

The read process cannot be 100 % efficient and some loss during the transfer process can be expected. The rate of loss is known as the **charge transfer efficiency** (CTE) and effects the pixel diagonally opposite the read pixel the most and the read pixel the least. Consequently the CTE results in a gradient appearing across the frame increasing outwards from the read pixel. For example in a 512 by 512 array the last pixel will go through 1,024 read cycles during a full frame transfer. If the CTE is 0.999999 then the last pixel read will have lost about 0.1 % of it electrons to CTE.

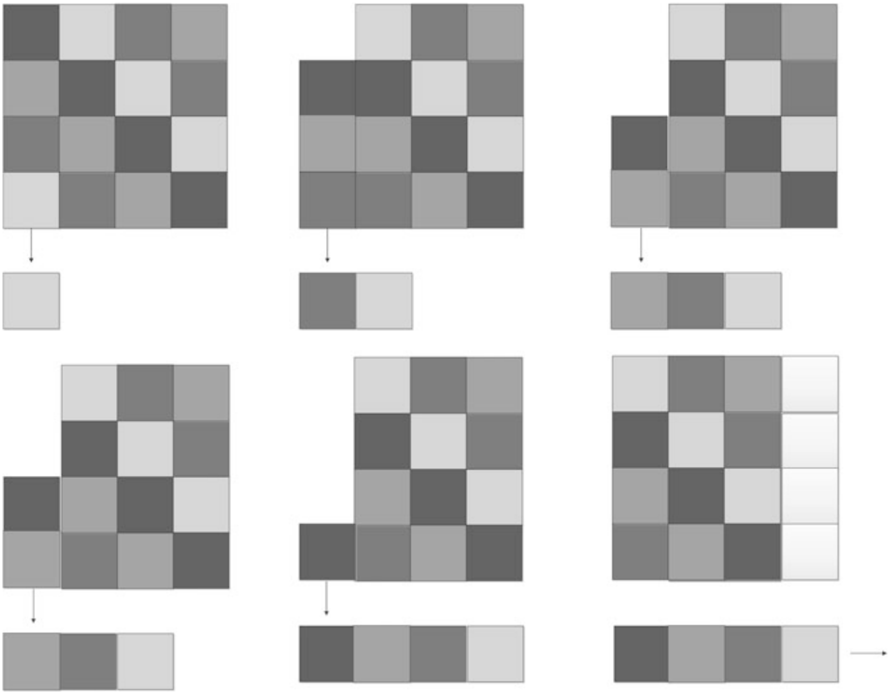

Fig. 7.2 Illustration of the read cycle of a CCD. An image is taken (*top left*) and the corner pixel is read. This pixel is then set to zero, and the charge of the pixel above it in the column is transferred to it, this is repeated until the whole column is moved down one pixel. The corner pixel is again read, and the cycles continue until the entire column is read and, therefore, empty. The adjoining column is then moved, and its adjoining column moved over etc. The Read cycle continues until all pixels are read

The process of reading the frame itself introduces errors. Using the read electronics in the system will have a tendency to add electrons to the system. This problem is exasperated if you increase the transfer rate, but also due to errors within the ADU itself. These errors, know as **Readout noise** are largely but not entirely dealt with by the subtraction of the bias calibration frame discussed in Sect. 7.5.

When we apply the gain to the electron count we, in most cases will get a non-integer number. However, as previously mentioned, the ADU is integer only so fractions are lost, either by a rounder processes or by just being intergerised so that at best the ADU will report the count with an uncertainty of ± 1 unless of course the gain is one, this is known as **digitisation noise**. Compared to other sources of noise, digitization noise is not that significant. However, both the pixels and the ADU tend to have a non-linear response as they approach their limit just as the photographic film becomes no linear. Non-Linearity is a major problem when imaging as a pixel whose count is in the non-linear range is effectively useless for most scientific uses. Nor is it immediate apparent that a pixel is in the non-linear range without first identifying the non-linear range and determining if the pixel lies within that range. In Sect. 7.6 you will go through the process of the identifying the non-linear range for your CCD.

When exposing the CCD to light, a number of thermal electrons leak into the pixels. These cannot be distinguished from those liberated by the interaction with a photon and can become a considerable source of noise. This **Dark Current** is both temperature and exposure time dependent and take the form of a Poisson distribution that we discuss in more detail in Chap. 5. The dark current should increase linearly with exposure time hence if you are undertaking a long exposure of a faint object the dark current may become very significant if left unaddressed. Likewise, an uncooled camera will have a significantly larger dark current than a cooled one. One of the most common reasons for bad image quality is either the failure to run the cooler or failing to deal with the dark current by applying a calibration frame as discussed in detail in Sect. 7.5.

The last source of noise we will discuss that is directly, although not wholly, associated with the CCD is flat field noise. In an ideal set-up, the amount of light falling on the CCD and the response to that light should be uniform, or flat. However, as we have discussed earlier, there are a number of sources of error that are structural in nature, such as variability in QE across the CCD and CTE. In addition, there is structural noise within the optics which contributes to the overall flat noise problem. These problems are addressed by **Flat Fielding** which is discussed in Sect. 7.5. However, even the best flat field cannot entirely deal this structural noise and for some instruments, such as space telescopes, flat field noise can be the dominant noise source.

7.3 Binning and Sub-framing

Most modern CCD camera will enable you to perform a function known as on-chip **binning**. Binning is a process whereby neighbouring pixels are joined to produce larger pixels. Most cameras will, when performing this function, average the count for each the bins. In which case if we have 2×2 binning with four pixels responding with a count of 450,445,430,424 the single pixel produced by binning will report a count of 437, remembering that a count is an integer value. Not all cameras do this, and some will report a value of 1,749 so be sure you know what binning process the camera uses before implementing it. On-chip binning happens before the data is downloaded. Binning not only reduce the number of pixels but also reduce the resolution of your final image. However, in general, most CCD/telescope combinations are over resolved so that a small amount of binning does not affect the resolution. Notwithstanding the loss of resolution, the On-chip binning has two major two benefits. Firstly, it reduces download time; this use useful when observing very transient events such as occultations. Secondly it reduces noise partly because of the reduced download time and the noise being averaged out over a number of pixels.

An alternative to on-chip binning is off-chip binning (or post-download binning). Off-chip binning is a software action and is performed after download, so the benefits of reduced download time and the associated noise reduction is lost. However, the noise gain from averaging the noise over some pixels is still obtained and for CCD that do not support on-chip binning, off-chip binning is a suitable alternative.

When performing on-chip binning, the calibration frames should have the same binning. Binning is not always a linear process so you cannot just turn an unbinned calibration frame into a binned one by using the software. Likewise, when performing off-chip binning the same process must be applied to the calibration frames before subtracting them.

An alternative to binning is sub-framing. Sub-framing is a process were only part of the CCD is read. Sub-framing is a software driven process and is not always supported. It involves taking an image of a target and selecting a region within the downloaded frame. Once that region is selected only those pixels in the selected region are download in future exposures. Hence with sub-framing no resolution is lost but you do not gain the noise reduction associated with binning. It is mostly used when download speed is needed, such as in focusing. The University of Hertfordshire's Observatory uses a fast CCD camera to observe asteroidal occultations of stars. These are very transient events often only a few seconds in length. A typical astronomical CCD has a download time of a few seconds so are not suitable for precision timing of such events. If we consider that the occulted star may only fill 100 pixels, by sub-framing, so that we only download the pixels with light from the target we can significantly improve the download time and hence the cycle time of the camera. In fact, this camera has been run at a thousand frames per second for such observations.

7.4 The FITS File Format

The Flexible Image Transport System (**FITS**) is the standard format for astronomical images and image-derived data such as spectra and may have either the .fit, .fits or .fts extension to the file name. The FITS file format is designed to carry metadata (data about the data), two-dimensional image data and three-dimensional data such as radio and millimetre data. The **FITS Header** holds the metadata in a series of labels known as **cards**. Each card will have a label stating what it is, for example, EXPTIME, exposure time, and a piece of data for example 10. This card tells the user that this was a 10s exposure, but it also tells your software it was a 10s exposure. You can add and change header cards using any number of astronomical packages. Just remember if you use something that is a standard card name it might cause problems later. Robotic telescopes, as well as taking the image, run the data through a **pipeline** a series of operations on the image. All these operations are noted by the insertion of a card, or cards, in the header. The addition of these cards allows the user to identify what has happened to the image and allows the processes to be reverse it needed (which isn't always possible).

The FITS header may also contain information about the translation between pixel location and astronomical coordinates systems such as Galactic. The translational system is known as the World Coordinate System (**WCS**). If the application you are using has the ability to display pixel WCS position but does not for the image you are displaying, it is likely that the WCS has not been calculated for that image. This is easy to resolve as we will show later but not always necessary.

Unlike other image formats such as JPEG or PNG, FITS files are not compressed, so they tend to be quite large; however they contain all the data taken by the camera, unlike more traditional imaging formats. If you are familiar with digital photography you may have noticed that you DSLR camera can save images in RAW format; when you select RAW the number of images you can store on your memory card is much reduced. This is because RAW is the terrestrial version of FITS. Most professional photographers will take images in RAW as they can then be easily manipulated. For the same reason, professional astronomers use FITS, as should you. If you need to make a JPEG image from a FITS, it is fairly straightforward, and most FITS display applications have a Save As option. You should, however, always keep your science data in FITS.

FITS is support by a wide range of computer languages with libraries available in C, C++, Perl, R and Python as well as many others. Using code to manipulate FITS files is very straightforward and very powerful.

7.5 Calibration Frames

Read noise can be easily addressed by taking a **Bias Frame**. A Bias Frame is an exposure of zero seconds, with the shutter closed. Hence, it forces the camera to go through a read cycle while not exposing it to any light. Therefore any electrons registered have come, not from photoelectric interaction in the CCD, but due to random fluctuations introduced by the read process, electrical components in and nearby to the camera and to a very small extent cosmic rays (see Sect. 7.7). When you take an image of your target, a **light frame**, the bias frame should be subtracted from your light frame. This subtraction process normally happens automatically on modern camera control software. When using a previously generated bias frame (and there is nothing wrong with doing this as the variation of the bias over time is very small), it has to be taken as the same camera temperature as the light frame. Subtracting the bias frame removes any read noise, as well as the pedestal that was added during the read process. Although pixel values within a bias may be quite high, depending on the camera and the software you are using, the standard deviation across the array should be very low, perhaps only a few counts. If it is not, then you might have a read problem or a faulty piece of equipment, For example, a bad camera cooling fan. In general a good bias frame is a median of a very large number of bias frames. Taking the median reduces the amount of random noise associated with individual frames and address to some extend cosmic rays.

With so many pixels on a modern camera, the camera on the Gaia space telescope have over a billion pixels, clearly there is going to be some random noise getting into the pixels. To address this, we take a **Dark Frame**. A Dark Frame is an exposure, with the shutter closed, of a duration equal to the light frame. The Dark Frame, as with the bias frame, is subtracted from the light frame to remove any inherent problem with individual pixels. Darks should be fairly linear, i.e., if you double the exposure time you should double the count within the pixel. You also should by now have noticed that a dark frame must also include a bias frame so, in general, the removal of the dark also removes the bias, but not always. Again it depends on the software you are using. You should also notice that a dark must only be linear after the non-linear bias has been removed. Again, as with bias frames, dark frames should be taken at the same temperature as the light frame. Typically they are taken either immediate after or before the light frame is taken and in many cases the software you are using will automatically subtract both the bias and dark frames from the light frame.

Not all pixel are created the same, even on the best science grade array some are more sensitive some less so. The hole in the secondary mirror of a Cassegrain telescope, the secondary, dust on the filters, the correcting plate or the CCD window and non-uniform illumination of the chip due the telescope aperture being circular and the array being rectangular all need to accounted for. This is done by taking a **flat frame** which unlike biases and darks are used to divide the light frames. require the camera being exposed to a uniform level of light until the median pixel count reaches a set limit (for most 16-bit camera this will be around 30,000 to counts). This has to be done for every filter mounted on the camera. Normally multiple flat fields will be taken, and their median value for each pixel used to create a science flat. Typically flats are taken during twilight using a position opposite the Sun as this is consider the flattest (i.e. most uniform) part of the sky, with multiple flats taken and their mean value used as the master flat for that filter., Flats frames taken in this manner are known as **sky flats**. Very good flats are extremely challenging to achieve, and because they account for the contamination of the optics, they are needed to be updated much more often than dark or bias frames. Hence, there is now a tendency to do **dome flats**. Dome flats use the illuminated inside of the telescope dome as their light source, rather than the sky. Dome flats may be done at any time and are more consistent because the environment they are taken in can be controlled. However, they do not cover the entire optical pathway (because they don't include the atmosphere), and they may not be so uniform as sky flats. There is currently considerable debate about sky flats versus dome flats, the methods used to produce skies and domes and their impact on many aspects of observational astronomy, as will be discussed later.

When taking an image of the object, the resulting image is known as a **light frame**. You will need to subtract the bias and dark frames from you light frame and divide the result by your flat frame (minus its bias and dark component) to create a **science frame**. Many professional astronomical telescopes will do this as part of the **Pipeline**, a series of computer programs that turns an astronomical image into a usable science image (Figs. 7.3, 7.4 and 7.5).

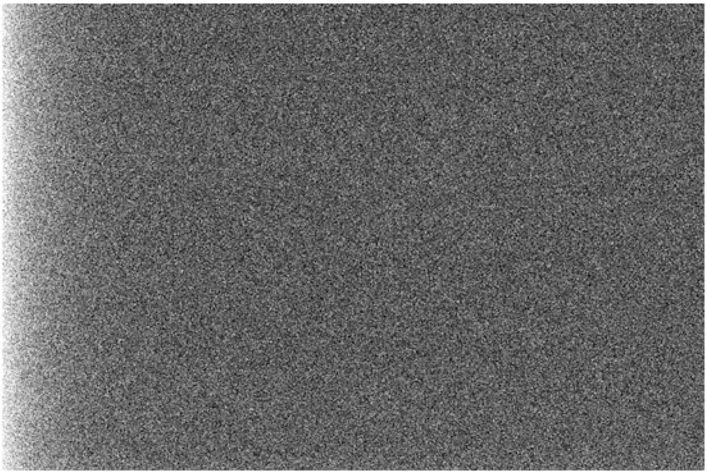

Fig. 7.3 Bias frame from an SBIG 1301 CCD. The *bright edge* on the left-hand side is the noise generated by the read process (Image University of Hertfordshire)

Fig. 7.4 Dark frame from an SBIG 1301 CCD. The *bright corner* on the left-hand side is the noise generated by the read process. The bright points scattered randomly across the frame are hot pixels, as discussed below (Image University of Hertfordshire)

Super Flats and Fringes

In some cases it is necessary to take more advanced calibration frames, for example when it is not possible to do dome flat or when very high precision is need, in which case a **Super Flat** may be required.

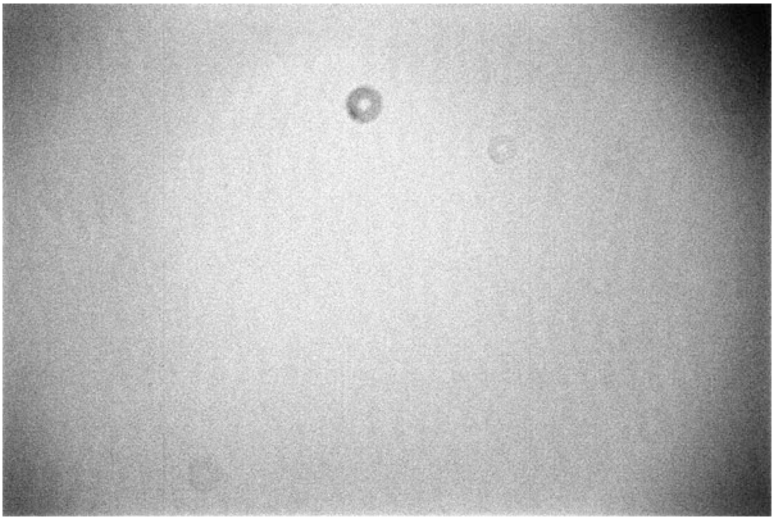

Fig. 7.5 Flat frame from an SBIG 1301 CCD. The gradient across the frame is partly caused by read noise but also by the chip not being uniformly lit. The *dark circles* are out of focus dust in the optical path. The *very dark circle* is dust on the CCD window (Image University of Hertfordshire)

In simplest terms a super flat is a normalised flat frame constructed by taking the mean of several on sky images which contain very few stars (none is preferred but effectively impossible) and no extended objects such as galaxies or nebulae. The stars in the field can be removed using a sigma clipping algorithm which removes pixels with values outside a set standard deviation and replaces it with either the mean value of the image or by the average of the eight pixels surrounding the pixels being clipped. Super Flats, as the name suggests are very flat and should be flatter than both dome or sky flats. However, modern dome small telescope flat fielding techniques using photo-luminescent have become very sophisticated and may be very close to super flats in quality.

Fringes are an unwanted interference pattern caused by long wavelength photons passing through the thinned substrate of back-illuminated CCD. The pattern has been likened to waves on water. You can think of backlit CCD as the reverse of front lit. Whereas in front lit CCD the wiring for the CCD is on the exposed surface of the chip, for backlit CCD it is at the rear and the light has to pass through the substrate on which the CCD is mounted in order to be detected. Backlit CCDs have some advantages such as being cheaper to manufacture and being more sensitive to light. More modern backlit CCD have reduced fringing, and it is unlikely that you will need to address it (although applying a super flat is one way of doing so).

7.5.1 Practical 1: The Bias and Dark Frames

In this, the first of the practical sessions, you will learn how to take dark and bias frames. This practical can be undertaken in daylight and in poor weather as the dome does not need to be opened, and the lens cap stays on the camera. It will give you a first introduction to your camera control software and the FITS Liberator Package.

Leaving the bias and read noise aside for the moment, as we take progressively longer exposures, even with the shutter closed we find a mysterious build up of counts within the pixels. These are thermal electrons produced by the CCD itself entering the pixel and are one of the main sources of errors in CCD based observations. In this practical you will investigate both the dark and bias frames in order to understand their impact on uncertainties in CCD based observations.

7.5.1.1 Taking the Frames

1. Set up the telescope by ensuring that no light is entering the camera by placing the lens cap on the telescope. The telescope does not need to be powered up and does not need to be tracking. Connect to the camera using your camera control software.
2. Using the camera control software, make sure you have the binning set to 1×1 and that you have sub-framing turned off. Turn off the cooler and let the CCD warm up to ambient temperature. If your control software has a warm-up option, you may use this to save time.
3. Set the camera temperature just under ambient, so if the uncooled camera is at 12 °C set it to 10 °C. If your camera control software has the ability to select between frame types, then select bias. If your software can do auto-calibration make sure this is also turned off. Once your CCD is down to temperature and is stable at that temperature, take five zero second exposures saving each frame. It would help if you saved each frame with the temperature as part of the name. Your camera control software should put the temperature in the header, but it might not.
4. Once you have taken your first set of bias frames, reduce the temperature by five degrees. Again wait until the temperature is stable and then take a further five frames. Continue this process until the camera reaches either its normal operating temperature or it reaches the point where the cooler is running at 100 % in order to maintain the temperature.
5. You will now need to take your dark frames. Warm the CCD up to ambient temperature and as before cool it down to just a few degrees below ambient. If you have the ability to set the software to dark frame, do so.
6. Once the temperature is stable take five dark frames with an exposure time of 1s.

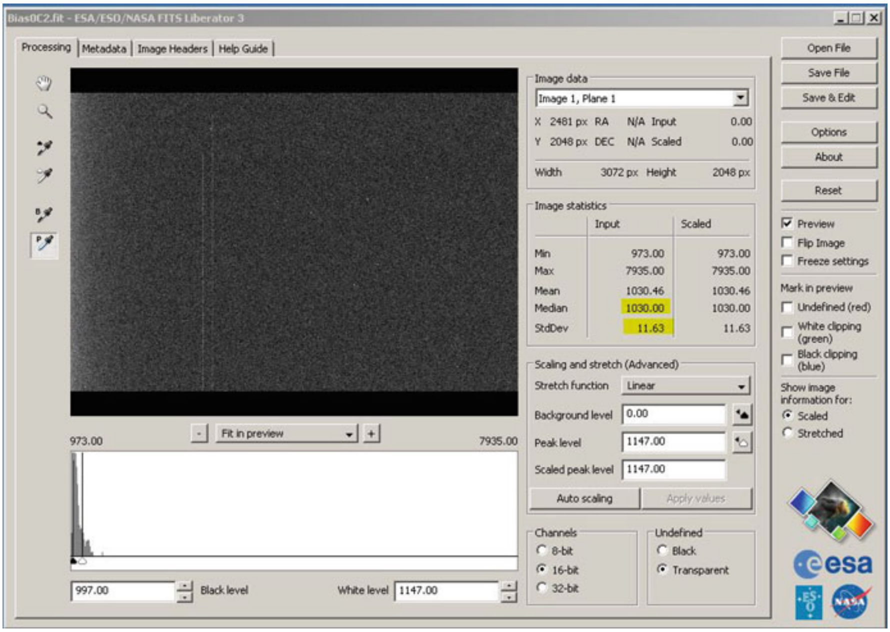

Fig. 7.6 The above figure illustrates a Bias frame loaded into FITS Liberator. The Median and StdDev figures required are highlighted

7. Increases the exposure time to 5s and take and save a further five darks. Continue this process for 10, 30, 60, 120, and 240 s. Once this is done, cool the camera down by 5° C and repeat the cycle. Take 1, 5, 10, 30, 60, 120 and 240 s exposures at the new temperature.
8. Continue this cycle down to the same temperature that you took your last bias frame.
9. Once you have downloaded all your frames open the each frame in FITS Liberator (Fig. 7.6). The image statistics panel on the right-hand side, next to the image shows the Image Statistics. Record the Median and the StdDev from the Input column for each image. This represents how the median pixel value in the array and the standard derivation of the array, i.e., how much it varies. Do this for all you images and record on spreadsheets as two tables, with one table for the bias frames and one for the darks. You can used the Image Header Tab in FITS Liberator to confirm the temperature that each image was taken at and the exposure time.
10. For each block of five images, for both the bias and dark frames take the mean of the medium count values and of the temperature making a note of the spread of the median count values (the maximum value minus the minimum value) as will be used for your uncertainties. Use these to make two new tables one for the bias frames and one for the dark frames.

11. Looking at the bias, plot mean bias median value (y) against mean temperature (x). Your error for each point should be ...

$$\pm \frac{\Delta x}{2\sqrt{N}} \tag{7.1}$$

where Δx is the spread of that point, and N is the number of measurements in that point, i.e., in this case, five.
12. Think about what the point at which the line crosses the y-axis represents and if the temperature has a large impact on the bias? From you Standard deviation data, does the bias frame vary much across the CCD? If it does, what causes this?
13. Moving on the dark frames. For each exposure, time plot means bias median value (y) against mean temperature (x) using Eq. 7.1 for you error bars. Do this for each exposure time, plotting all seven exposure times on the same graph if possible.
14. Remember that the bias is part of the read process, and so it appears in the dark frames. You will need to address with this in your analysis.
15. If you have access to another camera, and particular a different model of camera and has the time, repeat the process on a second camera.

Analysis

You should write your report up in standard laboratory style, Introduction, method, results, and conclusions. You should include all your tables and your plots as wells as notes about the camera and software used in addition to any problems you may have encountered. Your report should include...

1. A description of the shape of your curves and an interpretation of how the curves vary with exposure time and temperature?
2. What was your bias level, how does the bias vary with temperature and internally across the CCD?
3. How does the dark current vary with the exposure time and temperature (don't forget there is a bias in your dark, so you need to subtract the effects!)? How does the dark frame variation across the CCD compare to that of the bias frame?
4. How significant is the dark current when compared to the bias current and how might we deal with both the additional counts caused by the bias and dark and their associated uncertainties?

7.5.2 Practical 2: Flat Frames

In the second practical you will be taking flat frames. Flat frames are a vital part of the calibration process and are the calibration frame type that needs to be updated the most often. As you will see, they are also a primary contributor to uncertainties.

In this practical you will familiarise yourself with the flat field method, the operation of the camera and the use of SAO DS9. The analysis will give you an understanding of both the importance of flat fielding and the impact that poor or irregular flat fielding has on image quality.

Flat Fielding

1. Turn the dome flat field lights on. These might just be the normal white lights but do not use the dome red lights if the dome has them as you want the light going into the telescope to be as spectrally flat as possible. Take the lens cap off the telescope.
2. Rotate the dome so that the flat field board is facing the telescope. You may have to move the telescope in alt to get it directly facing the flat field board. Rather than power the telescope down, turn it's tracking off. If your telescope has a flat field diffuser put this in place.
3. Power on the camera, connect the control software and put the cooler on and set to the correct temperature. While you are waiting for the camera to get down to operating temperature make sure the binning is either off or set to 1×1, the sub-framing is off as is any auto calibration.
4. Change the filter to R or ŕ and take a short exposure. If you camera control software has the option to select the frame type as "flat" select it.
5. Take a very short exposure, say not more than 1s. Once it is displayed on the camera control software check to see the mean count near the centre, In MaximDL this is done by bring up the information window (ctrl-i) and selecting Aperture from the Mode drop down. If you are not using MaximDL, it is likely that your software has a similar capability. Failing that you could open the image in FITS Liberator and use the Mean value in the image statistic box. You are trying to obtain a mean count of around 30,000–40,000 for a 16-bit camera. If you are getting too many counts reduce the exposure time until it is the correct range. Likewise if the count is too low increase the exposure time.
6. Once you are happy with your exposure times, take five flat fields and save each one. Don't forget to record your exposure time in you lab book.

   You can perform the analysis of your flat field images in DS9. You will need to determine how flat your field is. To achieve this, you need to measure the average counts in a number of small regions spread over the field. This cannot be done with FITS Liberator hence the need to use DS9.
7. Open your first flat in DS9 and set the scale by using the scale menu select 99.5 %. and using the Zoom menu drop down select zoom to fit frame.
8. You will now need to set up you regions. From the Region, menu select Shape and then Box. Double click on the centre of the image and a box will appear. Single click on the Box and drag it out, so it is square and has a width about an eighth of the width of the frame.

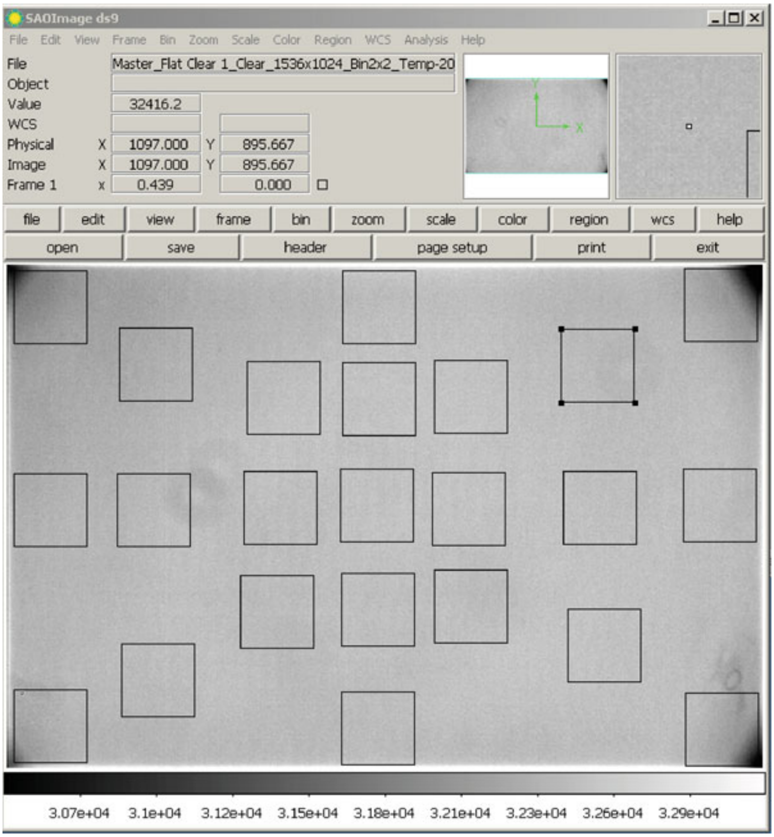

Fig. 7.7 Example of the DS9 region layout for the flat field practical

9. Move the Box to the top left-hand corner of the frame. From the Edit, Menu select copy and then paste (or press ctrl-C and then ctrl-V). This is creating a second Box co- located with the first. Drag this Box to the bottom left the corner. Repeat until you have something like Fig. 7.7.
10. From the Region menu select Select All (or press ctrl-A). This selects all the regions. Now again from the Region menu select Save Regions. Enter a file name and the select ds9 for the format and physical for the coordinate system. This will save you region layout for future frames.
11. Pick a region and click on it. This will bring up some of the information on the region including its centre. Record the centre making sure it is physical units, not, for example, RA and Dec. Use trigonometry to determine it's distance from the centre of the array (Hint you can find the array size in the header). With the region information box, open select apply and a second box with more information on the region. Find the mean count and record on a spreadsheet along with the region's distance from the centre.

12. Repeat this for the remaining four frames, although this time you can load your saved regions in, which is quite a time saver.
13. Plot distance against mean count on a spreadsheet with each frame on the same graph.

Analysis

Write your report up in standard laboratory style, Introduction, method, results and conclusions. You should include all your tables and your plots as well as notes about the camera and software used in addition to any problems you may have encountered. Your report should include. . .

1. A description of the overall appearance of your flat field image, including the small scale and large scale uniformity. Are there any unusual features in the flat field, if so what do you think they are caused by?
2. Your flat fields include a Bias and a Dark current. Given what you know about these two calibration frames, how do the errors they introduce compare to the flat field?
3. There might be variation between the edge of the flat field and the centre along each of the axes if there is can you explain it? What would the impact be if we image a star in the centre of the field and then at an edge? Can you quantify this effect?
4. When the observatory staff take flats they take multiple copies and combines them into a master flat. How do you think this is done. Are they just added?
5. Dark frames are temperature and binning sensitive. Are light frames also temperature and binning sensitive? How do you think the chose of filter affects the flat? Do you think we need to flat field for every filter, or will just one do?

7.6 Linearity of CCD

Unlike photographic film, which does not respond in a linear manner when exposed to light, the response of a CCD is linear over most of their dynamic range (around 1000–65,000 counts for a 16-bit cameras). Hence doubling the exposure time doubles the count in an exposed pixel. However, when the counts begin to get high, such as with very long exposures or very bright objects CCDs tend to become progressively nonlinear before become saturated and no longer responsive.

It is important to determine where this point for non-linearity lies so it can be avoided. This is normally determined by using a modified version of the flat field technique outlined in Practical 2.

7.6.1 Practical 3: Measuring CCD Linearity

In this practical you will be investigating how the properties of telescope design and the CCD detector can affect the accuracy of the results you obtain from your astronomical imaging observations. This practical can be undertaken either with the dome closed or by imaging a star the choice depends on weather conditions. The point of loss of Linearity of a CCD is one of the most import components of photometry, the measuring of a star's brightness and it is important that you understand why it happens, where it happens and how to avoid it.

Taking the Data

1. Turn on the dome lights. If you can dim them turned them down enough that you can still see the control computers keyboard.
2. Rotate the dome so that the flat field board (or if there isn't one just the inside of the dome) is facing the telescope. You may have to move the telescope in Alt to get it directly facing the flat field board. Rather than power the telescope down, turn it's tracking off. If your telescope has a flat field diffuser put this in place.
3. Power on the camera, connect the control software and put the cooler on and set to temperature. While you are waiting for the camera to get down to operating temperature make sure the binning is either off or set to 1×1, the sub-framing is off as is any auto calibration.
4. Select the narrowest filter. In most cases, this will be either OIII, Hydrogen α or and SII Filter. If you do not have, any of these in the filter wheel use an R or B (but not a V).
5. Take a 1s exposure, save it and open it in DS9. Scale the image to 99.5 %. Create a circular region, it needs to be fairly small say 100 pixels in area, and locate it near the centre of the field. You will need to save the region so you can load it into future frames. Look at the mean count within that region. If it is over 30,000 your exposure is too long or your dome lights to bright. Try reduce them until you have 5,000 counts or less.
6. Take a succession of light frames, increasing the exposure time with each. I suggest 2, 5, 10, 15, 30, 60, 90, 120, 180 and 240 s exposures. After you save each one, open it in DS9, load the region you saved in step 5 and make a note of the mean count. You can stop when the mean count hits about 64,000 (for a 16-bit camera), at which point the CCD is, of course, saturated.
7. Plot the mean count against exposure time. You may wish to try fitting a line to your data. If you have undertaken practicals 1 and 2 you will have some idea of the size of your error bars but remember that you are plotting the mean count, not the actual count.

Analysis

As previously, you need to write your report up in standard laboratory style. You should include you table of your exposure times, mean counts and your plot as well as the standard notes about the camera and software used in addition to any problems you may have encountered. In particular your report should include...

1. In what circumstances is the detector response of CCDs preferred to that of photographic film?
2. Describe the shape of your plot in particular how the gradient changes with exposure time if it does.
3. Try to explain the significance of any change in the gradient and its impact of performing precision measurements such as those need to detect exoplanets using the transit technique
4. Your plot may not be straight through the whole length. Can you suggest some a process that might cause this effect.

7.7 Bad Pixels and Cosmic Rays

It is unlikely that any astronomical will have every pixel working, and more pixels will be lost as the camera ages. The most common fault is a dead pixel, a pixel that will either not read or is no longer sensitive to light. Dead pixel will read as having either zero counts or a count near the bias level and are difficult to see in an image. Hot pixels are pixels that read artificially high, either at or close to close to the maximum value (65,536 for a 16-bit) camera and appear as bright, isolated points, in the image. You may also see dead columns. These can be caused by a faulty pixel preventing the read process in that column. These faults may cause problems when trying to extract data from the image, so you need to be aware of them. Most astronomical imaging software has a function that attempts to deal with hot and cold pixels although no process, besides physically avoiding that part of the CCD, is totally satisfactory. The most common method works by identifying pixels with values under or over a set value and replacing the faulty pixel with the mean or median value of its neighbours.

A CCD cannot distinguish between photons and other high-energy particles such as those emitted by radioactive decay naturally from the surrounding environment. Given that the CCD is effectively shielded few actual, local events, rarely have enough energy to penetrate the casing and register. However, cosmic rays, very high energy particles, many thought to be from supernova explosions, stream into the Earth's atmosphere, causing cascades of high energy high energy particles that pass through the Cassegrain aperture of the mirror and dumping their energy into the CCD. This cause a trial of high registering pixels that appear hot when read, which of cause they are not. Cosmic ray detections on CCD images are difficult to remove using standard methods as they appear at random and tend to leave a jagged trail. This means that the median of the neighbour process above will not produce a satisfactory result.

7.7.1 Practical 4: Bad Pixels and Cosmic Rays

In this practical you will use the CCD as a cosmic ray detector. In order for this to be effective you need to remove the hot pixels that themselves appear like cosmic rays. This practical is written for Diffraction Ltd.'s MaximDL package but you could use any suitable package.

Measuring the Cosmic Rays

1. With the dome closed and the lens cap on attached your control software to you camera.
2. Ensuring the camera is down to temperature, that the binning is set to 1×1, and there is no sub-framing, take two 15 min dark frames and save them as Frame 1 and 2
3. Subtract one frame from the other
4. the resulting image in DS9 and identify the cosmic rays in the image by putting a region on them.
5. Using the Hot pixel tool find the number of pixels affected by cosmic rays.
6. Do this for a number of differing exposure times and plot cosmic ray count against dark exposure time.

Analysis

Your analysis should include...

1. Your images including the images with the cosmic rays marked. Also include the results of the Hot pixel tool and your plot.
2. How do cosmic rays differ from hot pixels?
3. Given the size of the array how proportionally how many pixels are affected?
4. How does the exposure time affect the number of cosmic rays. Why do you think this is?
5. When undertaking precision photometry the appearance of cosmic rays have to be address. Given that dark frame subtraction is not effective, can you suggest another way of dealing with cosmic ray strikes on CCD.

7.8 Gain

As stated previously each pixel on a CCD is, in effect, a small capacitor storing electrons liberated by the effect of an incident photon, the value that is actually read is the number of electrons within the well. However, what is reported is the count. A pixel can often hold in excess of 100,000 electrons which is more than a 16-bit

Analogue to Digital converter (ADC) can report. Hence, the number of electrons is scaled. This scaling factor is the Gain G. Your observatory staff should be able to tell you the Gain of the camera you are using, and it should also be recorded in your lab book and your images as part of the FITS header.

As always, no two cameras are created equal, and there may well be a slight differential between the specified gain and the actual gain. In an ideal world, the gain should be the maximum well capacity in electrons divided by the maximum number the ADC can hold. In reality, the gain is seldom this high. In the next practical, we will determine the Gain of your camera.

7.8.1 Practical 5: Measuring Camera Gain

In this practical you will find the gain of your camera. You will need the camera gain when calculating uncertainness in your observations. If new the camera may very well be operating near the manufactures published level but with age it may move away. You will be using the skills you have developed during the previous practicals in this book.

Taking and Reducing the Gain Images

1. With the camera down to temperature and the lens cap on the telescope, take a series of darks with 1, 2, 3, 4 and 5s second exposures. Take two darks for each exposure time. Ensure that the binning is set to 1×1 and sub-framing is off. If you have the Dark frame option on you control software turn it on.
2. Take the lens cap off and put the doom lights on. Set the telescope up for taking flat frames as previously.
3. Take pairs of flats frames for 1, 2, 3, 4 and 5s exposures. Start with the 5s exposures. Using FITS Liberator (or another appropriate package) ensure that you are not entering the non-linear region of the CCD, hence the need to start with the long exposures. If you are entering the non-linear region swap to a different filter and try again.
4. Using an appropriate package (for example MaxImDL or a python script) generate a mean image of each of the pairs of dark frames and subtract these from the flats of the same exposure time.
5. Taking each pair of images subtract one of the pair from the other to make a variance image.
6. For each of you five variance images open them in FITS Liberator and find the mean value and the sigma value (σ).
7. Plot (in Excel for example) $\sigma^2/2$ against the mean and then fit a straight line to your points

Analysis

When writing up your report in the standard format, you should include…

1. What feature of you plot do you think indicates the Gain. How does it compare to the Gain reported by the manufacturer.
2. How does the Gain measures compared to the maximum Gain?
3. As you subtract the pairs of flats to create your variation image, do you think it makes a significant difference that is subtracted from which and why?

7.9 Saturation, Blooming and Other Effects

Each pixel can take a limited number of electrons before it becomes full, this is known as **saturation** and should be avoided as much as possible. When a saturated pixel is read, the reading process causes electrons to flow into the pixels above and below the saturated pixel, often more in one direction than another. This in extreme cases this causes a spike to appear across your star, a feature known as **blooming**. Once an image has a bloom in it, the area around the affected part of the image is effective a loss as far as science is concerned. Some cameras have anti-blooming gates, which act as capacitors between the pixels, thereby stop stray electrons. In general, for science applications, the disadvantages of cameras with anti-blooming out weight those for. If you have a camera with this feature and you intend to take science frames (as opposed to just nice images), then turn the anti-blooming off if you can and keep the maximum pixel value well below the saturation point. You often see in press images from large telescopes, stars with crosses and halos around them. These are **diffraction halos** and **diffraction spikes** caused by light being diffracted around the secondary and the secondary supports. You should probably avoid these becoming too bright when doing science frames, as they will effect your measurements. You should also we aware that these are sometimes added after the image has been processed for artistic reasons.

Another problem that may be encountered with long exposures is **Residual Bulk Image (RBI)**, another effect of saturation. If a pixel becomes saturated, further photon strikes can cause electrons to move out of the pixel and into the substrate on which the CCD is mounted. When you move to another target, these electrons leak back into the pixel from where they originated and create a ghost image of the previously observed star. Some cameras have an IR flash that can help prevent RBI by pre-filling the pixels. For cameras without this feature removing of RBI is troublesome as it persist for some time (the author has seen RBI last 20 min). One way to fix this is to warm up the camera and to allow thermal electrons to displace the RBI electrons although the best way to fix RBI is, of course, not to let it happen in the first place by not saturating.

On occasion when taking an image you may see a strange, almost ray like feature, vignetted across the image. A very bright star, planet, or the Moon, being just outside the field of view of the telescope, but close enough that light is still entering the telescope is the normal cause. In general there is not an alternative solution for this other than not image under those conditions. You should not confuse this with dome clipping, where the edge of the dome is very near, or actually in, the field of view of the telescope and stray light from inside the dome is reflecting into the aperture. Nor should you confuse these with very straight single or double lines crossing your image, these are satellite and aircraft trails.

You might also see fractal-like structures appearing at the edges of an image. These are ice crystal forming on the window of the CCD. Within your astronomical camera, the CCD is sealed within an airtight, windowed box. The box is normally filled with an inert and very dry gas such as argon or nitrogen. Over time, water vapour gets into the window and can freeze on the inside. If this is happening, warm the camera up again, and then cool it slowly in steps. This should stop the icing, but tell the observatory staff, as they will want to address this issue as it might lead to CCD failure.

Chapter 8
Imaging

Abstract Astronomical imaging using a telescope and the CCD camera is the foundation of modern astronomy. We show through practical examples, how to plan, take, and calibrate astronomical images for use in science. We work through the operation of the telescope including pointing, focusing and camera operation. We discuss problems associated with non-stellar sources such as galaxies and nebula and planets.

8.1 Introduction

Imaging is one of the four corner stones of observational astronomy. The others being photometry, astrometry and spectrography. We can consider this to be; What does it look like, how bright is it, where is it and what is it made of? Imaging is an important part of photometry, astrometry and spectrography as it is the collection of photons that make these aspects possible. When undertaking imaging we need to understand what we are trying to achieve. If we wish to do perform photometry then we need to keep the camera within its linear range, which is not so important with astrometry, although we should avoid saturating the target. With spectrography we will be using a spectrograph rather than an imaging camera. In spectrograph we are unlikely to over exposure but we will need to undertake careful guiding to keep the target within the spectrograph's slit. If we are imaging to determine morphology of a target, then a long an integration time may be needed, whilst avoiding saturation, and taking multiple images which are combined and manipulated to improve the visualisation of the target.

First light is the term used for the point where a telescope takes its first image, the commissioning image. Although it is unlikely you will be the first to use the telescope this might be the very first time you have used a telescope to gather photons from an astronomical target. The previous chapters should have prepared you for you first observations, this chapter will guide to through that first observation from the planning to the data reduction. The chapters that follow this one assume that you have successfully completed the practicals that follow in this chapter, or at the very least have experience in using astronomical telescopes and CCD.

M. Gallaway, *An Introduction to Observational Astrophysics*, Undergraduate Lecture Notes in Physics, DOI 10.1007/978-3-319-23377-2_8

8.2 Planning

As stated previously telescope time is a precious commodity, and you should plan your evening observing in advance to get the maximum out of the telescope.

If you are observing as part of a practical, you should read the practical in advance and understand its aims and processes. Reading or rereading the practical in the dome is a waste. You should know what telescope you will be using and its optical characteristics especially the pointing limitations, the limiting magnitude and field of view as well as the filters available

You will either have to choose your target from a supplied list of options or have to select a target that meets a set of criteria or the target may might have been stated explicitly. You will need to ensure that the target will be visible at the time and date you will be observing, and that is within the limiting magnitude of the telescope. Remember that for extended objects the quoted magnitude is an integrated magnitude. If you know the approximate area of the object you can get an idea of its point source magnitude that you need to know to determine if the target is bright enough to be observed as well as gauging the exposure time. You should write down the targets name, the RA and DEC in J2000 and its magnitude in the band you are observing. If you are not asked to use a specific object then it always advisable to have a backup object. If the Moon is going to be up it is best to avoid the region of the sky near the Moon. Some telescope control applications have a Lorentzian Moon avoidance function built in, which warns if the telescope is pointing too close to the Moon, depending on the phase. If the Moon is full and you are doing photometry or deep sky imaging, you may wish to observe at a later date which applies if there is a bright Solar System object nearby.

For each target, you should produce a finder chart. A finder chart will help you confirm that the target you are looking at is the target you want to be looking at. Most star fields are indistinguishable from one another. Even if you target is obvious, such a galaxy, having a finder chart that is larger than the telescope field of view helps. If the telescope moves close to the target but not close enough to put in the field of view, a finder chart can help in finding where the telescope is pointing and how to move it so that your target is in the centre of the field.

There are a number of finder charts tools to help you find your target. Perhaps the most popular in the professional community is Aladdin, which can be accessed via SIMBAD. Additionally, DS9 has an image server option that allows the downloading of survey images. The DS9 image servers can be found in Analysis ≫ Image Server, where are a number of image servers listed; we suggest using one of the DSS (Digital Sky Survey) servers as there are mostly optical images. You can search by object name or coordinates as well as determining the image dimensions. Before you print off the finder image, I suggest inverting it so that the stars appear black against a white background as this make annotation easier. Mark the target on the finder chart and mark the orientation of the axis. An example of a DS9 finder chart is shown in Fig. 8.1.

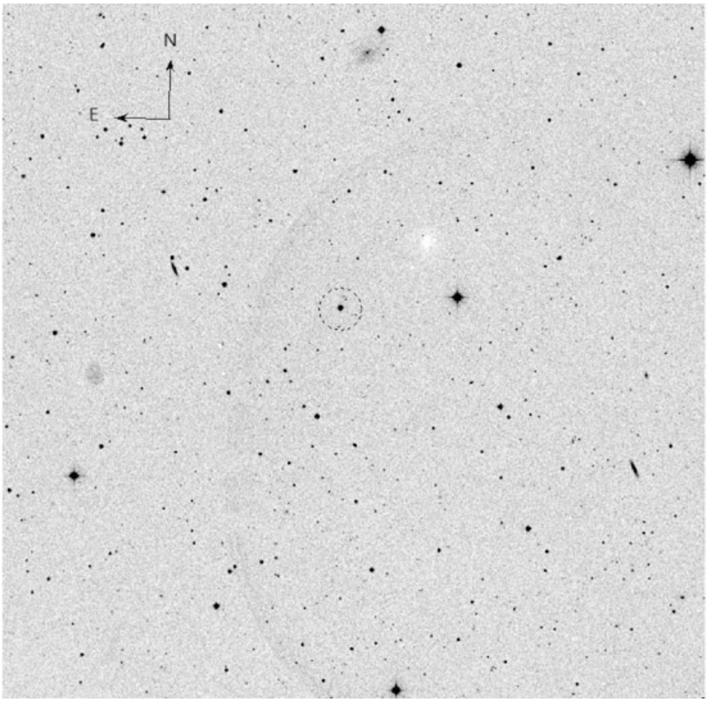

Fig. 8.1 Example of a simple finder chart using the ESO-DSS image server and DS9. Note the orientation arrows (Image European Southern Observatory)

The best time to image an object is when it crosses the meridian. At this point, you are seeing the object through the minimum amount of atmosphere for that object for that night, which will reduce the need to refocus. If you try to image each object as it passes through the same part of the sky, it will reduce the amount of time your telescope is slewing. Telescope processes that involved using the telescope when it is not imaging such as downloading, slewing, swapping filters and focusing are known as overheads. Always try to minimise the ratio to observing time to overhead time, thereby keeping efficiency high, you should do the same.

Although you can work out the meridian crossing times for each object you intend to observe, it is time-consuming and is easiest done with a piece of planetarium software such as Stellarium or The Sky. Open the planetarium program and set the time and date for the start of your observations. Using the software identify where in the sky your targets are and work out the best order in which to observe them, remembering that some will be setting while others will be rising.

8.3 In Dome Pre-observation Preparation

Before you make your first observation proper, you will need to take calibration frames, check the pointing of the telescope and focus up. These are a vital part of the observation processes, and their importance should not be overlooked.

8.3.1 Calibration Frames

Before you leave the observatory, you will need to create your calibration files that were discussed in Sect. 7.5. You may do this either before or after imaging. If you arrive and it is cloudy this is an ideal opportunity to take calibrations frames, if it is clear take the opportunity go to straight to imaging and take calibration frames at the end. A summary of the calibration process is shown in Fig. 8.2. As discussed earlier many observatories like to take **sky flats** which use the twilight sky as their light source. However for a small observatory the amount of time to get sky flats for narrow band filters is long, and it might not be possible to take all the required sky flats in one session. An alternative is **dome flats** where the inside of the observatory dome illuminated with the dome lights is used as the light source. Often the telescope should point at a part of the dome that at has been specially prepared for flat fielding, for example by the use of a flat white screen. A more modern process is to use electro-luminescent screen designed for the instrument you are using. These are designed to give a very constant level illumination. Although it may not be as flat a source as the twilight sky, it is very stable and gives for very flat master flats.

In order to take your calibration frames you will need to cool the camera down to working temperature and take a series of Bias frames at the same binning as you expect to (or have) used during your observing run. You need multiple biases to make a master bias; eight should be enough. Depending on what software you use later it is advisable to name the files such that they can be correct identified and are postfixed with a number, for example, Bias$-4 \times 4 - 1$.fit, Bias$-4 \times 4 - 2$, Bias$-4 \times 4 - 3$ would be an acceptable naming sequence where the 4×4 is the binning, and the number is the frame number. Most camera control software should do this for you, but you might need to set it up first you may also need to set the frame type card in the header to Bias, if the camera hasn't done this already.

You will also need to do the same with your dark frames, although these will also be exposure specific. You should already have some idea from your planning what exposure times you are to use, but you should have a reasonable range of times, say 1, 5, 10, 15, 30, 60, 120, 240 and 300s. As with the bias frames, you should take a series of dark frames for each exposure time you will using. If your telescope is small enough to have a lens cap, I suggest putting this on when taking darks, as well as turning the lights off as it reduced the chance of stray light. If you can set the

Fig. 8.2 Flow charts showing the processes required to make master bias, dark and flat frames and the (process to convert a light frame to a science frame)

camera software to take a series of darks automatically, as you can in MaxImDL in the Autosave tab of the camera control window, it might be a good time to take a break while the camera takes the frames.

In theory you should also take multiple flat fields for all the filters you intend to use, However, good flats are challenging and time-consuming so you may wish you use the flats supplied by the Observatory staff, who might also be able to supply bias and dark frames if time is short. If however, you wish to take your own you will need to take multiple flats for all the filters you are using during your run. If you use supplied flats ensure you know if they are master or single frames.

If you are taking flats, either dome or twilight, the process is broadly the same. In the case of sky flats, you will be imaging the part of the sky directly opposite the Sun (which will be set) as this is considered the most flat. If it is cloudy or foggy in the area of the sky where you intend to take the flats, then you will need to do dome flats. If you intend to do dome flats, follow the procedures outlined by the observatory staff, which might require pointing at a particular part of the dome, or using special lighting. You should take dome flats for every filter and binning you intend to use and, if possible, as many of each as you can in the time you have. A good flat fills each pixel up to about half its full well capacity, about 30,000 counts for a 16-bit camera. Your observatory staff should have some idea of what exposure time you need for each binning and filter to achieve this (note it is also camera and telescope dependent). If you do not have this information, it will be trial and error to find the correct exposure time. Do not keep the frames where the exposure time is too short or too long as this leads to errors. You only need to be around the half its full well capacity. It does not matter if you are a few thousand counts out, as long as the exposure time for each set is consistent when you are taken them. If you are taking sky flats and you can see stars in the frame, the sky has become too dark for flat fielding. If your flat exposure item is short perhaps under 5 s, it means your source is too bright, and you might have errors caused by the shutter in the flat field. A neutral density filter or neutral density film placed in the optical pathway can help to increase the exposure time without affecting the flat.

Once you have calibration frames, you will need to create **Master Frames**. These are calibration frames produced by taking the mean pixel value of a number of calibration frames. This is done in order to reduce noise that would be present in single frames. Depending on the level of error, you are willing to tolerate you can use single frames as master frames; this is especially true if you do not intend to do photometry.

In order to obtain your master bias calibration frames, you will need to find the mean of the frames; this is a pixel to pixel mean and not the mean across the frame. Be very careful when making masters. All the frames used to make the master must be of the same size, i.e. have the same binning (for dark, bias and flats), have the same exposure time (for dark frames) and the same filter (for light frames). Normally if you do not have the same binning level, the process will fail, but this is not the case with filters and exposure time. If you make a mistake at this point and it goes unnoticed, it will effect all you images. Most image processing software has some form of image mathematics. If yours does not, then it is reasonably straight forward to create master frames using Python although you will need the PyFITS and Numpy modules installed.

To turn your bias frames into master bias frames, pixel to pixel co-add all the bias frames of the binning and then divide every pixel by the number of co-added frames.

For the master dark frames, pixel to pixel to pixel co-add all the frames with the same binning AND exposure time. Divide every pixel in the resultant frame by the number of co-add frames. From the result of this pixel to pixel mathematics, remove the bias master that has the corresponding binning. The result is a master dark for that binning and exposure time.

To creating master flats pixel to pixel co-add all the frames with the same binning AND filter. Divide every pixel in the resultant frame by the number of co-add frames. From the result of this pixel to pixel remove the bias master that has the corresponding binning and the master dark with the corresponding exposure and time binning. Find the mean pixel value within the frame and divide every pixel by that value in order to normalise it. The result is a master flat for that filter and binning.

Applying the masters to light frames is straight forward, and your imaging processing software would be able to do this for you. From your light frame, pixel to pixel subtracts the master bias with the same binning, then from this subtract the master dark with the correct exposure time and binning. Finally pixel to pixel divided the resultant frame by the master flat with the correct binning and filter. The result is the science frame.

8.3.2 Pointing

Once you have all your calibration frames done, open the dome (assuming it's dark). Some telescopes require you align to two or three bright stars. If you need to do this, do it now. Your observatory staff should be able to help you perform a pointing alignment if you have not done one before. Even if you don't need to use alignment stars, it is always a good idea to check the alignment by slewing to a bright star and taking a short exposure. Centre the star in the image and, if the telescope has this ability, sync the telescope to the star. Syncing recalibrated the encoders to match the position of the target star. Sync to the wrong target can cause problems so make sure that the star your are syncing to is the one you think it is. If you can not see the star, then slew around in small steps taking images until you can either see the star or it is clear that your telescope is out of alignment. Note that you might be able to see some stray light coming from the star at the edges of the frame before you see the actual star that can provide a hint where the star is in relationship to the frame.

If you unable to find the target star with a reasonable distance the use the finder scope. Most telescopes have a small telescope with a wide field of view, co-mounted on the main telescope. These should have cross-hairs or if you are lucky, an illuminated dot, in the centre. Try placing you star in the centre of the finder scope. If you can't see the star in the finder, you have a more serious problem but most of these are easy to resolve.

If you cannot find the target star, there are some possible reasons. You may have an incorrect RA and Dec, the time or date or location could be incorrect on the telescope controller, the telescope could have been incorrectly parked or the encoders that tell the controller the position of the telescope in Alt Az, could be offset. It should be easy to check that the time and position is correct, make sure that you are using the time without daylight saving, it is often the cause of errors.

If you can, synchronise the telescope to an RA and Dec, in addition to syncing with a known star, a process known as **plate solving** may help. Plate solving involved automatically identifying the stars in the image and using their relative positions matched against stellar catalogues, to determine the centre of the image, which can be used to synchronise the telescope. A number of telescope control packages, including CCDsoft and MaximDL have this feature. It is also available as online services from astrometry.net. Note that if the telescope is a long way out of focus most plate solving programs will struggle as they will be unable to identify stars in the frame. In which case you may wish to jump to the focusing section of this practical and then return to aligning the telescope.

If the time and location is correct and plate solving fails to correct the pointing problem, try to **home** the telescope if it has this feature. Homing moves the telescope to the absolute zero position of the encoders if you have a minor problem with the encoders this may resolve it. If this fails, try parking the telescope. This will put the telescope into a known position if there is a pointing problem you might find that the telescope does not return to the park position. If this is the case, unlock the telescope and move it to its park position by hand, although you should check with the observatory staff beforehand if this is the correct procedure.

If none of these processes clear the pointing problem, then you may wish to escalate the problem to the observatory staff.

8.3.3 Focusing

There is a tenancy for users new to telescopes, to struggle with focusing. However, once you know, the correct technique focusing is straight forward. Focusing is achieved by moving the focal plain of the telescope so that it intersects with the CCD or retina. Focus may be accomplished by either moving the principle optical component, i.e., the mirror or moving the detector by using a focuser, a mechanical or electro-mechanical devices that moves the position of the sensor. If you have the facility of moving the mirror and using a focuser, ONLY use the focuser. The only reason you should need the mirror focus is it you are changing the detector and in most cases this should only be done by observatory staff.

It is likely that you will have to check and adjust the focus several times during the night. This will partly due to the contraction or expansion of the telescope due to temperature changes but also to changes in air mass and the slight movement of the camera as the telescope moved between horizon and zenith.

The air mass is the path length that the light takes through the atmosphere with an object at the zenith having an air mass of 1.0. Air mass can be calculated from the angular distance that an object lies from the zenith, the zenith distance, z. The air mass X can be found from the sec of z.

Before I take a science image, I always like to take a short focusing imaging if I have the time. This is just long enough so that I check that the field is in focus, and the telescope is pointing at the correct location. Most of the time no refocus is needed, but it is better to waste a few 10-s exposures than several 300-s ones.

Independent of the focuser design the procedure for focusing is the same. Initially, it would seem that a star, being a point source at infinity, should occupy only one pixel, which is not the case. The light is spread out over several pixel by a combination of diffraction within the optics and atmospheric turbulence. The amount that the light is spread is known as the **Seeing**; if you plot pixel position against count you will find that plot, for a focused star, appears Gaussian. This plot is the point spread function **PSF** of the star. A star in focus appears as round, solid and small object on the CCD, whilst an out of focus star will have an extended often doughnut-like structure (caused by the obstruction of the secondary) appearance.

Many amateurs use a diffraction mask to assist in focusing. This is a grid structure often made of cardboard or plastic, that is put in the aperture of the telescope. The mask makes a diffraction patten which creates a distinctive pattern when focused. The most common mask in use is the **Bahtinov mask**, named after its inventor. The mast produces three diffraction spikes around a bright source, normally a star. As the telescope is focused the three spikes begin to intercept at the same point, which should be the star being focused on. For large telescopes, over about 0.4 m Bahtinov mask can become troublesome to use, and many telescopes over this size have astronomical cameras rather the Digitial SLRs and auto focusers. Consequently, they tend not to be used on larger telescopes (Fig. 8.3).

When you start up you will need to perform a focus. This is done by point the telescope at a bright star and taking an image (look through the telescope if using just an eyepiece). Then move the focuser in one direction, either out or in and take a

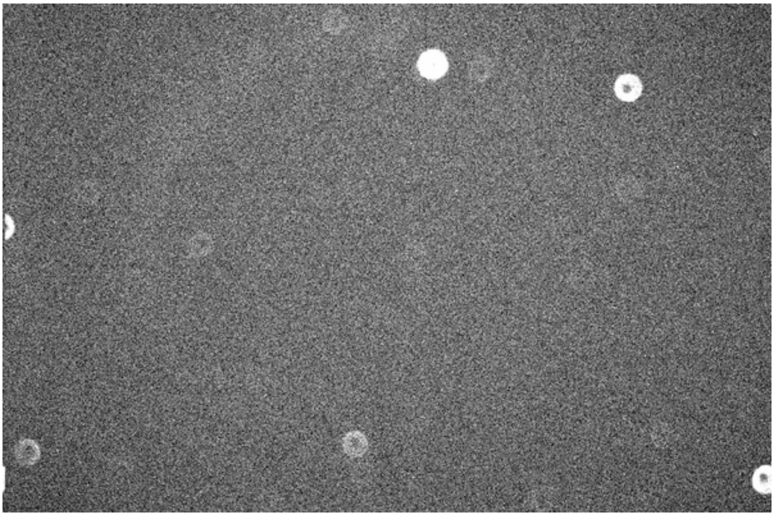

Fig. 8.3 An out of focus science frame. we can clearly see that instead of point sources we have the characteristic doughnut caused by the obstruction of the pathway by the secondary. The object near the *top centre* is, in fact, bright enough to fill in, but it too is out of focus (Image courtesy University of Hertfordshire)

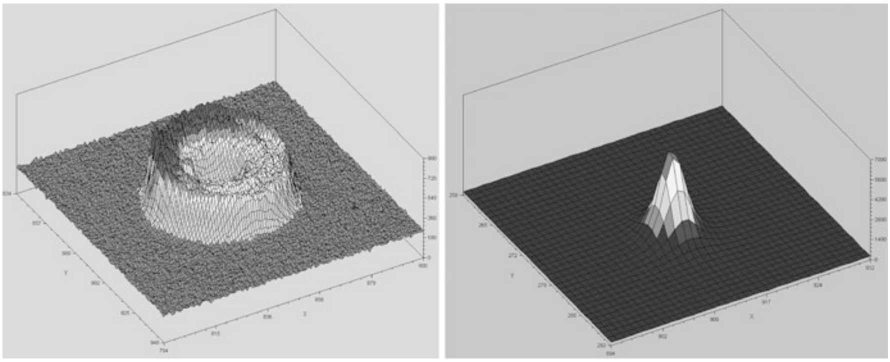

Fig. 8.4 The point spread function of an out of focus star (*left*) and in focus star (*right*). The doughnut PSF structure can be clearly seen, although in the image this star had “filled in” the doughnut. This contrasts to the eps on the *right*, which shows a high, near Gaussian profile, with a distinctive peak indicating that a good focus has been achieved (Image courtesy University of Hertfordshire)

second image. If the image of the star is getting smaller, you are moving to focuser in the right direction. If it is not getting smaller, move the focuser in the opposite direction. Repeat this cycle until the star is as small as you can get it.

There is a tendency for the bright stars to fill in when focusing, so when you have finished focusing on the bright star move on to a fainter star and repeat the cycle until all the stars in the field are focused. Note that you will need to refocus throughout the night as the altitude of the target, the air temperature and possibly the filters used, can all change the focus point.

Your camera control may have some focusing tools, which can help your initial focus. For example, it may have the ability to plot the PSF of the star you are focusing. If you have this feature, it will be easier to determine when the star is in focus; which will occur when the peak is at it’s minimum diameter. Also, you might have the ability to download only a portion of the CCD, a feature known as sub-framing. This enables you to take faster images of the object you are focusing on, but do not forget to turn it off when you have finished. Lastly if you have a powered focuser, you may have the ability to autofocus. The auto focuser moves the focuser in and out until the Full Width Half Maximum (FWHM) of the stars eps is minimised. The FWHM is a parameter often encountered is physics and is the width of a peak measured at half the peaks height (Fig. 8.4).

8.4 Guiding

In most cases, the tracking on the mount you are using will not be good enough to undertake very long exposures without some kind of trailing. Many CCD cameras come with a secondary CCD built-in or an accessory which consists of a small, low cost, CCD and a small mirror or prism to guide a focus a proportion of the image onto the CCD, these are know as off-axis guiders. Alternatively your telescope may use a separate co-axial telescope and CCD for guiding. With your camera control software, you should be able to access the guider CCD. Either this will automatically pick a star on which to guide or you will have to pick a star. If your field is very crowded, contains a large galaxy or a nebula, the auto guider may have problems identifying a suitable candidate star, in which case you may have to select one for it. The functionality of your guider will depend on the software you are using so, once again, check with the observatory staff before using. In some cases, you may not always need the auto guider. For very short exposures the amount a star moves may very well be less than one pixel in this case auto guiding may be unnecessary, whilst for longer exposures the need for auto guiding will depend on the tracking quality of the mount, although as telescope mounts are designed for sidereal tracking, when imaging solar system objects, which do not move sidereally, you will only able to take short exposures before you see the objects trail even with auto guiding.

If an auto guider is not available if you do not wish to use it for some reason, or it is struggling to identify a suitable guide object, an alternative to tracking is stacking. We will discussing stacking in more detail later. However broadly it is a technique whereby a single exposure is generated by the summation of several shorter exposures.

For each of your targets, you will have to either setup the auto guider or if it automatically picks a guide star, check that it is indeed guiding. Generally changing a filter should not be a problem for the guider. Most telescope control software reports guider data as it exposures, and you may wish to abort the exposure if guiding is lost for than a few tens of seconds.

If you can adjust the exposure time of the guide CCD, it is often useful to increase the exposure time in order to either make the guide stars more prominent or to increase the rate of correction. Beware setting the exposure time too short as it can cause problems with the control computer.

8.5 Target Specific Problems

In this section, we will discuss some of the challenges specific to certain target types. Remember that there is a fundamental difference between science images and the images you see on the cover of books and glossy science magazines. The processes used to make these images pleasing to the eye often means that the pixel values are no longer linear. Aside from applying calibration frames you should not perform any action that is likely to change the linearity of the frame, for example by applying a Gaussian filter, unless you intend not to make any measurements on the image. You may wish to scale or **stretch** the image which will make the fainter structure more visible. Stretching an image changes the way it is displayed. A modern CCD has a range of pixel values far in excess of what can be detected by the human eye when viewed as a grey scale, a problem that also applies to most monitors. By stretching

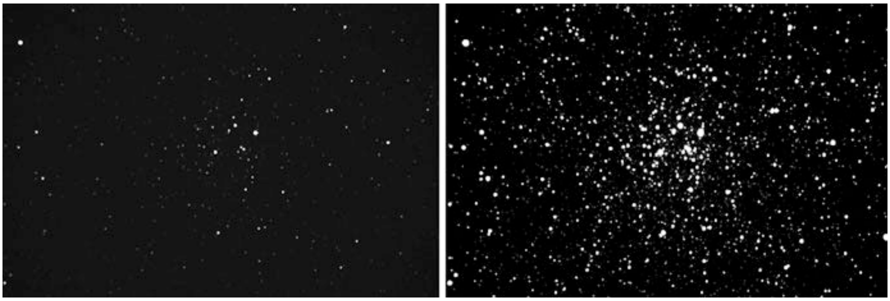

Fig. 8.5 A clear image of the Open cluster M26. The *left-hand image* is how it would normally be displayed whilst the *right-hand image* is stretched. The pixel range being displayed is reduced so that faint stars are now displayed without making the brighter stars too bright. Stretching in this way only effects how the image is displayed and does not change the pixel values (Image courtesy University of Hertfordshire)

your display only the pixel range that is of interest, which makes the detail more visible. Figure 8.5 shows an unstretched and stretched version the same image of M26. it is clear that we can now see many more stars now the image has been stretched. This only changes how the image is displayed and does not alter the pixel values, so the frame remains scientifically useful.

Most applications perform stretching using a frequency plot which shows the distribution of the pixel values. There will be a large number of pixels with a low count, representing the background with few pixels at higher levels. With most applications, the range of pixel values displaced is either set by a pair of drag bars on the plot or by a pair of sliders. By sliding the bars, you can enhance the display of fainter details. Typically this is done by moving the minimum value near the centre of the background peak and the maximum bar to a lower value, with the best location often being slightly right of the background peak. Many camera control applications remember the stretch settings, but the best settings depending on a number of factors. If you take an image in one band and perform a stretch and then change filter, do not forget to change the stretch setting otherwise you may not see the target in the image.

8.5.1 Imaging Open Clusters

Open Clusters, also known as Galactic Clusters are the site of the most recent star formation in the Galaxy and as the spiral arms are delineated by star formation they are always found within the confines of the Milky Way. Hence, we do not find open clusters in, for example, Ursa Major but we find many in Cassiopeia. Open clusters form from collapsing molecular clouds which have become dense enough and cold enough that gravity overcomes the thermal motion, turbulence and magnetic fields that can prevent collapse. However, do not be fooled by science fiction shows, the density of these clouds is still lower than any vacuum ever achieved on Earth!

As the cloud collapses, local, smaller regions become over dense and these further collapses into stars. Overtime radiation pressure, stellar winds and supernova from the more massive stars formed within the cluster, strip the cluster of the gas and dust that was not turned into the stars. It is this dust and gas that makes up the bulk of the mass of the cluster, not as you would expect, the stars. In fact, star formation is quite inefficient with only about 1–3 % of the mass of the initial cloud becoming stars. With the stripping of this gas, the cluster becomes gravitational unbound and begins to break up with the stars becoming members of the galactic disk. Depending on the mass of the cluster this can take anywhere between tens of millions to a billion years.

As a point of interest, open clusters appear to mass segregated with more massive stars being found near the centre of the cluster and lower mass star near the edge. Why this happens is a matter of debate. Models suggest that over time gravitational interaction will segregated the cluster. However, large-scale star formation models suggest that the most massive stars tend to form in the densest part of the cloud which is typically at the centre. Perhaps it is more likely that both processes contribute to the overall distribution of stars within the cluster. None the less, Open Clusters are some of the less challenging objects for the proto-astrophysicists to image.

The principle problem undergraduates encounter when exposing an Open Cluster is over exposing and therefore saturating the brighter stars, especially when an exposure time is not adjusted between bands. Over exposure or in fact entering the non-linear region of the CCD which occurs before saturation, can render the image effectively useless for science and is not immediately obvious when an image is displayed. It is important to understand a number of effects. Firstly the size of the star on the image is not representative of its brightness as the processes that causes the light to be spread over multiple pixels, diffraction and atmospheric seeing, effects all stars in the field the same and to the same extent and is, therefore, independent of the brightness of the source. Visual detection of saturation is problematic as computer monitors are unable to display the full pixel range in grey levels and even if it could the human eye would not be able to distinguish between the grey scales over the entire range. To resolve this you will have to perform and image stretch (also known as scaling); this has nothing to do with changing the shape or size of the image, rather it is how the image is displayed. In general stretching does little to improve the detection of saturated pixels but it does have an import application which will be discussed later.

How is saturated detect? The simplest method is to do a surface plot or a slice of some of the brighter stars in the field. Most astronomical imaging or image processing applications should be able to perform this function, including MaxIm DL, APT and SalsaJ. To illustrate Fig. 8.6 shows a B-band image of the open cluster M26. The image is inverted i.e., black and white are reversed, in order to make it clearer. From visual inspection, none of the stars appear saturated. However when

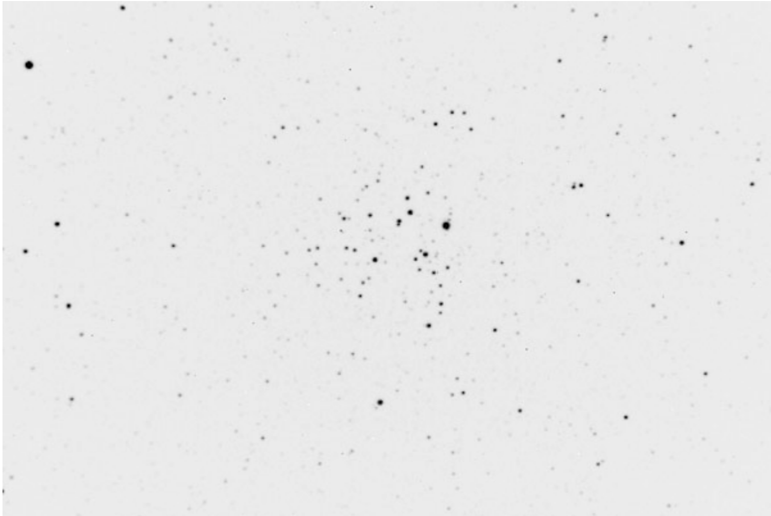

Fig. 8.6 Inverted B-band image of the open cluster M26 (Image courtesy University of Hertfordshire)

we perform a surface plot of a bright star within the field, Fig. 8.7 we see it does not have the characteristic Gaussian shape we would expect, rather it is domed. This is an indicating that the exposure was long enough that the pixels to become non-linear. Note that if the top of the peak was flat, this would indicate saturation, although, of course, the pixels would have gone non-linear before this happen. The smaller peak in Fig. 8.7 is another, less bright star, caught within the plot region. As can be seen, this is broadly Gaussian and, therefore, acceptable. Figure 8.8 shows the same region as Fig. 8.7 but, in this case, is a slice instead of a region plot. However, we can see the non-linear nature of the bright star within the plot and the Gaussian nature of the less bright star also intersected by the slice.

Ideally you want the brightest star in the image to be just inside the linear range of the CCD. If you have a saturated star in the field, you may wish to retake the light frame with a reduced exposure time. If you know the count and the magnitude of an unsaturated star (see the chapter on photometry as to how to measure the count) as well as the magnitude of the saturated star, you can find out the count of the saturated star if it was unsaturated using Eq. 8.1 where M_{ref} and M_i are the magnitude of the unsaturated and saturated star respectively and I_{ref} and I_i are their integrated count; you can scale your exposure time such that in the revised exposure the star is unsaturated. Note that this is band specific and to some extent, spectral type specific and uncertainties can be high.

$$I_{ref} \times 10^{\frac{M_{ref}-M_i}{2.5}} = I_i \tag{8.1}$$

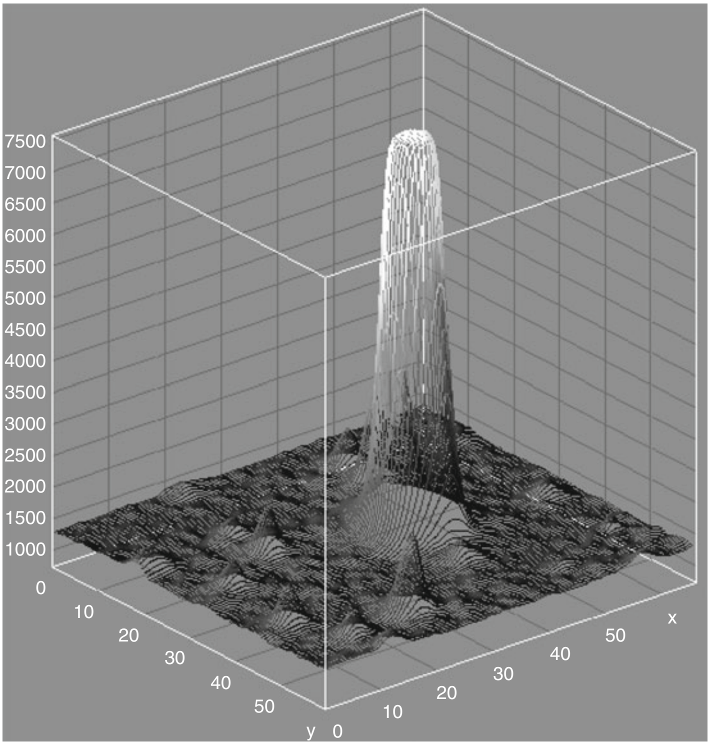

Fig. 8.7 Surface plot of a bright star within Fig. 8.6 using SalsaJ. The dome-like structure of the plot suggests that the star in within the non-linear region of the CCD. A flat top, as opposed to a dome, would indicate saturation. Note the smaller peak which is much more Gaussian-like than, the larger peak. This is a less bright star within the region of the plot, and its Gaussian nature indicates that its pixel values lie completely with the linear region of the CCD

8.5.2 Imaging Globular Clusters

Globular clusters are very different objects to open clusters. They are old, very dense, found exclusively in the Galactic halo and, in general, are considerably more massive than open clusters. Unlike open clusters no star formation has even been detected within a Globular cluster and the stars contained within have lower levels of elements with atomic numbers higher than two (which in astronomy parlance are known as metals with the ratio between hydrogen and metals being the **Metallicity**) indicating that they have formed early in the evolution of the galaxy. These low metallicity stars are known as **Population II** stars, while the younger, higher metallicity stars such as those found in open clusters and also includes stars such as the Sun, are known are **Population I**. There is a possibility that ancient, ultra low metallicity, **Population III** stars exists, most likely as very low mass and, therefore, slow burning, M-Dwarfs but these have yet to be detected.

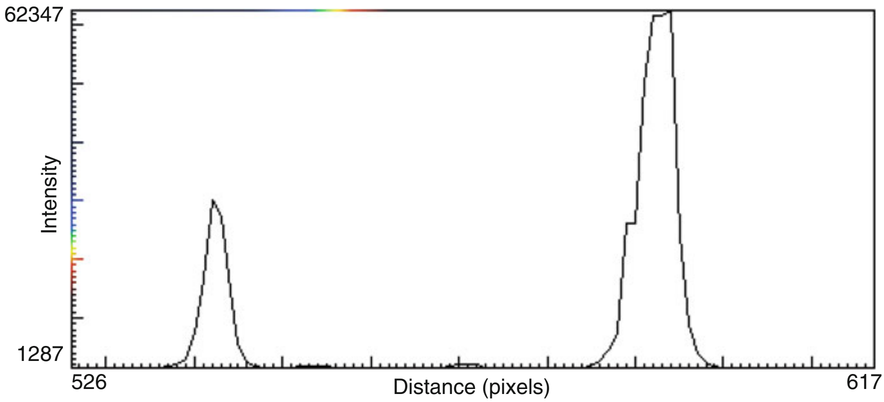

Fig. 8.8 A slice though a bright star within Fig. 8.6 using SalsaJ. The large peak is clearly non-Gaussian indicating that it lies within the non-linear region of the CCD (a flattened peak would also indicate saturation). The small peak is from another star also in the slice. This appears broadly Gaussian in nature as we would expect and hence lies in the linear range of the CCD

The challenges for imaging a globular cluster are very different from those of an open cluster. As globular clusters are in the halo, they lie at much greater distances than many open clusters and, therefore, any very bright star is likely to be foreground stars rather than a cluster member. However, the globular clusters is a very dense ball of stars, perhaps up to several hundreds of thousands of stars with an average density of perhaps one per cubic parsec. Although this does not sound significant, near the centre of the cluster the average distance between stars may decrease to a few hundred astronomical units.

Consider the line of sight as we pass from the edge of the cluster towards the centre. Near the edge the number of stars within the line of sight is low and individual stars are seen. However, as we move towards the centre, the number stars with the line of sight increases to the point that we are no longer able to resolve individual stars out.

The secret to imaging a globular cluster, therefore, is to enhance the detectable stars near the edge whilst not saturating the core, where we are unable to resolve the stars out. This can achieve by a taking multiple images and stacking them in a non-linear fashion. This can be used for general imaging and astrometry but not for photometry, as it makes the pixel values non-linear.

Figure 8.9 shows a science frame of the Globular Cluster M3 taken in clear. The exposure time was 60 s and no other action other than the removal of the bias, dark and flat frames has been undertaken, except for inverting the image. We can clearly see stars near the edge of the cluster but the centre *appears* saturated.

Figure 8.10 shows a slice through Fig. 8.9. The apparent rise in the background near to the centre of the field is in fact due to the stars in the centre of the cluster merging in the image. We can also see that there are a number of regions in the centre that are saturated. The sharp peak located to the left is due to a cosmic ray.

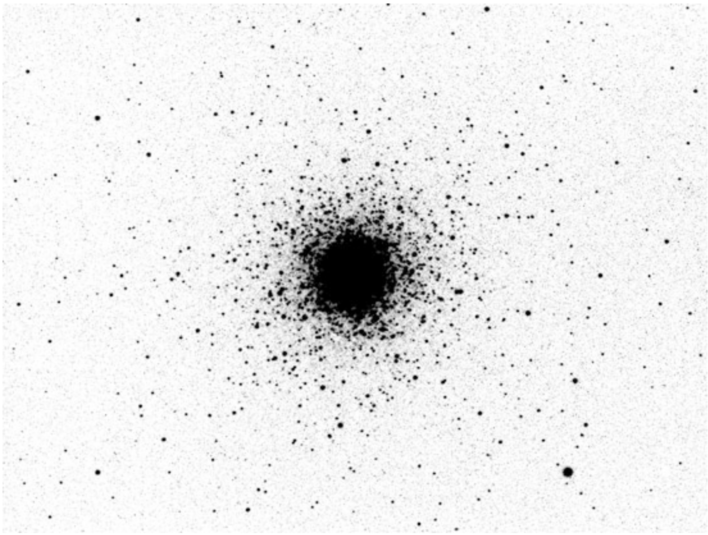

Fig. 8.9 An inverted science frame of the Globular Cluster M3 taken in clear (i.e. without a filter). This image has yet to undergo any post-production correction to improve the visibility of the stars near the *centre* (Image courtesy University of Hertfordshire)

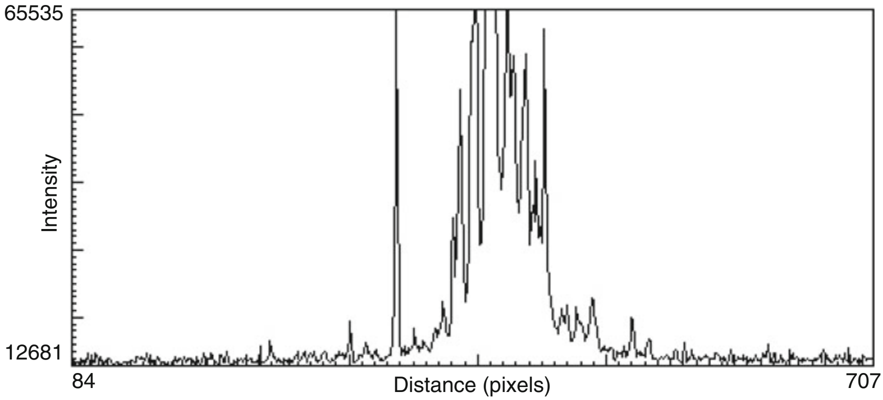

Fig. 8.10 A slice though Fig. 8.9 using SalsaJ. We can see that near the centre of the globular cluster there are many saturated stars The sharp peak to the left of the main body is caused by a cosmic ray

8.5.3 Imaging Galaxies

In many ways, the challenges related to imaging globular clusters applies to galaxies. Spiral galaxies have bright cores comprising a dense congregation of old stars surrounded by a thin disk of stars and gas with the spiral structure being delineated by the regions where stars are forming. Elliptical Galaxies are, no longer

Fig. 8.11 An image of the spiral galaxy M81. This was taken in three filters with *blue* being a Johnson B filter, *green* a Johnson V filter and *red* an $H\alpha$ filter. The use of an $H\alpha$ filter enables the spiral arms, the regions of star formation, to be distinguished (Image courtesy University of Hertfordshire)

star forming (or are doing so at a very low rate) and have become virialised. In essence this means that the kinetic energy of the stars within the galaxy has become evenly distributed, there is no disk, as with a spiral galaxy, just a dense ellipse of stars. Imaging elliptical galaxies is just a matter of trying to keep the galaxy in the linear region of the CCD. Remember that the most quoted magnitudes for galaxies are the integrated magnitude and not the surface brightness, it is the formed you need to use to determine the exposure time. Spiral galaxies can have considerable structure which is often very faint. This leads to a dichotomy as the nucleus is very bright so you will want a short exposure whilst the arms and dust lanes are faint and require a long exposure. In general it is better to stay within the linear range of the CCD but take multiple images and mathematically manipulate them to bring out the structure. As spiral galaxies are star-forming they produce can a considerable amount of $H\alpha$ emission which is limited to the location of the star formation, i.e. in the arms. Hence, the arms tend to be brighter in $H\alpha$ than the nucleus. The addition, therefore, of an $H\alpha$ filter when imaging can help delineate the arms, Fig. 8.11 svhows the effect of imaging a star-forming galaxy in $H\alpha$.

8.5.4 Planetary and Emission and Dark Nebulae

Planetary nebulae (PNe), called so due to their planetary like appearance in early telescopes, are the shells of stars that have been discarded during the red giant phase. The hot progenitor star, often a white dwarf, ionizes the expanding envelope which leads to the emission of forbidden lines. Planetary nebulae have complex structures that are very poorly understood. Many exhibit an hour-glass like structure and often appear circular or elliptical in small telescopes. Planetary nebulae are rare in the Milky Way despite the large number of known possible progenitor stars; suggesting that they are very short lives, perhaps only 10,000 years.

Planetary nebulae can be challenging if their angular size is small, as many are, as this will make resolving detail difficult. Although it is possible to image PNe in broadband filters, their very nature lends them to narrow band imaging as this is where the bulk of the emission lays. The use of narrow brand filters will also reduce the sky contribution improving signal to noise. However, for very faint PNe very long exposures are needed which can be achieved by taking multiple images and stacking them later. If they are very small, the introduction of a barlow or similar device into the optical pathway can increase the plate scale.

Emission Nebula, of which PNe are a sub-class, are clouds of gas often star forming and with embedded stars. These stars ionise the gas which then de-ionises emitting light. Just as in the case of PNe, emission nebula can be very diffuse and faint but can also be very bright in some regions especially if there are massive stars illuminating the cloud. One such example is M42, also known as the Orion nebula, as illustrated by Fig. 8.12 M42 is also a **diffuse nebula** having no well-defined boundary as can be seen. In astrophysics, di-atomic hydrogen is known as

Fig. 8.12 An image of the emission nebula M42. This was taken in three narrow band filters with *blue* being an [OIII] filter, *green* a $H\alpha$ filter and *red* an [SII] filter. The use of narrow filters enables subtle features within the nebula, to be distinguished. Note the very bright core which is host to a number of massive stars (Image courtesy University of Hertfordshire)

Hydrogen or just H, disassociated hydrogen is known as **HI** whilst ionised hydrogen is known as **HII**. Hence M42, which is mostly ionised, is an HII region, as well as a diffuse and an emission nebula.

The core of M42 is illuminated by a number of very massive stars which can cause the saturation of the core. Only by stretching the image, reducing the dominance of the core emission do we see the fine structure. There are a number of computational filters that can enhance detail, for example, a logarithmic stretch, although this will make any results non-photometric. However, in many cases it is the morphology we are interested in which case sacrificing the linearity of the image is acceptable.

Dark Nebulae are dense cold molecular clouds. Because of their nature they are only visible in the optical when located in front of an emission nebula. Although they often appear dark in front of the emission nebula. Dark Nebula can be seen in Fig. 8.12, although perhaps the most famous example is IC434. Hence when imaging dark nebulae the integration time will depend on the brightness of the background source.

8.6 Stacking and Dithering

For unresolved astronomical objects, it is likely you will be producing one image in each band unless you are undertaking time series photometry that is discussed in Chap. 10. For resolved objects such as nebulae and galaxies, it is likely that your will have to have a long integration time to bring out the detail within these objects and to improve signal to noise. Very long integration times pose a number of problems. Firstly if there is a bright region in the image, for example, a foreground star, it may saturate and bloom before you have enough photons from your target. Secondly, the number of cosmic ray strikes on the CCD is a function of time, as you increase the exposure time the number of cosmic ray strikes goes up proportionally. Also, the longer you expose, the more likely that something will go wrong, for example, a plane or a satellite passing through the shot or tracking errors. Lastly, there just might not be enough good weather or even dark skies in one session, to complete the full exposure.

To get over these problems astronomers can take multi frames of the same target and pixel to pixel stack them. This technique is discussed more fully in Chap. 9. A stacked frame is not limited to the same bit constraints as a camera; so for example, images from a 16-bit camera can be stacked to form a 32-Bit FITS, solving the problem of very bright objects in the field.

Images have to be aligned before they can be stacked unless your telescope has perfect tracking. However, the alignment process enables us to deal with a number of issues. Variations within the CCD (for example bad pixels) can be evened out over the chip by **Nodding** or **Dithering** the telescope.Dithering is performed by moving the telescope a few arc seconds between exposures. Faults move with the telescope, and so they appear in a different parts location in each frame, this can be used to minimise their effects.

So we see that when we make our observations we must be aware of the post-imaging processing that will be needed and take a number of exposures at an exposure time appropriate to what is needed, both for the initial image and for later processing.

8.7 Practical 6: The Messier Marathon

The practicals in this chapter are designed to give you the skills to take science grade astronomical images in an efficient manner so that you can utilise your precious clear sky time to the maximum. You will be imaging a wide range of targets that require the implementation of different imaging approaches, not only will you gain experience on the telescope but also a range of astronomical software packages

A common challenge amongst amateur astronomers is the Messier Marathon, trying to see how many Messier objects one can image in a single night. A successful Messier Marathon careful requires planning and perpetration, exactly the skills a good observational astronomer needs. If you are a Southern Hemisphere observer, the Messier Marathon can be quite challenging due to the limited number of southern objects and the low altitude of many of those that are visible. If this is the case, we suggest you use the Caldwell Catalogue instead or as an addition. As noted in Chap. 6, the Caldwell catalogue contains a further 109 bright objects that are not in the Messier Catalogue.

8.7.1 Aims

In this practical you will attempt to image as many as Messier objects as possible in a single observing session. Not only that, you will make colour composite images meaning you will need to image in at least three filters. This practical take careful planning and teamwork to do successfully. You will be familiarising yourself with using the telescope and the camera to image a wide range of object types from bright clusters to faint galaxies.

8.7.2 Planning

To successfully complete this practical, you will need some careful planning. Starting up you planetarium software and set it to the time and date you expected to observe. Bring up the constellation boundaries and names and have a look at what will be visible and determine if it is rising or setting. Check the position of the Moon, if there is any as for some of your targets, you may need to avoid the Moon.

You will now need to look at the Messier catalogue. You need to observe at least one open cluster; one barred spiral galaxy and elliptical galaxy, a spiral galaxy, an HII region or supernova remnant, a planetary nebula and a globular cluster. Using your understanding of what is currently visible pick the best-placed target of each kind. You will also need backup targets for each class of object.

Once you have selected your targets you need to determine what order you will observe them and, based on their magnitudes, how long your exposure is going to be. Your should also be aware of the angular size of the objects, some maybe too small for you camera to image some, for example, M31, will be too large. You should image them in at least three bands. If you have narrow band filters (such as OIII Hα and SII), the HII region and the planetary nebula will benefit from being imaged with these. Although the use of narrow band filters will increase integration time, low surface brightness objects that emit strongly in these lines benefit greatly as the sky tends not to be so bright around forbidden lines (although there are still emission lines there) so your signal to noise ratio will improve for these sources.

8.7.3 Getting Ready

Before you leave for the observatory make sure you have all your documentation, your lab book, something to put you images on, such as memory stick and appropriate clothing. It is best to assume that it will be colder than forecast as you can always take a sweater off, but you can't put one on you don't have!

1. Once you arrive at the observatory you need to start up the telescope and the instrument. A member of Observatory staff (or your tutor) should be able to help you if you don't know how to do this. The first thing you will need to do is start the camera coolers, if they are not already on, as it takes a while to bring the camera down to operational temperature.
2. Once the camera is down to temperature you will need to take calibration frames. If you have been recently observing with this telescope and camera and have bias, dark and flats for the binning, exposure times and filters you intend to use, then you can most likely use them instead of taking new ones, although they may limit your exposure times. If your flats are more than a week old, it is normally worth taking new ones. Dark and Bias frames are normally good for a month, as long as there have not been any major changes in the setup of the telescope or new equipment installed in the dome.
3. Once you have taken your calibration frames, if you need them, open the telescope dome and go through the pointing and focus routine. Once this is done, you can move on to imaging.

8.7.4 Imaging

1. For each of you targets you will need to slew to the target using the coordinates or if you have a GOTO facility, the targets name.
2. Take a single fairly short, image in clear or the nearest equivalent filter, for example, a Johnson V. Use this to check your focusing.
3. Confirm that the target is in the centre of the field. If it has low surface brightness, the object might not be visible without stretching and even with stretching it might be challenging to distinguish. If you cannot identify the target, use the stars and you finder target to attempt to locate the object.
4. Once you are happy with the centring of your target and the focusing, turn on and calibrate your guider and check that you are tracking. Remember that guiding is a process requiring an interactive response between camera and mount while tracking is just the sidereal movement of the telescope.
5. Pick your first filter and start imaging your target. Once the first image is displayed check that it is not under or over exposed. You may need to stretch the image. If your exposure is too short, you can work out how much longer you need by examining the pixel value of the brightest target. Pixel value should effectively scale with exposure time, although of course the, bias and pedestal do not. However, the contribution to your pixel value of these components should be quite small, and their non-linear nature gives you a degree of headroom when scaling up exposure times. For example, if you are exposing for 10 s and your target's maximum pixel count is 10,000, assuming that your maximum linear pixel count is 50,000, you can safely exposure for 50 s because the bias and pedestal, which is likely to be around 100 counts, remains constant with time; this being the case the maximum pixel count is more likely to be 49,600 for a 50-s exposure.
6. Once you have done this process for all three filters, and you have saved your images and noted your exposure times and files names in the log book, you can move on to the next target. However, you may wish to take multiple images in each band and stack them in software later as this reduces noise as we will see in Chaps. 9 and 11. If this is the case, once you have your first set of RGB images (or whatever filter series you are using) and you have confirmed your exposure time, you may wish to use the automation service available in your camera control package. This will take a series of images in multiple filters without requiring your interaction. However, I strongly recommend you always take the first set manually to confirm that the chosen setup is correct. Likewise, you should be aware how long it will take to run and be at the telescope before it finishes otherwise the instrument might stand idle when it could be being used.
7. Although not entirely necessary, many amateur astronomers like to take a clear frame in addition to the three coloured filters. The clear is used to form a **luminance frame** in the a modified version of the standard RGB colour images, known as an LRGB. LRGB tend to have better definition of fine structure, but not all imaging software can produce LRGB images although MaxImDL can.

8.7.5 Shutdown

Once you have all your light frames, do not forget to take your calibration frames, if you haven't already. Make sure you have everything in your lab book you need including the details of the telescope and the time you finished imaging.

It is likely that your instrument has a specific shutdown procedure that hopefully you know if you do not, please check with a member of observatory staff. In general this procedure requires the parking of the telescope which put the instrument in a known position so that when it restarted it knows where it is pointing, The dome will need closing and, if it a tracking dome, may also need parking. You should shut the camera and its cooler down and turn any ancillary equipment that you might have turned off, on and visa verse. For example, some observatories use dehumidifiers to protect the optics.

8.7.6 Post Imaging

After taking you images, you will need to produce your science frames. In order to do this, your will need to subtract the correct dark frame from each of your flats and then normalise the flat.

1. Subtract the dark frames from your light frames
2. Divide the result by the flat to give a science frame.

   Note that given the dark frame includes a bias it is not strictly necessary to have bias frames. However if you do not have a dark frame with the correct exposure time for some reason and you need to make one by scaling another dark frame you will need to subtract the bias first, as it doesn't scale, and then add it back on after the scaling. As bias frames are so quick and easy to take and can be useful, I suggest you always take them even if in the end you do not use them. The process of making science frames is discussed in more detail in Sect. 8.3.1 and in Chap. 9.
3. If you have multiple images of each band, you should stack these to improve signal noise. However beware, when you stack frames, the nature of the stack can destroy the scientific usefulness of the data. For the Messier Marathon this is not an issue as you are attempting just to produce an image but for photometry and spectrography you should only stack via addition although you may make a mean (but not a median) stack if all your exposures are of the same length. Details of how to align and stack images are in Chap. 9.
4. Once you have your separate three frames, four if you intend to do an LRGB, you will need to tidy them up and stretch them before combining them into your colour image. Hot pixels and cosmic rays will appear as unusually coloured pixels when you create the colour image and are easier to deal with at this

point using a cloning tool. If your astronomical imaging software has a hot pixel removal tool it is worth applying at this point, although I find that often they fail to get them all especially cosmic rays.

5. Combined the image using for favourite software into an RGB or LRGB image and then colour balance them as discussed in Chap. 9. Once you happy with you image export it in TIFF or JPEG. You may then use image standard image manipulation tools such as Photoshop or The Gimp to remove artefacts and to crop the image.

8.7.7 Analysis

When completed you should write up you work and include your images, finder charts, exposure times. During the write up pay particular attention to any problems you encounter and describe what you did to the images, why you did what you did and how it influenced them. In particular look at how the signal to noise improves when adding calibration frames and science frame stacking.

8.8 Planetary Imaging

It is often thought by students that the easiest object to image are planets, as they are so bright, but this is not true for most teaching observatories. In fact Planetary Imaging is extremely challenging for the very reason that planets are so bright. Most astronomical camera cannot cope with the number of photons being received from the brighter planets, Jupiter, Mars, Venus and Saturn and over expose even with their shortest exposure setting and with a photometric filter in the pathway. Another problem is that the telescope is not designed to track solar system objects but stars. Auto guiding is not of much help here as they would need to guide on the planet, which is not a point sources and cause problems with guiding software. Hence, even if we wished to take long exposures, we would suffer from trailing.

An obvious solution to this problem is to reduce the amount of light coming into the camera. This could be done either by **stopping down** the telescope, i.e. putting a cover on the front to reduce the aperture size or putting either a twin polarising filter, or Neutral Density Filter in the optical pathway. Another challenge is that to see the detail in the planet, high magnifications are needed. At high magnifications seeing becomes a major limiting factor. As we have seen, the turbulent motion of the atmosphere limits the resolution of the telescope which becomes very obvious with solar system objects. Again we could reduces our exposure time, to below the turbulent time scale, which can be smaller than a hundred of a second. Far shorter than most astronomical CCD can image.

8.8.1 Lucky Imaging

There is, however, a simple solution to these problems that was first implemented by amateur astronomers but is now being picked up by professional, the use of webcams. Although insensitive and noisy compared to astronomical cameras their speed and relative low-cost lead to their modification by the amateur into planetary imaging cameras. However, in recent years low noise astronomical webcams, both colour and black and white, have started appearing on the market and these are a cost effective solution to the problems with planetary imaging using conventional astronomical camera. Professional ultra-high speed cameras, which may be cooled and have noise and sensitivities levels close to standard astronomical camera as well as frame rates as high as a thousand frames a second, are also becoming available but at a considerable cost.

The advantage of using a high-speed, but relatively insensitive camera, is that their frame rate is typically faster than turbulent time scale. Imaging software such as, Registax are capable of processing the video file that comes off these cameras, frame by frame. By picking only the best aligned frames, where turbulence is not present and stacking those frames, ultra high resolution imagery is possible; well below the typical seeing and often limited by the diffraction limit of the telescope or the pixel size of the camera. Resolutions of 0.2 arcsec using this technique, known a **lucky imaging**, are common. This is typical of the resolutions achieved at a very high observatory such as in Hawaii or Chile, or with adaptive optics.

There is yet another use for **Lucky Imagers** as these devices have become known. Their very high frame rate allows the observation and precise timing, of very transient astronomical phenomena, which are too short for normal CCD based astronomical camera to capture. Examples of these are asteroid occultations, which may be a few seconds long, and lunar meteoroid impacts on the Moon.

Given the very light nature of most lucky imagers, the time taken in removing and then replacing an astronomical camera, the fact that a guide camera is not needed for planetary imaging and that most guide cameras are mounted on refractors that are ideal for planetary imaging, the is an advantage in using the guide telescope for planetary imaging. If you have a choice, and you might not, we would suggest going down this route for most planetary imaging.

8.9 Practical 7: The Speed of Light

In 1676, a Danish astronomer, Ole Römer noticed that the eclipses of Io, in front and behind Jupiter, lagged behind that predicted by orbital mechanics and that the lag seemed to be dependent on the distance of the Earth from the planet. Römer realised that this implied that the speed of light was finite and managed to measure it close to current estimates. With modern high speed camera and increasing computing power it is possible for you to follow in Ole Römer's footsteps.

8.9.1 Aims

In this practical you will image a solar system object using a high speed astronomical camera or webcam with the purpose of identifying the point in time in which an occultation occurs and looking at the difference between the predicted moment that it should occur (without accounting for the speed of light) and the actual moment occurred. You will then be able to calculate the speed of light.

8.9.2 Planning

This project needs careful planning as you will to identify an occultation or transit that will take place during your observing time. If Jupiter is visible, then a transit of Io is the best candidate as they are common and obvious. If not you could use an a planetary occultation of a star. Try to avoid using the moon as although lunar occultations of bright stars are common (planets less so), the moon is very close and will introduce large errors into your results.

You will need to use a piece of planetarium software to find the uncorrected time of the occultation or transit. Stellarium is excellent for this, just remember to turn off the option to simulate light speed found in the view window. You will also need to find the distance between the Earth and the Solar System object at the point of the occulation or transit. Again Stellarium can do this for you.

You are likely to generate a large video file, so make sure you have some way of getting this off the camera control computer. You will also need to carefully measure the moment the transit is observed so ensure you have an excellent well calibrated timepiece.

8.9.3 Observing

1. It is best to turn up at the observatory well before the transit is due so that you can make sure that everything is setup and that your are in focus and pointing correctly.
2. Start imaging your target about 15 min before the event is due, but there is no need to record at this point.Remember that you will be unable to auto guide and the telescope and even with the tracking on the target will move in the field. Hence, you will need to track the target manually.
3. You should see the two bodies involved in the occultation or transit in the frame and be able to make a reasonable judge of how close you are to the event occurring. Start recording when you think you are about 10 min from the transit. Try Make a careful note when you started recording, try to determine accuracy to

about a second and the frame rate the camera is running at. Depending on your object, the event could be delayed for as much as an hour from that reported by Stellarium.

4. Once you are happy that you have captured the event you wished to observe. Stop and save the video file, back it up and transfer it to an another media for removal from the camera control computer. Then before shutting the observatory down as normal.

8.9.4 Analysis

1. Using the video file, play the video until the moment that the transit or occultation occurs. You might need to rewind and play the video a number of times in order to identify the frame number.
2. Dividing the frame number by the frame rate gives the number of seconds after recording started that the event occurred. using your starting time and this figure, along with the uncorrected time and distance from Stellarium, you can calculate the speed of light.
3. As with previous practicals you need to write your report up in standard laboratory style. You should include notes about the camera and software used in addition to any problems you may have encountered. You should also include details of your preparation especially details concerning the occultation or eclipse and its expected, non-corrected, predicted time. Pay special attention how your result compares to other, more modern methods of determining c, what your uncertainties are, where they come from and what you would do differently if you were to do this same experiment.

Chapter 9
Image Reduction and Processing

Abstract Image processing is the process to turn an image from a science image to something more pleasing to the eye or highlighting structure of interest. Many image processing renders the image unsuitable for photometry. In this chapter, we discuss image processing techniques such as stacking, Sigma clipping, Drizzling, Stretching, as well as the application of image processing filters. We touch on how to make colour images from black and white astronomical images using commercial packages.

9.1 Introduction

In this chapter, we will discuss the processes by which you take your raw light image or images and create an improved image either for aesthetic or scientific reasons. Although these two goals are not always opposing, most of the processing required for the production of high quality glossy astronomical images renders them all but useless for scientific purposes. Hence we should distinguish between **Image Reduction**, which we can consider the process required to turn a light frame into a science frame, and **Image Processing**, which goes beyond Image Reduction to produce an image that is pleasant for the viewer.

It is, therefore, important to understand your final goal. If you wish to perform photometry or astrometry you will not be making coloured images nor will you apply highly sophisticated mathematical filters to your image. However, if you wish to look at the structure of the object then coloured images and filtering are useful tools. Likewise, if you are looking at producing spectacular images, then your stacking technique will be different for stacking photometric images. If you wish to process your image you may wish to use processing software such as Adobe's Photoshop or The Gimp to tidy up an image.

Observational astrophysics rather than astronomical imaging is the principle subject of this book; therefore, this section concentrates mostly on image reduction rather than image processing. However, most image reduction and image processing techniques are the same. We will, however, touch on image processing and the production of colour images. However, a full treaty on the matter is beyond the scope of this book.

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The basic work flow of image reduction and processing is: – calibration, alignment, stacking, combining and enhancement. The last two of which only apply to image processing. Combining is just to the process of turning grey scale (i.e., Black and White) images into a colour one. Enhancement covers a number of processes from enhancing faint detail, making stars smaller and less dominate and even to adding diffraction spikes!

There are a number of general and astronomy specific software packages for image reduction. Some, for example, MaxImDL are both image acquisition and image processing packages, others, such as StarTools are for image processing only. Packages such as Photoshop and GIMP are non-specialist image processing software, and your will need some further software prior to process your images in these packages. Some packages, for example, DeepSkyStacked (DSS) and FIT Liberator, perform very specific reduction tasks, such as image format conversion.

9.1.1 Calibration

Calibration of images must take place before processing, i.e. the dark bias and flat frames must be applied. Figure 8.2 showed the pathway needed to create and apply. If you intend to stack images, you must apply the calibration frames *before* stacking. As stated previously most camera control software allows you to apply calibration frames (for example MaxImDL) as does many of the post-imaging packages, such as AstroImageJ and GIMP.[1]

The order of calibration is the subtraction of Dark and Bias frames from the light frame followed by the division of the resultant light by the normalised flat frame. The resultant frame is the science frame If you only have single images, at this point you may wish to perform hot and cold pixel removal. Bad pixels are not so much of an issue if you intend to undertake frame stacking.

Calibrated image should be visually checked for quality. Check the images might be time-consuming if there are a large number. However, stacking a poor image adds nothing to the overall image quality and flaws such as vignetting and aircraft trails are difficult to remove except by exclusion. Images may be checked for quality before calibration. However, a poor or misapplied calibration frame might go undetected if they are not checked post calibration.

9.1.2 Stacking

Once the calibration frames are applied and you have the science frames you will need to stack them if you have taken multiple science frames in the same band. In order to stack you need to align the frames first. There are several packages that

[1]GIMP requires a plugin in order to undertake calibration.

can do this including MaxImDL, Deep Sky Stacker (DSS) and AstroImageJ; both DSS and MaxImDL will align and stack in what appears to be a single operation, although of course, it isn't. The easiest alignment method and in my option the most successful is to use the WCS in the header to align, although this is only possible if the frames are plate solved. Pattern alignment is used by both DSS and MaximDL. Essentially an algorithm identifies the stars in the field and uses them to make a series virtual boxes with stars at each corner. There *boxes* are then matched by rotation, translation and scaling of the frames, until they match. An alternative to this is aperture or star matching. Here the user does the hard work by identifying the same group of stars within each image, the software using this data to find the required transformations needed. Lastly, some applications have the option of comet alignment. Comets and other solar system boxes tend to move against the background stars over multiple frames. Traditionally stacking methods would stack the stars at the expose of the comet or another body. Comet Stacking (or if this is not available, single star matching, where the Solar System Body is selected rather than a star) stacks to the body being observed rather than the background stars.

Once the frames are aligned, they need stacking. If this is going to be done during the alignment process it is vital that the stacking process is selected *before* the alignment is started. There a number of ways to stack a series of frames, collectively know as, not unsurprisingly as a stack. The process you use will depend on the frames, the target and what operations are going to be undertaken after stacking.

The simplest form of stacking is **Summing**. This process simple performs pixel to pixel addition that keeps the pixel information linear and, if stacked to a higher bit FITS, can be used to stack images with stars that would otherwise saturate. Summing is ideal when photometry is being undertaken. Ensure that if photometry is intended that the exposure time is set to the sum of all the co-added frames. Hopefully, the software will do this, but it doesn't always.

Mean stacking, as the name implies, is the, same as summing but the final pixel value is divided by the number of frames. This process will result in a linear science frame if all the frames in the stack have the same exposure time. Mean stacking has a positive effect on bad pixels if the image is dithered, as the effect of the bad pixel is spread over several pixels.

Median stacking uses the median value of the pixel to pixel stack. Median stacking has similar results as mean stacking, although I feel that it handles artefacts, defects and noise better. However, it does not preserve pixel linearity so if you wish to do photometry, I suggest avoiding Median stacking.

Sigma Clipping or Kappa Sigma Clipping, determines the standard deviation over the pixel stack and remove pixels with values displaced more than an a set number of standard deviations, Kappa, away from the mean before reducing the pixel value to the mean. Sigma clipping is therefore very effective at removing cosmic rays or bad pixels as there tend to have out laying values on a single image, assume you have dithered. However, it produces a non-linear pixel response.

Drizzling is a stacking technique developed by Richard Hook and Andrew Frucher in order to improve resolution. In essence, the process stacks the images

Fig. 9.1 The stacked image of NGC281 (the Pacman). The stack consisted of 15 300 s hydrogen α frames. The stacks were aligned using the WCS and stacked using $\kappa = 3$ sigma clipping (Image courtesy University of Hertfordshire)

to a much finer grid that imposed by the pixel layout of the CCD. This process should improve resolution for images that are under-sampled i.e. the pixel size is larger than the seeing.

9.2 Stretching the Curve

We have already encountered stretching when undertaking initial imaging. Stretching is the process of changing how pixel value maps onto the display. Stretching is required because the full range of the pixel value far exceeds the ability of the human eye to perceive small ranges in brightness or a monitor's ability to display such changes. Effectively we want to bin the pixel values. For example, most images are typically as 8-bit, the display a brightness level from 0 to 255. For colour displays, this would be three or four values (red, blue and green and sometimes a black level). Clearly this is much smaller than the pixel range within your camera, so we might decide, as anything less than the mean pixel value is likely to be background, to set any pixel with a value less than the mean to a display value of zero. Likewise, anything near the high end of the pixel value is likely to be a star, so we might with to set any pixel value in the top 10 % of pixel values to represent a display value of 255.

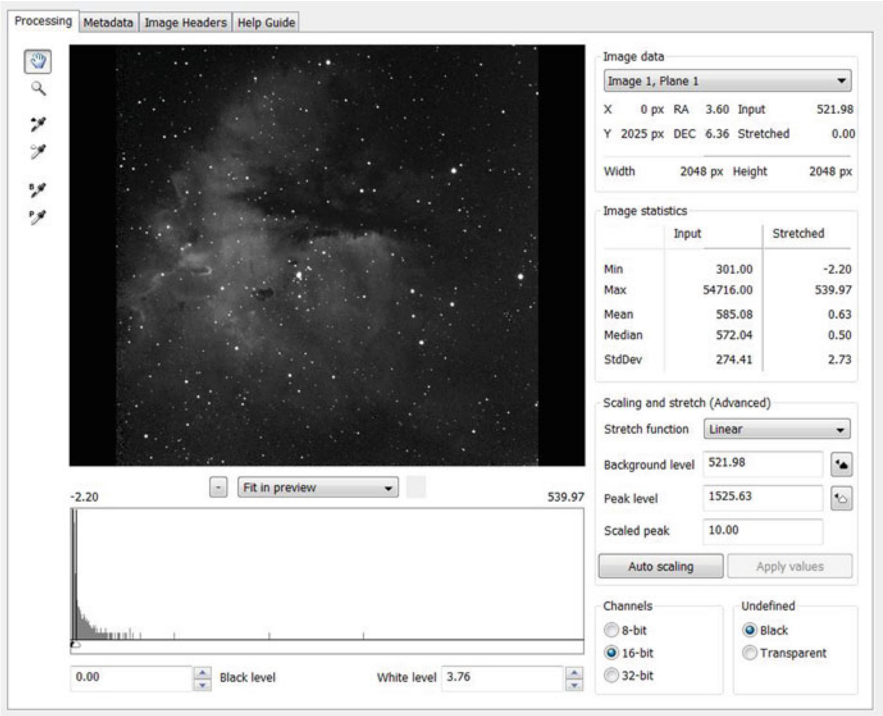

Fig. 9.2 Example of a manually stretched image, in this case, Fig. 9.1 being stretched in FITS Liberator (Image courtesy University of Hertfordshire and ESO)

There are a number of ways we can perform a stretch. The simplest, and the one we have already encountered is to stretch manually using the pixel value distribution histogram. The Histogram shows the most common pixel values, with the peak typically being the background. By moving the upper and low display limits we reduce the pixel value range that each display pixel corresponds to therefore enhancing detail. Stretching in this manner is known as a **Linear Stretch**.

An alternative method is to apply a built-in non-linear stretch to the entire image, the most common of which is a **log stretch**. A log stretch enhances low pixel value pixels while reducing the impact of high-value pixels. In a log stretch, a pixel with a value of 10 would be translated to a value of 1 while one of 60,000 would be translated to 4.77, for a Log10 stretch. Other pre-set stretches include Sin, LogLog and power law stretching. After a non-linear stretch has been applied it is normal to have to adjust the display range, i.e., to make a linear stretch on top of the non-linear stretch, in order to get the most of the image (Fig. 9.3).

If you have both faint and bright and faint structure, then linear and built non-linear stretching are going to be unsuitable and a curve stretch will be needed which can be done both before and after the colour image is built.

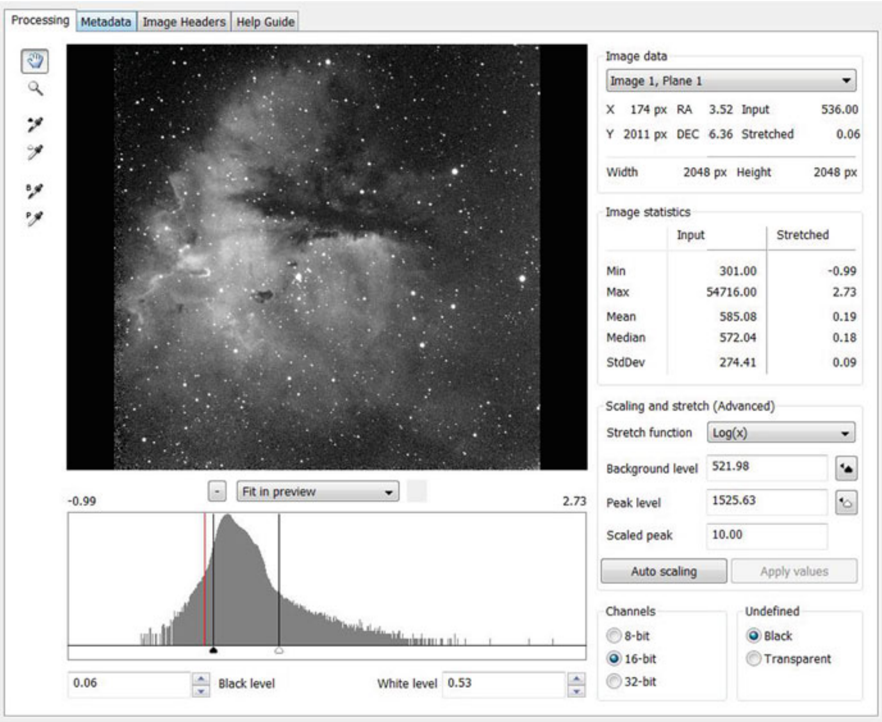

Fig. 9.3 Example of a Log stretched image, in this case, Fig. 9.1 being stretched in FITS Liberator. Note the difference in the pixel value histogram between this stretch and the linear stretch shown in Fig. 9.2. Also, note that there is a manual stretch on top of the log stretch. The whole histogram range is not being displayed (Image courtesy University of Hertfordshire and ESO)

Both MaxImDL, Photoshop, Gimp and a number of other image processing packages have this function. The principle is that the relationship between pixel value and the display value is plotted as a line and that line is manipulated, in order to enhance the required details. Where the gradient of the line is flat, there is high compression of the pixel value while where it is steep there is low compression. Hence, the pixel values covered by the high gradient area of the curve are enhanced and the ones in the low gradient area are suppressed.

For example, looking that the pixel value histogram for Fig. 9.1 we can see a main peak and a secondary peak with a pixel value around 3,800 counts. By a modified log stretch using the curve option in MaxIm DL both the pixels covered by the principal peak in the histogram and those in the secondary peak can be enhanced; the results of this process can be seen in Fig. 9.4.

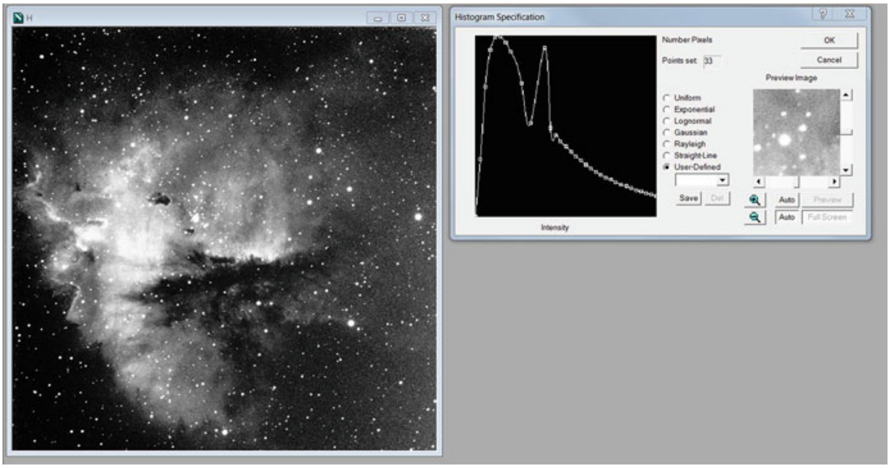

Fig. 9.4 Figure 9.4 above shows the effect of apply a stretch curve to a FITS image. In this case, the image is Fig. 9.1 and the curve is being applied in MaxImDL (Image courtesy University of Hertfordshire and Diffraction Limited)

9.2.1 From Grey Scale to Colour

Once you have your individual filter images reduced and stretched, you may wish to combine them to make a colour image. Typically you will make a colour image from three images and assign a red, green and blue channel to each. Normally you would assign a filter to the channel that most represents that filter pass band. So for example an image taken with a Johnson V filter would be assigned to the green channel while an I filter would be assigned to the red channel.

Of course, it is not always possible to simply match a filter to a colour channel, for example for narrow band filters. There are conventions to help. For example if you have V, B and Hα the Hα is assigned to the red channel. If you already have an R filter set the Hα may be stacked with the R band image a four colour image could be created. For the standard H$-\alpha$, OIII, SII set there is a standard known as the **Hubble Palette**, where SII is mapped to the red channel, Hα to green and OIII to red.

The majority astronomical imaging processing software will let you assign a frame to a channel, align those frames and then turn them into an RGB image. With Non-Astronomical packages such as Gimp and Photoshop you have to load the images as separate layers and assign colour channels to each layer, then align the channels by hand before combining them into a single colour image. You should be aware that Photoshop does no handle FITS files and GIMP does not always handle them well. If you are using an imaging processing package, I suggest therefore that you convert you files from FITS to an uncompressed standard format such as TIFF. FITS Liberator is the ideal tool to undertake this conversion and to perform you initial stretch.

In some cases, such as when a target has a lot of fine structure it may be possible to add a fourth **Luminance** channel. A Luminance is a low signal to noise frame, often taken without a filter. Adding this as a luminance channel to create an LRGB image enhances fine detail lost in lower SNR, filtered images. Many amateurs spend as much time on the Luminance frame as on the rest of the image. If you do not intend to use the images for science, then a Luminance is strongly recommended. If you do intend to use the images for science the time spent on the Luminance can be best used for improving the SNR of the science frame. You can imagine that in an RGB image each pixel has three values assign to it, representing the mix of red, blue and green. In an LRGB, image there is a forth value which tells the display how bright a pixel shall be. As the human eye is much more sensitive to changes in brightness than it is colour, the effect of adding the luminance frame is to hence the difference in brightness between regions.

Be aware that when processing LRGB images it is good practice to work on the combined RGB frame and the Luminance frames separately in order to get the colour balance and stretching correct. You then break up the RGB frame into separate channels and then recombine into an LRGB.

There are almost as many approaches to creating an RGB (or LRGB if we have a luminance frame) as there are software packages. For example, MaximDL makes the transition from single images to colour image seamless buy using a one-step approach. Users select the image combination they wish for example RGB, LRGB or CYMK (Cyan, Yellow, Magenta and Black for litho printing) and then assign a channel to each one. Once accepted Maxim will created the image which can be saved in most standard formats.

AstroImageJ takes a rather differing approach. In AstroImageJ (which cannot handle luminance data yet) the individual frames the are loaded into stack. We can consider a stack as a cube which the x and y-axis being the pixel position and the z-axis being the frame, or slice in AstroImageJ terminology, number. The user assigns a colour to each slice and the cube is merged vertically to form a colour image.

Producing a colour image in Photoshop or GIMP is a little more complicated than using an astronomical imaging package. Hence, I will walk though the process for both applications.

I will start with Photoshop, I assume you will be using Photoshop CS6[2] and will be producing a LRGB image from four frames that have been previously stacked, aligned and converted to TIFF or JPEG. Open you four frames and then add a fifth new frame. The setting for that frame should be the same as one of your open frames it does not matter which. Set *Background Contents* to Background Color and *Color Mode* to RGB Color. Select the new image and the channel tab. You will see a list of for images, RGB, Red, Green and Blue.

Select the red channel and then select the red image. Select the whole of the red image with a ctrl-A and then copy it with a ctrl-C (or just used the Edit drop down

[2]Although this instruction should work with most previous versions.

menu). Now return to the new image and select *paste in place* from the Edit menu. Do the same for the green and blue channels until you have you colour image. You now need to add the luminance frame. Select the Layers tab then Layer New Layer. Set the new layer mode to Luminosity. This will bring up a second layer above your RGB image (which will be called Background). Select the new layer and then select your luminance frame. Select all and then copy then return to the new image and paste in place. You now have an LRGB image.

The process is somewhat similar in GIMP. Open you three RGB frames as we would in Photoshop except we do not need to open the luminance frame yet. Also, we do not need a fifth frame. From Colors select component and then compose. Choose RGB as the Color Model and select the three frames corresponding to the three filter colours; pressing OK results in a colour image. To add the luminance frame select the colour image and then select the luminance file from File Open as Layers. Highlight the new Luminance layer in the Layer bar and change its mode to value. Now left click on the luminance layer and select merge down. You now have a LRGB image.

9.2.2 Image Filtering

Now you have your colour image you may wish to enhance it in order to bring out the finer detail in the image. This can be done by applying any number of mathematical process to the image, a process known as, somewhat confusingly, as filtering. Filtering an image is fundamentally different to stretching an image. Filtering is non-linear and once accepted permanent. It does not change how the image is display, it changes the pixel values. Hence, when filtering an image you can no longer use it for the purposes of science. However for complete we will detail a number of the more common image processing filters.

- **High Pass** We can consider the data in an image to a series of waves representing the rate of change between pixels. If the values change rapidly between pixels, the data is high frequency. The High Pass Filer accentuates the high-frequency regions of the image, increasing enhancing these areas at the expense of increased noise.
- **Unsharp Mask** The Unsharp Mask performs edge detection and enhances the edges. In most implementations, you need to define the with of the edges, the aggressiveness of the process and the floor at which a pixel is considered the background.
- **Wavelets** The Wavelet process identifies specific frequencies (as selected by the user) within the image, and either enhances of de-enhances them. The process is very powerful but requires considerable skill to use.
- **Deconvolution** Deconvolution is a complex mathematical process that de-blurs the image by modelling the blurring from the Point Spread Function (PSF). If there is no blurring in the original image Deconvolution will have no effect.

- **Erosion** Erosion filters removes bright edges around stars making them both sharper and smaller.
- **Kernal Filters** A kernal filter applies a kernel, convolution matrix, or mask to the image. This is basically an matrix, the matrix is moved over the image which is multiplied by the values with the matrix. The effect of the Kernal Filters is dependent on the contents of the matrix. It may blur, sharpen emboss or enhance edges depending on its parameters
- **FFT Filters** An Fast Fourier Filter applies a Fourier Transformation to the image (remember an image can be treated as a series of waves). FFT Filters are used to reduce patterning on a image such as stripping.
- **Digital Development Filters** Digital Development Filters are a series of mathematical filters designed to make CCD images have film like response, i.e. non-linear. They bring out faint detail and reduce bright sources.

Chapter 10
Photometry

Abstract Photometry is one of the key observation pillars of astrophysics. New techniques have improved the precision of photometry to the point that Exoplanet transit observation are possible with small telescopes. This chapter details the methods by example to perform photometry, the reduction of images taken for photometric purposes and the interpretation of the results.

10.1 Introduction

Photometry, the measurement of the number amount of electromagnetic radiation received from an object, is one of the cornerstones of modern astrophysics. The ratio of two fluxes, in astronomy a colour, are an important method for classification and identification of stellar objects.

Time-dependent photometry is used to understand the physics of many astronomical objects such as Cataclysmic Variables, Accretion Discs, Asteroids and Cepheid Variables. Cepheid Variables are an important component of the distance ladder that leads to the measurement of the most distant objects in the observable Universe.

High precision time dependent photometry is one of the principle methods for the detection of planets surround stars, **Exoplanets**, used in astrophysics.

The advent of the CCD, coupled with increased computer power, has considerably improved the accuracy, precision and ease of photometric measurements. Modern small telescopes, such as those found in many university teaching observatories, are capable of the milli-magnitude measurements. Such precision is needed to observe and detect exoplanet transit and are once again making a contribution to science and our understand of the Universe.

There are fundamentally two approaches to photometry, Aperture and PSF and three approaches to the calibration of the results absolute, differential and instrumental.

When we measure the light from a distant object, we measure not only the light coming from the object but also from the sky, stray interstellar light and scintillation noise; in addition to the false signal caused by the remnants of the bias, dark current. These sources are collectively known as the **background**. The method by which we deal with the background is the difference between Aperture and PSF Photometry.

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In **Aperture Photometry**, we drop an aperture, which is normally a circle, but not always, onto the image of the target star. We then integrate the count constrained within that circle. The diameter of this circle is important. The circle should contain only pixels associated with that star and be wide enough to capture the vast majority of the signal from the target. The circle's diameter typically a circle with a radius three times the FWHM PSF of the star.

Remembering that the PSF should be Gaussian, you should see that an aperture of this size contains 99.7 % of the light from the star. Because the PSF is principally a function of seeing, the aperture size should be constant for the whole field. The difference between a bright star and a dim one on the image is not the width of the PDF but the height.

Clearly the integrated count within the aperture also includes a contribution from the background. It is not possible to determine directly which are from the source and which the background so we must assume that the background is constant at local levels. By dropping an annulus onto the aperture, where the inner ring of the annulus is beyond the aperture we can address the problem of a noisy background. Typically the area contained within the annulus is the same as that enclosed by the annulus, but if not it should always be larger. We can take the background to be the integrated count within the annulus normalised for the area of the aperture. Hence, stars falling into the annulus must also be avoided as they would artificially heighten the background. We subtract this area normalised background from our integrated aperture count to in order to get an instrumental count. As the count will depend on the exposure time, we need to divide the count by the exposure time to get counts per second.

The key drawback with Aperture Photometry is that it requires an aperture and annulus, without contamination from other stars, be fitted to the target. In very crowded fields, such as those found is globular and open clusters, this just is not possible.

The solution to this is PSF fitting. Given that a star's profile is Gaussian and a Gaussian is determined by its width and height it is possible to estimate the count by fitting a Gaussian to the PSF. Hence PSF Photometry. In this way, even if the PSF of two stars overlap we can achieve a reasonable estimator of the count. In most cases of PSF photometry, the PSF is used to detect and remove the stellar contribution to the entire image count and the remainder is used to determine the background.

In general PSF, photometry is not as accurate as aperture photometry in ideal conditions but outperforms aperture photometry in crowded fields. Many applications used for photometry can do both PSF and aperture photometry and you should judge which one you should use based on the field and the FWHM of the stars in the image.

You may have thought about the implications of unresolved binaries on photometry, two or more stars are so close that they appear as one. Fortunately as luminosity increases by approximately by $M^{3.5}$, low mass companions contribute little to the overall luminosity of the system. Even in a system comprising of a pair of stars of the same mass, the colour is not overly affected. Although any distance based on the measured brightness of such a system would be less accurate than a single system.

There are two approached to calibrating raw **instrumental magnitudes**. The first of these is Differential Photometry, which uses references stars, preferably of the same spectral class and in the same field of view and hence imaged at the same air mass.

Differential Photometry is often used to determine the magnitude of variable stars and to identify exoplanet transits. The magnitude of the reference star does not need to have been accurately determined (as long as it isn't variable itself).

The measurement needed in Differential Photometry is the difference in magnitude between two or more source, δMag, rather than an apparent magnitude.

Very high accuracy and precision results can, and are produced using this method. Determining objects apparent magnitude differentially to anything better than a few tenths of a magnitude is challenging, mostly due to the dearth of stars with very well determined magnitudes. Both the Landolt (UBVRI) and Henden (BV) catalogues provide well determined Standard Stars. However, these are often not accessible for observers at high latitudes.

The second method, **All-Sky Photometry** uses a well-calibrated optical system to determine magnitude directly from images. All-Sky Photometry is mostly used for large-scale surveys and also for regions well away from a suitable calibration star. Photometric accuracies of a few hundredths of a magnitude are possible for a well determined system, which is a factor of 10 poorer than that achieved by Differential Photometry.

In the rest of this Chapter, I will concentrate on Differential Aperture Photometry as this is the most commonly encountered by undergraduates. Differential Aperture is versatile and many of its aspects form the foundations of other forms of photometry.

10.2 Measuring

In this section, I will walk through how to do standard aperture photometry that will yield an uncalibrated instrumental magnitude using the **Aperture Photometry Tool (APT)**. APT is freeware and, as it is writing in Java, will run on almost any platform. Later on I will be using a similar programme, AstroImageJ to perform differential photometry. Like APT, AstroImageJ is freeware and multi-platform. APT is more suited to field photometry, measuring all the stars in a frame while AstroImageJ is more suitable for time series photometry, measuring just a few objects but over many frames.

10.2.1 Aperture Photometry

Before we start the photometry, I suggest checking the FITS header of you image to determine if has been plate solved. If the header contains cards with the heading CTYPE1, PLTSOLVD or PA, which are WCS keywords, then it has been plate

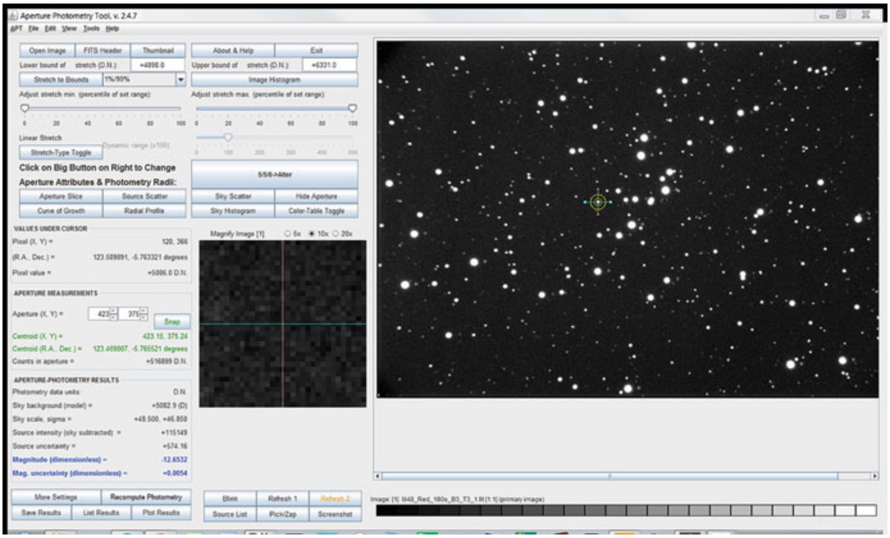

Fig. 10.1 Screen shoot of Aperture Photometry Tool with an image of the open cluster M48 loaded and a star selected with an aperture and annulus overlaid (Image courtesy of the University of Hertfordshire)

solved. This will make things a lot easier later on. If it has not, you can either use a local plate solving service, which some applications have, or you can push the image up to the astrometry.net website for solving. It is not necessary to do this, but in my experience it is highly desirable.

Once an image is open in APT we need set the parameters for our photometry, which can be found in the more setting button. We select Model 2 and Model D. This gives us sky subtraction and takes into account that some pixels are intercepted by our apertures.[1] When you now left click on a star you get a series of concentric circles displayed, this is the aperture and the annulus (see Fig. 10.1). Most likely these are set incorrectly. We need to know the FWHM of the frame in order to set this correctly. If you have used an auto focuser or the image is plate solved you might find the FMHW listed in the header which should be in pixels (but always check). If not you will need to plot the PSF of a field star. In ATP, this is just done by selecting a field star and pressing the Aperture Slice Button.

In ATP, this displaces both an x and a y slice through the PSF of the selected star. The star might not be completely circular so the two slices may not match in which case just select the widest of the two. In the case of Fig. 10.2, half the count of the peak is about 14,000 as we need to the subtract the background before we half the height and then add it back. The full width at this point is about 5 pixels (which happens also to agree with the number in the header). Good practice is to have diameter of the aperture to be three times this, 15 pixels.

[1]Do not forget to press the apply button!

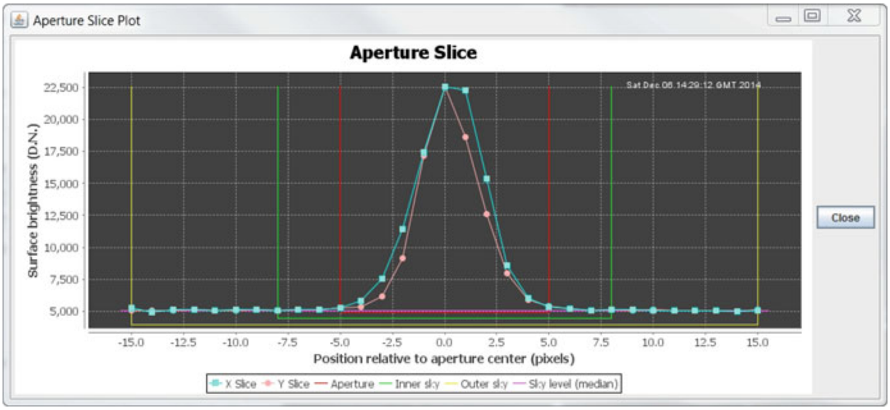

Fig. 10.2 Screen shoot of Aperture Photometry Tool Aperture slice showing the PSF of the star selected in Fig. 10.1

The radius of your aperture is set under the Big Button, which brings up a popup with a direct entry box and a slider being used to set the aperture radius. The multiplier radio buttons (1×, 5×, 10× and 20×) set the size of the annulus based on the aperture size. As the aperture radius is integer I would set this to 8, rather than 7. Note that unchecking Center(x,y) lets you make elliptical apertures that are useful when doing extra-galactic photometry.

Now when you click on the stars, in the image window, you will see in the Aperture Photometry Result Panel. You might notice the magnitudes looks strange, often negative. The magnitude is negative because this is an instrumental magnitude and you will have to calibrate our data to into apparent magnitudes later.

Although, APT reported the magnitude, it is the Source Intensity (Sky subtracted) and the Sky Scale Source Uncertainty that we wish to record, along with the RA and Dec of the object. Be aware some software time normalises the results

10.2.2 Photometric Reductions

If you need photometry in more than one filter, you will have to undertake some form of photometric reduction. These reductions turn your counts into reduced standard magnitudes account for the instrumentation you are using, under the conditions and the air mass at which the observations took place. For most purposes, there should be stars within the field with known magnitudes in the bands you are observing in. If not hopefully you have images of nearby stars with known magnitudes and at the same air mass as the target.

The reduction process outlined here and in the following practical is discussed in considerably more detail in Brian D Warner's excellent book, "A Practical Guide to Lightcurve Photometry and Analysis" also published by Springer. There are techniques that claim better accuracy, but the gains are marginal given the complex nature of photometric reductions and the inherent uncertainties in small telescope system that are often based in non-ideal sites.

When imaging for photometry purposes it is best to take multiple images very close together in time. Multiple imaging increases the number of data points in your photometry and, therefore, reduces uncertainty and ensures that they all taken at the same air mass. These advantages apply to reference star images as well.

Previously in this Chapter I discussed how to set an aperture size and measure the integrated count within the aperture and how to account for the sky. Your software, should unless it's very simple like DS9, be able to automatically time normalise the data. If it does, not do not forget to divide the background subtracted count by the exposure time in seconds. You should determine the instrumental integrated background subtracted count for all the reference stars you are using and for every image. Take the mean count value for each star and use this as the instrumental count (and also calculate uncertainty from this). The mean instrumental count needs to be changed to an instrumental magnitude by using a modified version of Pogson...

$$m = -2.5\log(count) \tag{10.1}$$

where m is the raw instrumental magnitude and count is the count within the aperture less the background.

The count does not represent the number of photons received rather a ratio between the electrons generated and the **gain** of the camera. You will find the camera gain in the camera manual or written into the FITS header as EGAIN. Note that gain might be also determined by the binning. The gain is applied to the count before it is logged. Hence, Eq. 10.1 becomes...

$$m = -2.5\log(count \times gain) \tag{10.2}$$

and your errors become

$$\delta(m) = -2.5\log 1 - \frac{count \times Gain}{\delta count \times Gain} \tag{10.3}$$

where $\delta flux$ is your calculated count error.

If you now perform aperture photometry on a star in a science frame and apply Eq. 10.2 you will find you obtain an odd result, most likely a magnitude with a very negative value. The reason for this is that your photometry is not reduced. The process of reduction takes your raw, instrumental magnitude and converts it into a standard magnitude, one for a standard set up observing above the atmosphere.

The basic reduction formula in magnitudes is...

$$M_f = m_f - k_f' X + T_f(CI) + ZP_f \tag{10.4}$$

where:- M_f is the standard catalogue magnitude of the reference star in band f, m_f is the instrumental magnitude of the reference star in band f, k'_f is the extinction coefficient in band f, X is the airmass, T_f is the colour correction for band f, CI is the colour index and ZP_f is the current band f zero point.

The **Zero point** is an offset between the instrumental magnitude and the standard magnitude. In broad terms, it is a measure of the sensitivity of your optical system to light. Hence, space-based telescopes also have a zero point. A zero point can be either a flux or a magnitude. In the case of the flux, it is applied by dividing the time and gain as normalised count by the zero point. If it is expressed as a magnitude, it is just added to the instrumental magnitude as per Eq. 10.4.

The **Extinction Coefficient** reflects the proportion of light that is lost as it travels through the atmosphere, for a given air mass and hence is wavelength and, therefore, filter dependent as well as being altitude specific. The **Colour Index** is the ratio of flux received from a star over two bands, or, more understandably the difference in magnitude between two bands for that star. So, for example, the star HIP 2027 has an R magnitude of 7.574 and a V magnitude of 7.640 and hence a V-R colour index of 0.065.

Different wavelengths are subjected to differing amounts of extinction as they pass through the atmosphere, with more blue light being scattered than red (hence the sky appearing blue). A blue star will, therefore, be affected by the atmosphere more than a red. Hence the need to apply an extinction coefficient that is airmass and filter specific. Note that the air mass is a function of the altitude at which the object is observed, applying...

$$X = 1/\cos(z) \tag{10.5}$$

where $z = 90\,\deg - altitude$ gives the air mass X. Note that this the altitude angle of the object not the altitude of the observer.

The colour transformation $T_f(CI)$ is effectively stable and can be used on successive nights unless there have been significant changes in the optical system, likewise the extinction coefficient. The Zero point is related to the entire optical pathway and will change between individual nights of observing so it should be recalculated for every session. The airmass X is a function of the actual observation.

You may have the terms first and second order extinctions in the context of photometric reductions. The first order applies to the air mass coefficient whilst second order applies to the colour transform. In some cases, the first order extinction can be ignored if you are using calibration stars at the same air mass. However, as this might not be the case and given that they are largely immutable it is worth having their values at hand.

The following practical will guide you through the process needed to identify the individual components of Eq. 10.4. Once this is completed, you will only need to identify the Zero Point during subsequent observations while the optical system is unchanged.

10.3 Practical: Find the Zero Point and Transformations

Although raw, instrumental magnitudes are useful if you wish to undertake observations that may span more than one night or that require colour information. You will need to reduce your photometric data so that observations between nights, or objects of different colours or observations at differing air masses, can be compared. In order to transform instrumental magnitudes into standard magnitudes, you need to find your systems extinction coefficient, colour transformation and zero point.

This Practical draws on many of the skills you will have obtained in previous practicals especially the Messier Marathon practical, The Bias and Dark Frame practical and the Flat Frame practical. If you have not done these, it is recommended that you read them before attempting this practical.

10.3.1 Aims

In this practical you will find the nightly zero point of an image as well as the colour transformation and extinction coefficient for you optical system. You will also need to determine levels of uncertainty in standard magnitude when these parameters are used to convert instrumental magnitudes.

10.3.2 Preparation

In order to achieve the practical aims for you will need to image three fields in at least two photometric bands, preferentially B,V and R at a single binning value. These fields are the First Order Field, the Second Order Field and the Uncertainty Field. The fields you will need to observe should contain a number of photometric standard stars. Ideally these will be Henden, Hipparcos Red-Blue Pairs or SDSS Blue Red Pairs (lists of these are carried in Light Curve Photometry and Analysis Warner 2007). Your Second Order Field will need to be imaged as it crosses the local meridian at least three times, without over exposing the photometric standards. The First Order fields will need to be imaged a number of times at differing air masses, again I suggest a minimum of three times per air mass and at a minimum of three air masses. Again you should ensure that you do not over exposure the photometric standards.

You will need an Uncertainty Field, again containing photometric standard, that will be used to determine the uncertainty in your photometry calibrations. The calibrations should be imaged close to the meridian but not at the same air mass as either of your other two fields.

Once you have identified your fields you will need to produce finder charts with the photometric standard marked with their magnitudes and the time that you should start and stop imaging and the exposure time for each band.

Unlike most of the other practicals in this book this practical requires very good photometric conditions. There can be no moon at the time of imaging and the weather must be very good with no cirrus. Good seeing is desirable but not so important as the Moon status or weather.

10.3.3 Imaging

When you arrive at the observatory as usual, get the camera down its operational temperature and ensure it is steady. You will need to take multiple flat, bias and dark calibration frames in order to make master frames that you will need to apply calibration frames your lights to get science frames. Normally a flat field does not make a significant contribution to photometric uncertainties, unless one isn't used, misused or is of a very poor quality. This practical is an exception and so you will need good and recent flats.

1. As normal, before you start science imaging make sure the telescope is focused, that the binning is set correctly, the auto guider is operating correctly and that the camera is not sub-framing.
2. Slew the telescope to the first field some time before you intend to image and take a point image to ensure that the telescope is point at the correct field and is centred.
3. As the target reaches the required altitude start imaging. Take the images as a filter sequence i.e., VRBVRBVRB rather than VVVRRRBBB.
4. You will have some time between image sets. Use this time to check that the images are of an acceptable quality, with no trailing, overexposed stars or plane trails.
5. Once you have imaged all your First and Second Order Fields together with the Uncertainty field, shut the telescope down (if required). Ensure you have all the data on a removable format, including the calibration files.

10.3.4 Photometry

1. In order to perform photometry you will need to turn your light frames into science frames as outlined in Chap. 9.
2. Once you have your science frames you will need to find the seeing of each image. If you are using APT, this can be done by dropping an aperture of a star and taking an aperture slice from which you can get the FWHM. In MaxImDL, you need to use the Graph function (found under View, select line and draw a

line through a star. Doing so will give you a slice of the PSF from which you can obtain the FWHM. In AstroImageJ, this is done by drawing a line through a star and performing a Profile Plot (found under Analyse. In theory, the FWHM should be the same for all of your images. If it is significantly different then do not use the outlying images.
3. You now need to set the aperture and annulus size. Set this to three times the Half the FWHM. In general it is best to let the software determined the size of the annulus but if it does not five times the aperture radius is normally a suitable radius. If you have a very dense cluster, you may wish to have a much smaller annulus size in order to avoid contamination from nearby stars.
4. Using ATP you will need to set the source algorithm to Model 2 and the Sky algorithm to Model D.
5. Drop an aperture on each of your reference stars and note the source count and the count uncertainty. In APT, this is done by double clicking on the target. box. APT uses a very sophisticated uncertainty model that goes beyond a simple signal to noise ratio. If you are not using APT to do the photometry use the SNR to calculate the uncertainty instead.
6. Using the camera gain, which should be in the FITS header, time normalise the count, if needed, and apply Eq. 10.2. This will give an instrumental magnitude for each reference star in all fields.
7. Record each instrumental magnitude along with the target stars position, the filter in which the image was taken, the air mass and the magnitude of the standard star.

10.3.5 Reduction

There now follows a series of steps that will find the various components of Eq. 10.4 for your system. These processes require the plotting of data and the fitting of a line to that data. In Warner, it is suggested that a spreadsheet is used. You could use a specifically designed application, such as PhotoRed or perhaps a mathematical application such as Matlab to plot the data and fit the line. However, whatever you use, you should always visualise the data. It is easy for outliers to change dramatically the straight line fitting. If you see an outlier try to determine what caused it, it might be a hot pixel, a cosmic ray or even a misidentified reference star, and address it if possible. If you cannot correct it, you can remove it from your dataset as long as you make a note of it in your lab book, with an explanation why it was removed.

In this section I will be assuming that you are using V and R filters for notation, with V and R being the standard magnitudes of the reference stars. The notation v and r is used for the instrumental magnitudes and $< v >$ and $< r >$ the mean instrumental magnitudes. If you are not using V and R just substitute.

10.3.5.1 Second Order Extinction and the Zero Point

In the context of a V filter observation, we can rearrange Eq. 10.4 to. . .

$$V = v - k'_v X + T_v(CI) + ZP_v \tag{10.6}$$

As, the Second Order fields are all imaged at the same air mass we can safely ignore the first order part of Eq. 10.6, reducing it to. . .

$$V - v = T_v(CI) + ZP_v \tag{10.7}$$

You should be able to see that Eq. 10.7 is the equation of a straight line with $T_v(CI)$ being the gradient of the line and ZP_v being the intercept.

You need to find the mean values for your v band and r band instrumental magnitudes ($< v >$ and $< r >$ respectively) for each standard star in the field. Plot $V- < v >$ against $V - R$ and Fitting a straight line to this plot yields the V band Colour Transform T_v as the gradient and the V band Zero Point ZP_v. Performing the same plot with $R- < r >$ against $V - R$ produces the Colour Transform and the Zero Point for the R band.

Once this is done, you need to find the **Hidden Transformations**. These are effectively adjustments from you instrumental set up's colour index to the standard one. They are hidden because that do not appear in Eq. 10.4. They are found by plotting $V - R$ against $< v > - < r >$ and fitting a straight line to give T_{vr} as the gradient and ZP_{vr} as the intercept. Undertaking the same process for $R - V$ against $< r > - < v >$ gives T_{rv} as the gradient and ZP_{rv} as the intercept.

10.3.5.2 First Order Extinction

We now move on to our First Order Field. We know that the Colour Transformation and the Zero Point are air mass independent and that they are constant during a night's observation. So, if we vary the air mass at which we observe calibration stars, the change in instrumental magnitude when adjusted for Colour Transformation and Zero Point, is due to the change in X. Hence, we can find the Extinction Coefficient.

In order to determine K'_v we need to plot the adjusted mean instrumental magnitude $< v >_{adj}$ against air mass.

$< v >_{adj}$ and $< r >_{adj}$ is found by applying. . .

$$< v >_{adj} = V - v - (T_v \times V - R) \tag{10.8}$$

Plotting $< v >_{adj}$ against X and fitting a straight line gives the First Order Extinction Coefficient k'_v as the gradient, which should always be positive. It should also always be equal to, or greater than, k'_r. The interception point should be very close to the Zero point of the filter in question.

10.3.5.3 Error Analysis

You will now have to look at the uncertainties in your magnitude calibration system. Because the steps involved are complex the easiest way of determining the uncertainty is by comparison with known magnitudes.

Using the instrumental magnitude for each of your calibration stars in your Uncertainty field calculate the instrumental colour index for each calibration star using. . .

$$CI_i = (((v - k'_v \times X_v) - (r - k'_r \times X_r)) \times T_{vr}) + ZP_{vr} \tag{10.9}$$

where X is the air mass for that image and filter. Use the mean value for each star observed for the star's instrumental Colour Index, CI_i.

You can now transform the instrumental magnitudes to standard magnitudes by using the mean instrumental magnitudes for each calibration stars and applying. . .

$$M_v =< v > -(k'_v \times X) + (T_v \times CI_i) + ZP_v \tag{10.10}$$

By looking at $M_v - V$ and you will be able to determine the level of uncertainty in your photometry and see if there are any systemic errors.

10.3.6 Analysis

As usual you should write your analysis in a standard format including all the data, plots and calculations. Give special attention to discussing the level of your uncertainties given that a typical survey has photometric uncertainties errors of 0.1 mags. Where do you think your errors are and how do you think you could improve them.

In your discussion include how you would be able to determine just the Zero Point if you were observing on another night. Observers of exoplanet transits might need photometric uncertainties of 0.001 mags. How do you think they can achieve this given the levels of uncertainties in your results?

10.4 Differential Photometry

One of the questions asked in Practical 8 is how are exoplanet transits detected if millimagnitude precision is needed to give the level of uncertainty in our photometry as determined in the Practical. The answer lies in the Second Order Extinction Coefficient determination. When we image the reference stars in the same field as the object of interest, we can ignore the First Order coefficients. For most observations we are not interested in the change of colour of an object over time, although that itself is interesting and contains a considerable amount of physics. Rather we are interested in the change in magnitude ΔM in a single band. As both the Colour Transform and the Zero Point are time independent, these can be ignored under these circumstances. Hence in this case we are only interested in the difference between the instrumental magnitude of the target and the reference, although in general multiple reference stars are used in order to reduce errors. Obtaining millimagnitude uncertainties using differential photometry is possible using a small telescope. However, you should be aware of the sources of errors so they can be minimised.

There are a number of sources that contribute to system noise with regards to photometry. **Read Noise** (σ_{read}) is simply the noise generated by the electronics during the read process. It is effectively the noise within the bias frame (although not the bias current that is systemic). The bias will vary across the chip and to some extend between bias frames. In general, the bias noise is small, with much of the variation taken out by using a master bias, i.e., the mean of a number of biases. However, if the bias associated with the pre-exposure dark the bias noise should not be a limiting factor unless very short exposure images in use.

As discussed in Chap. 7.1, the dark current is the build up of thermal electrons within the CCD. It is effectively but not completely, removed by the subtraction of the dark frame. So it is another source of random uncertainty (σ_{dark}). Again the use of master dark frames to reduced dark current noise. Remember however that a dark frame includes a bias. Although the dark current will not vary across the frame as much as the read noise it is, unlike the read noise, dependent of exposure time and operating temperature.

Flat fielding, as discussed previously, is a challenging issue. For bright objects, the quality of a flat field and the noise of the flat field **(flat field noise)** normally is not significant. However, for a faint source or when measuring sources in multiple locations in the frame, it can become significant because it causes variation across the field due to the field not being entirely flat. Therefore, a variation in pixel location can result in a different pixel response.

In general when performing single target photometry, the best practice is to keep the target in the centre of the frame as this is where the frame is most flat. However, when undertaking differential photometry, this might not be possible for both reference star and target. We might require moving between the source and the reference, **nodding** so that each appears on the same pixels. Thereby reducing flatfield effects.

As stated previously, because of atmospheric effects the light from a star does not appear as a point source occupying one pixel, but rather a circle (hopefully) occupying multiple pixels. However, pixels are square and are spaced with gaps between them, a small amount of light will fall between them and the edges of the star will cut through pixels rather than completely fill them. This leads to **Interpolation Errors**. In the most part, the software will deal with this problem when undertaking photometry, but users should be aware of the problem, if, for example, we are using DS9.

Another photometric precision problem is scintillation noise. Effectively small scale turbulence in the atmosphere scatters the light causing the random, non-detection, of photons. Scintillation occurs on small scales of about 30 cm and over time scales often less than a second. Hence, it can be reduced with larger aperture telescopes or increased exposure time. It is interesting to note that, for larger telescopes, increased exposure time is often not an option, and it would lead to saturation.

Our last source of noise is **Sky Noise**. Random scattering, absorption and emission of photons causes Sky Noise. It should be the largest contributor to the uncertainties and is the dominant source limiting how faint we can observe. Fundamentally, besides moving to a very high site, preferably in space, Sky Noise is going to the principle limit on the photometric precision.

It should be clear that, most of these sources of error, besides Interpolation Errors, form part of the background of the science frame. Hence, the best way to reduce these error's impact is to improve the signal to noise ratio.

10.5 Practical: Constructing an HR Diagram of an Open Cluster

The Hertzsprung-Russell is one of the most important concepts in modern astrophysics. Its construction tells us about the evolution of stars and can be used to determine the age of a star cluster by the identification of the main sequence turn off. Constructing and interpreting an Hertzsprung-Russell diagram is a vital skill for any undergraduate astrophysicist.

10.5.1 Aims

To construct an HR diagram and identify the main sequence turn off. The construction of the HR diagram will require taking and reducing images of an Open cluster in two bands. In additionally, you will need to undertake aperture photometry on both images and plot the results. You should also prepare a backup cluster.

10.5.2 Observing

Once at the Observatory you should follow the standard imaging procedures:-

- Cool the camera down to temperature and open the dome;
- Once down to temperature take flats if needed;

- Slew the telescope to a focusing and alignment star. Confirm alignment and focus the instrument;
- Slew to target. Take a short exposure image to confirm you are on target and still in focus;
- Turn on tracking and select first filter. Take several images of the target at differing exposure times until you are happy that you have a good linear image.
- Undertake the same procedure for the another filter;
- Take bias and flat frames.
- Save files to a portable media and shutdown Observatory

10.5.3 Reduction

Once you have the images, you need to reduce them, perform photometry and plot the results.

- Produce science frames by subtracting the bias and dark frames from your light frames and dividing by the flat
- Open the first science frame in ATP (or a suitable alternative);
- Select a star and perform an aperture slice. Use the FWHM of the slice to set the aperture and annulus radii. Set the Sky algorithm to median sky subtraction;
- Selected source list and then create a new source list. Process the source list;
- Hitting the List results button will display your photometry results. Export this file and then clear the contents
- Load the second image and repeat the photometry process listed above;
- Load both photometric results files into a spreadsheet. Move all but the centroid RA, centroid Dec, magnitude and magnitude error columns.
- If you know the filter's zero points apply it to the magnitude columns. This is not required if you do not have them;
- Cross match the two source lists (for example in TOPCAT) so that you have both magnitudes for most stars in the field (some will always be missing a magnitude, in which case discard them);
- Plot magnitude in filter A minus Magnitude in filter B verse Magnitude in Filter;

10.5.4 Analysis

Write your report up in the standard format. You should include all your plots and observational logs as usual. You should discuss your errors and their implications. You should identify the main sequence in your plot and the main sequence turn off. Discuss if all the stars you imaged are cluster members and the implications if they are not. What do you think is the benefits of imaging a cluster, rather than a normal star field?

Chapter 11
Errors

Abstract All measurements, no matter how carefully undertaken have errors. These errors need to be quantified so that differing experiments can be compared. Results may have precision, self-consistent, and/or accuracy close to the real result. Results may have random, and/or which Systemic Errors be addressed. This chapter expains the techniques for calculating and minimising the errors within an experiment as well as methods to propagate the errors when performing mathematical operations on the results.

11.1 Introduction

You will often see scientific results quoted as $x \pm y$ where y is the error. Errors also called perhaps more accurately uncertainties, is an expression of how much confidence can be put in the result. It does not mean a mistake has been made or that the scientists think their result is wrong in some way. In fact, all measurements are to some degree affected by uncertainness because a level of randomness is a fundamental part of the Universe as expressed by both quantum mechanics and chaos theory.

However, understanding how much of your result is comprised of random events is an important skill and you will be expected to express uncertainties for all your experimental results. Not only do the uncertainties express a level of confidence it also allows comparisons of differing experiments. Say for example one user does an experiment that results in a result of 2.5 ± 0.1, while another user uses a different method to make the same measurement and gets 2.6 ± 0.2. Although the quoted results differ in value, they are experimentally consistent due to their uncertainties.

The terms accuracy and precision are much more constrained in science than in day to day life, where they are used interchangeably. The term accuracy reflects how close a measurement is to its real value while precision refers to how self-consistent the results are. It is possible to be precise without being accurate and accurate without being precise. For example, later on in this book you will be asked to undertake differential photometry, in which you compare two stars, one with a known magnitude and measure the brightness difference. If the magnitude of your

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known stars were one magnitude out, your observations could be very precise but not accurate. When you use the words accuracy and precision use them correctly and with care; it is a very common mistake to use as if they are interchangeable – they are not.

Uncertainties can take two forms random and systematic. Random uncertainties are caused by random and unquantifiable effects, for example, cosmic rays and sky brightness. Systematic errors on the other hand cause a constant but unknown shift in the data, for example, this might be the bias on a CCD or to some extend the Dark Current. We don't know what the actual value is but we can allow for it.

Another common mistake is to confuse significant figures with decimal places. The number of significant figures is a measure of precision while the number decimal places reflect accuracy. Hence, 11.230 is a result to five significant figures but three decimal places while 0.001 is to three decimal places but only one significant figure.

When analysing your data, you should only use a level of accuracy that reflects your equipment and your ability to read them. For example the count of a pixel is always an integer, hence to talk about a fraction of a count is meaningless.

The are a number rules for dealing with significant figures and decimal places For significant figures all non-zero number are significant so 12.23 would be to four significant figures. A zero between non-zero digits is also significant as are zeros after a decimal point hence 101.25 and 101.00 are both to five significant figures. Between a decimal point and a non-zero number when the whole number is less than one, are not signification. Hence, 1.002 is to four significant figures but 0.002 is to just one (although it is to three decimal places)

Mathematics processes in regards to decimal places and significant figures are straight formatted. When you multiply two numbers, your results should always have the same number of significant figures as the smallest number of significant figures contained within your products. So, for example, $2.11 \times 3.13 \times 10$ should be 660.43. However our least significant figure is 10.[1] with two significant figures so our result should be written as 660. If the first figure dropped is five or great roundup else round down.

When you add or subtract figures, your result should have to the same number decimal of decimal places as the least accurate figure. Hence $2.11 + 3.11 + 10.$ should equal 15.22 however 10 is only accurate to two decimal places so our result should be 15.00. As with multiplication round the highest value dropped figure up if it is five or more, down if it less.

[1]Note that 10. is to two significant figures because of the decimal place, whilst 10 is to one significant figure and 10.0 us to three significant figures.

11.2 Handling Errors

One of the most common methods of finding uncertainties is to make multiple measurements and take the mean or median as the result and the standard deviation as the error. The Standard Deviation is also known as the Root Mean Squared (RMS) for reason that will become clears.

We can assume that, in most cases, our random errors will have a Gaussian Distribution, which plotted looks like Fig. 11.1. In this case, we can use the RMS to measure the distribution of the errors. The standard deviation of a sample if found by applying Eq. 11.1, where $\bar{x}$ is the mean of the sample, x_i is the value of the ith member and N is the number of items in the sample.

$$\sigma = \sqrt{\frac{1}{N}\sum_{i=1}^{N}(x_i - \bar{x})^2} \tag{11.1}$$

If our distribution of values is Gaussian, then 66 % of our sample will be contained within one standard deviation (1σ) of the mean. At 3σ it is 99.7 % and at 5σ that is the level consider for proof as 99.9 % of the sample is contained with the boundaries. The square of the standard deviation is known as the variance and is a measure of the spread and is always non-negative.

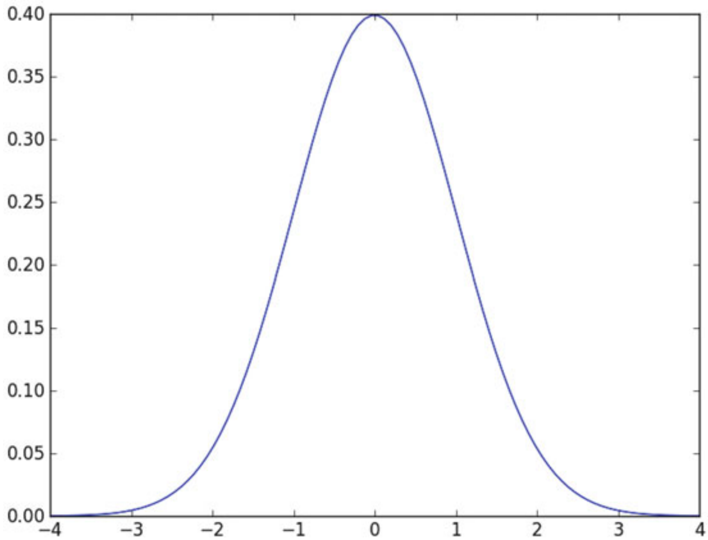

Fig. 11.1 A Gaussian plot. The centre of the peak is located at the mean value of the sample, the location of the peak is partly representative of systemic errors while the width is characterised by the standard deviation of the sample

When you are looking at your data beware of outliers. Outliers are results that appear on unexpected parts of the distribution curve, the plot of value against the count. Very large outliers can dramatically change your results, especially when dealing with small sample numbers. If you have an unexpectedly small or large result when compared the rest of the sample you may wish to investigate why this might be, often it is an indication of an actual error (i.e. a mistake has been made) rather than true result. If you cannot find the mistake, you may exclude that result but you must note it in your results as it might have been significant. In such cases may also be worth increasing your sample size to ensure that the it is indeed a statistical anomaly.

11.2.1 Systemic Errors

As stated above there are two kinds of errors, random errors and systemic errors. Systemic can also be divided into errors that can be addressed directly and those that cannot. For example, the bias added during the read process is largely unknown directly but is removed by the subtraction of a bias frame. Likewise, the dark current is offset that can be largely subtracted by the dark frame, although there is an additionally random element in the dark current. The non-uniformity of the illumination of the CCD and the pixel sensitivity are largely constant over short time scales and can also be consider in part systemic. Unfortunately perfect flat fields are extremely difficult to produce and quickly becoming dated as dust collects on the optics. So although flat field subtraction deals with the Systemic Errors there is always a degree of random errors inside a flat that we will see, can be significant. However, systemic errors will always be of the same sign and magnitude for all your readings.

There may well be other unadjusted or unknown Systemic Errors. For example the shutter on you camera might be sticking slightly so that you 1 s exposure is, in fact, 1.01 s. This will increase the number of photons you detect in 1-s exposure. Likewise, you might be using a reference star to determine the magnitude of a star with an unknown magnitude. If the magnitude of the reference star is over reported by two magnitudes, you will have a two magnitude systemic error.

Systemic Errors are extremely difficult to detect. To avoid them, you should ensure that your instrument is correctly calibrated. If you suspect that you might have a significant systemic error then compare your results with a known one. If your mean results constantly appear outside of the uncertainty range of the known results you might well have an unknown systemic error, especially if they constantly appear on the same side of the known result.

A tale of caution. In 2011, the OPERA experiment mistakenly announced that they have observed neutrinos travelling faster than light. It turned out that an ill-fitting fibre optic cable was causing a systemic error and, in fact, the results were entirely consistent with previous experiments.

11.2.2 Random Errors

Random errors (also known as noise) are an inherent and unavoidable aspect of any system. However you should endeavour where possible to minimise random errors by, for example, cooling the CCD, and quantify those you cannot.

In order to reduce errors, your should perform many observations if possible. As you can see from Eq. 11.1 you should be able to see that increasing the sample reduces the standard deviation by $\sqrt{\frac{1}{N}}$. A simple way of determining if your errors are large is to look at the spread of your sample, i.e., the maximum value minus the minimum value minus the mean value, and multiply this by 0.66. This will give you an approximate indication of your 1σ error; although you should beware of large outliers and small samples sizes, which tend not to be Gaussian in distribution.

A common trick used by marketers of beauty products is to say, for example, that 85 % of woman agree that the product has a positive effect. This result seems to suggest that the product works, when, in fact, the sample size is too small for the results to be meaningful.

Many of your errors will take the form of **Shot Noise** which is noise that is a result of the quantised nature of both electrons and photons. Shot noise has a Poisson distribution rather than a Gaussian. The Poisson distribution is characterised by Eq. 11.2,

$$P_{\mu}(\upsilon) = e^{\mu}\frac{\mu^{\upsilon}}{\upsilon!} \tag{11.2}$$

In terms of Astrometry, your random errors are largely caused by the pixel size and the seeing, the Gaussian distribution of the starlight caused by random changes in the atmosphere. There will also be errors from the distortion of the image due to the optics and problems inherent in the processes being used.

You will often hear the term **Signal to Noise Ratio** or SNR in astronomy. The SNR is just the ratio of the part of your signal (or count in the case of optical astronomy) that is coming from your image and that part that is coming from random fluctuation in the system. If you have low signal to noise, your uncertainties will be high and vice versa. Generally speaking an SNR of 100 is good, and 200 is excellent, but this does depend on what you are trying to achieve.

Many of the astronomical applications used for photometry, and some for imaging, will report the signal to noise (or just the noise of the background). You should use these figures with the understanding that they are under-estimating the uncertainty. The measurement report is based on the variation in the Sky Signal. However, many things contribute to errors in observational astronomy, and although the sky signal is the most dominant source of error, it is not complete. Flatfield, dark and bias noise, Interpolation errors and charge transfer errors as well as scintillation noise all, contribute to the overall noise level.

When dealing with spectrography rather than imaging, the problem of noise becomes more important. Although we vertically collapse the spectra so that we have a single line of pixels with each pixel representing a small wavelength range, that pixel includes light from the source and sky lines – that are not uniform. The subtraction of the sky during the reduction process should remove most of this noise leaving just the shot noise. Binning can help to improve the signal to noise. By joining adjacent pixels, we increase the signal while, hopefully, averaging out the noise to some degree as the error scales with $\frac{1}{\sqrt{N}}$. However, as the signal intensity varies across the spectrum so does the signal to noise. Therefore, the signal to noise ratio of the spectra is the mean pixel signal to noise.

11.2.3 Working Out the Error Value

For photometry, the simplest method for determining the error is to use the report signal to noise level. The S/N then becomes an error in magnitude by applying Eq. 11.3, where $\sigma(m)$ is the magnitude error, S is the signal in flux or counts and N is the noise, also in flux or counts.

$$\sigma(m) = \pm 2.5 \log \left(1 + \frac{S}{N}\right) \tag{11.3}$$

Equation 11.3 does not deal with errors that do not in some way link to the background variation. This is best addressed by taking repeat observation (assuming the source is not intrinsically variable over short time scale). Once you have your measurements, by finding the range and then apply Eq. 11.4, where Δx is the error in the sample, x_{max} and x_{min} are the maximum and minimum value in the sample and N is the sample size.

$$\Delta x = \frac{x_{max} - x_{min}}{2\sqrt{N}} \tag{11.4}$$

For large samples Eq. 11.4 is not suitable, and we have to find the mean value of the sample and its standard deviation as per Eq. 11.1 and then apply Eq. 11.5 where Δx is the error in the mean of the sample, σ is the sample standard deviation and N the sample size.

$$\Delta x = \frac{\sigma}{\sqrt{N}} \tag{11.5}$$

11.2.4 Propagation of Uncertainties

It may come as a surprise but when you add two results, you do not add the errors. There are very specific mathematical rules that govern how to treat errors when performing mathematical operations on results. This methodology is known as the Propagation of Uncertainties.

When adding or subtracting two values x and y and their errors Δx and Δy to find z then error in the resultant value Δz is Eq. 11.6.

$$\Delta z = \sqrt{(\Delta x)^2 + (\Delta y)^2} \tag{11.6}$$

For the multiplication of two values the propagated uncertainty is Eq. 11.7.

$$\Delta z = |xy| \sqrt{\left(\frac{\Delta x^2}{x}\right) + \left(\frac{\Delta y^2}{y}\right)} \tag{11.7}$$

For the division of two values, the propagated uncertainty is Eq. 11.8.

$$\Delta z = \left|\frac{x}{y}\right| \sqrt{\left(\frac{\Delta x^2}{x}\right) + \left(\frac{\Delta y^2}{y}\right)} \tag{11.8}$$

When raising a result x to a power n there error in the result z is found by using Eq. 11.9.

$$\Delta z = |n|\, x^{n-1} \Delta x \tag{11.9}$$

When applying a function f to x and y the propagated error is found as per Eq. 11.10.

$$\Delta z = \sqrt{\left(\frac{\delta f}{\delta x}\right) + \left(\frac{\delta f}{\delta y}\right) (\Delta y)^2} \tag{11.10}$$

Chapter 12
Solar Astronomy

Abstract The Sun is the most obvious astronomical object and the nearest star. It is dynamically active with many changing surface features. We discuss Sunspot, flares, prominences and other solar features and, via a practical example, show how to view the Sun in a safe manner.

12.1 Introduction

With a diameter 109 times, the Earth's and containing 99.89 % of the mass of the Solar System, the Sun dominates the Solar System and our daily lives. It provides the energy for almost all biological processes on Earth and its long term, and arguably short term; evolution has a major impact on life and human civilisation.

The Sun's composition is typical of a star of its age, 75 % Hydrogen, 24 % Helium and a smattering of other elements, including carbon, and oxygen. High densities and temperatures within the Sun's core, ($150\,\mathrm{kg\,m^{-3}}$ and $15 \times 10^{15}\,\mathrm{K}$) create the environment for a sustained Thermonuclear reaction. The conversion of hydrogen into helium via the Proton-Proton Chain. The first part of this chain, the creation of deuterium, is the choke point in the process and is sensitive to conditions. When the core is too hot or dense, more deuterium is created, and more energy released. The release of more energy causes a slight expansion in the core, dropping the core temperature and pressure and thereby suppressing the fusion rate. Hence, we have a feedback loop whereby the radiation and thermal pressure within the Sun in matches by the gravitational pressure; the Sun is in Hydrostatic Equilibrium.

The nuclear synthesis of helium from hydrogen results in the liberation of gamma rays and neutrinos. Neutrinos weakly interact with matter, and these are lost, escaping the Sun in a few seconds. The gamma rays, however, are continually absorbed and re-emitted as they pass through the layer above the core, the Radiative zone and are converted into lower energy photons. Eventually, the Radiative zone gives way to the Convective Zone, and energy is now transported to the surface by the convection of hot gas. Once the photosphere, the visible surface of the Sun is reached, the gas temperate as dropped to 5,500 K the energy created in the core is re-emitted once more as optical and ultraviolet light. The typical time taken for a photon to escape the Sun via this process is of the order of a million years.

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The atmosphere of the Sun above the photosphere is made of several layers with many of the events you will observe will occurs within these layers. The coolest layer is that immediately above the photosphere and is about 4,000 K which is cool enough for some molecules exists. It is this region that generates many of the molecular spectral lines we see in the Solar spectrum (see Chap. 13). Above this cool layer is the hot region know as the Chromosphere. This region is the source of many emission lines including the important Hydrogen Alpha line. Above this lies the Transition zone, a region of very rapid temperature rise, which is largely visible in the Ultraviolet. The Transition zone is, in turn surrounded by the Corona, an extremely hot region, up to 20 million K, that can be considered the extended atmosphere of the Sun. The Corona not normally visible to a ground-based observations except during a solar eclipse.

12.2 Solar Features

In the practical section of this Chapter, you will be asked to observe a number Solar features. There can be broken up into two groups, photospheric features and chromospheric features. How you observe them depends on their location and their temperature.

12.2.1 Sunspots

The Sun has a very powerful magnetic field the drives many of the processes that we see on the photosphere. It is believed that the source of this field is the Tachocline, the boundary between the radiative zone and the convective zone. The radiative zone and core rotate as a solid body while the convective zone and the boundaries above rotate differentially. This differential rotation drives a magnetic dynamo that is the source of the Sun's magnetic field.

Magnetic field lines escape the photosphere and loop high into the Sun's atmosphere before descending and re-entering the photosphere. These field lines suppress convection, preventing the flow of hot plasma from below. The reduced convection results in a temperature drop at the photosphere of the order of 500 K. Although the photosphere is still hot, about 5,750 K, the 500 K. The temperature drop is enough for the region of suppressed convection to look dark against the bright photosphere. This dark spot is a Sunspot. Sun spots always appear in pairs (although this is not always noticeable to an observer), reflecting the a North and South poles of the magnetic field lines. The order in which the Sunspots appear, for example, a North spot may be in front of a South Spot in the direction of rotation, in reversed between solar hemispheres.

The Sun goes through a period of increasing magnetic activity, characterised by an increasing number of Sunspots, every 11 years – the Sun Spot Cycle. After a cycle, new starts the new Sunspot slowly reappear with their order reversed.

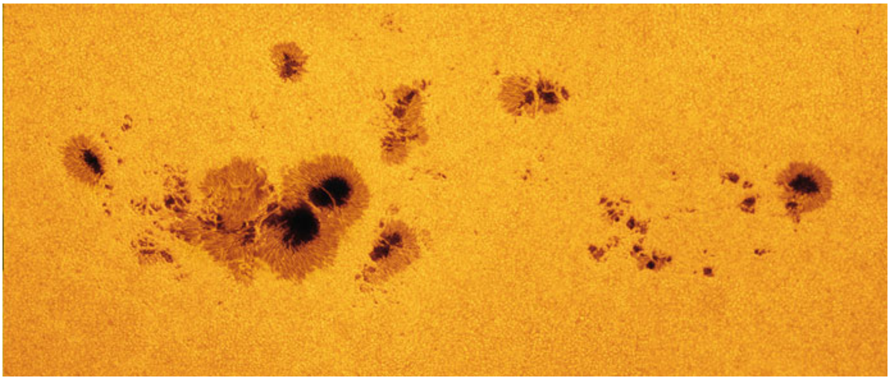

Fig. 12.1 Image of a large Sunspot archipelago captured by Alen Friedman on 10th July 2012

Hence, there is a double cycle made of two 11-year Sunspot cycles. The reasons behind the 11 and 22 cycles are still unknown. However, if you are observing the Sun, it is useful to know where in the sunspot cycle you current are. Sunspots can be seen in a number of filters and are best seen in white light. Figure 12.1 shows a large group of sunspots observed in white light. The very dark area is known as the umbra and the surrounding paler area, the penumbra. Spectrographic observations of sunspots show Zeeman line splitting caused by the intense magnetic field. The similarity of the region surrounding the sunspot to iron filings near a magnet should not be overlooked.

12.2.2 Faculae

Faculae are bright patches located just above the photosphere. They appear in areas of intensive magnetic activity, and it is this activity that is thought to cause their brightness, being 300 K above the surrounding photosphere. Hence, Facula may be considered to be the reverse of a sunspot. Facula are challenging to observe, and such observations are best undertaken near the solar limb. Figure 12.2 shows a Faculae group.

12.2.3 Granules

Granules are small scale structures found on the photosphere and are convection cells of hot material coming up from deeper within the Sun. They are dynamic structures, changing rapidly in comparison to other solar features. The dark edges are the cooler gas descending. Figure 12.3 illustrates a typical granulated solar surface.

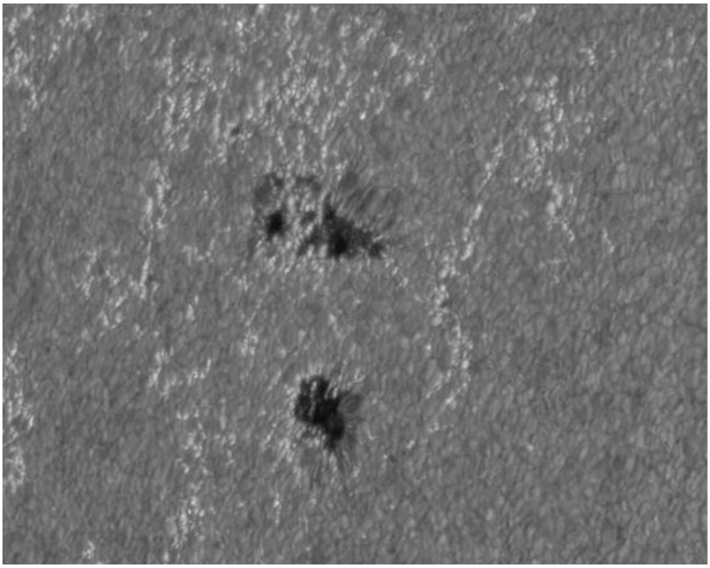

Fig. 12.2 Image of Faculae (*white*) on the surface of the Sun surrounding a sunspot pair (Image NASA)

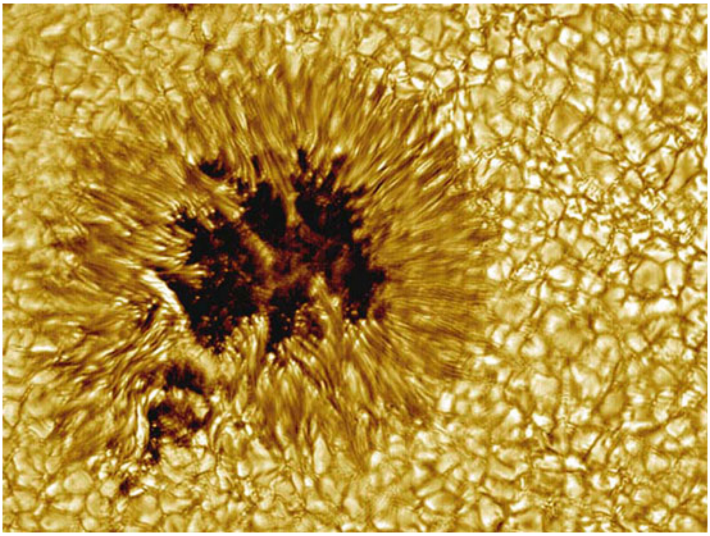

Fig. 12.3 Image of convection granules surrounding a sunspot (Image NASA Robert Nemiroff (MTU) & Jerry Bonnell (USRA))

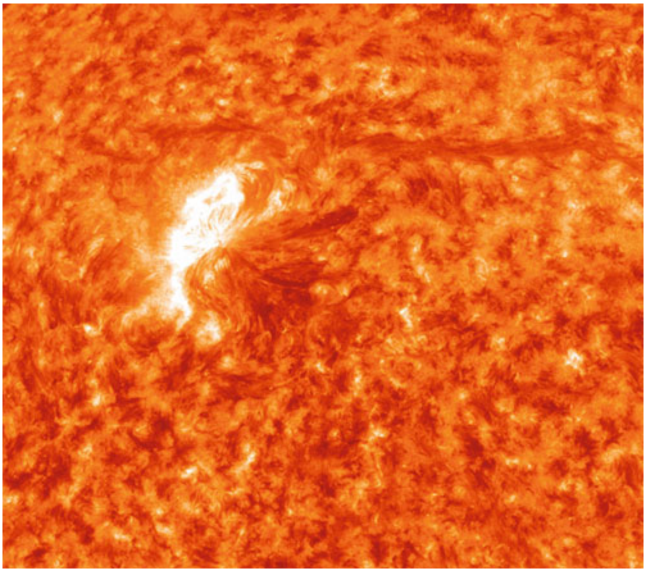

Fig. 12.4 Image of solar Plage (Image Credit: NASA/SDO)

12.2.4 Plage

Plages are bright extended beach like features (hence the name) located above and around sunspots. They are very bright in hydrogen alpha that leads is to the observation that although in white light sunspots darken the Sun in hydrogen alpha the presence of the Plages results in brightening. Figure 12.4 shows a Plage sitting over, and obscuring, a sunspot pair.

12.2.5 Filament

12.2.6 Prominences, Filaments and Flares

Given that the surface of the sun is a plasma it is not surprising that there is an interaction between the magnetic field lines extending from the sunspots and the hot plasma. The plasma becomes trapped in the magnetic loop and enables the observer to view the loop extending high into the chromosphere. When view from the side these loops of trapped gas are known as a Prominence (Fig. 12.5), when views from above they cast a block light from the hotter gas below creating a dark line across the surface of the Sun. In such an orientation, they are known as Filaments (Fig. 12.6). They are however, the same feature. Stresses forms in the field lines,

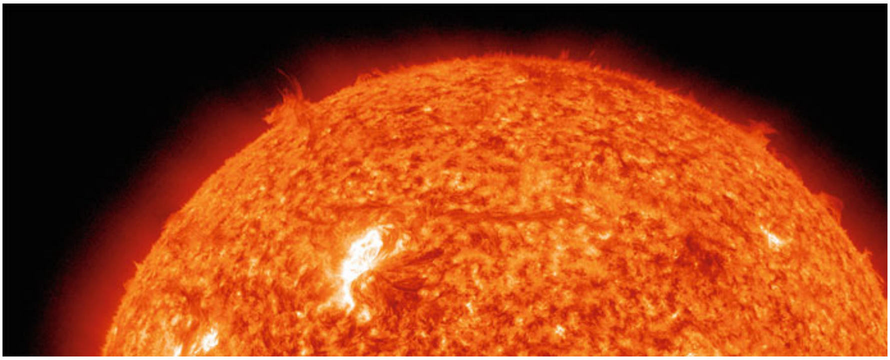

Fig. 12.5 Image of solar prominences (Image Credit: NASA/SDO)

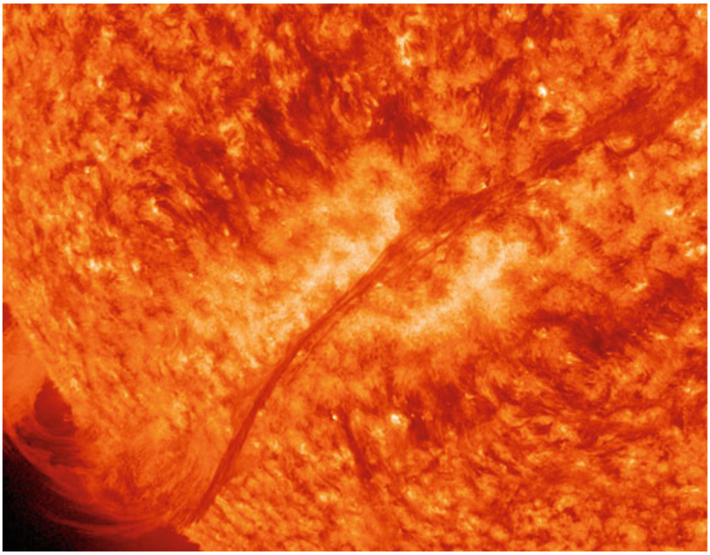

Fig. 12.6 Image of Solar Filament (Image Credit: NASA/SDO)

caused by the loading of the trapped gas. Eventually, the stress on the field lines causes they to break and re-connect onto the surface of the Sun, releasing a very large amount of energy in the process. The energy release is often accompanied by a sudden brightening in the trapped gas – a solar flare. Often a large amount of coronal material may be ejected into space at high speed; a Coronal Mass Ejection. If this material interacts with the Earth's magnetic field, it can lead to bright and extensive aurora.

12.3 Observing the Sun

The Observation of the Sun is one of the most hazardous activities an astronomer can undertake from the point of view of risk to both equipment and self. The Sun should never be observed with the naked eye, nor even with sunglasses.

The safest way to observe the Sun is by projection; either with the use of a pinhole camera or a telescope or a purpose built instrument such as a Sunspotter. These instruments focus the Sun's light onto a white surface, and it is the reflected light of that surface that is observed not the Sun directly. When aligning a solar projection instrument, it is best, as it is with all solar observing to use the telescope's shadow to point at the Sun, rather than looking at the Sun itself. This method is safe, although it is difficult to record the images and shows only limited photospheric details, almost exclusively sunspots.

Solar filters are unlike normal astronomical filters. They are designed to cut out a significant amount of light over a broad range of wavelengths including UV and IR. You should not use traditional, non-solar filters except when using them behind solar rated filters.

Filters are either white light, i.e. they just reduce the amount of light pass through to the observer across the spectrum, or they are narrow line filters. If you suspect a filter is damaged in any way, do not use it; even a pin hole could damage your eye or the camera; this is especially true of film based filters such as produced by Baader. Before using film based filters, I always inspect them for holes using a bright light source, which is not the Sun.

Hershel wedges are becoming increasingly popular for solar observing. The wedge consists of a prism that reflects most of the incoming light away from the observer and absorbs the unwanted UV and IR components. Hence, Hershel wedges are white light devices. However, the should always be used in conjunction with another light reducing filters such as a polarising filter or a neutral density filter as Hershel wedges do not remove enough UV and IR for normal, safe operation without post filtering.

Specialist white light solar filters that mount over the primary aperture are available. There filter the harmful light out before it reaches any optical component. They tend to use a aluminised coating on glass and are an advancement on the mylar film filters as they are less prone to damage. When used with a webcam and a refractor, such filter can be used to make high-resolution images of photospheric activities such as sunspots and granules.

For observations of the chromosphere either a specialist hydrogen alpha telescope, such as manufactured by Lunt or Coronado is needed or hydrogen alpha filters that can be mounted on a suitable refractor. In both cases, there will normally be filtering in front of the first optical component of the telescope and after the last. Often these will both have a degree of polarisation for the amount of light reaching the camera of the eye, to be adjusted. Such filters are ideal for the observation of flares, prominences plages and sunspots.

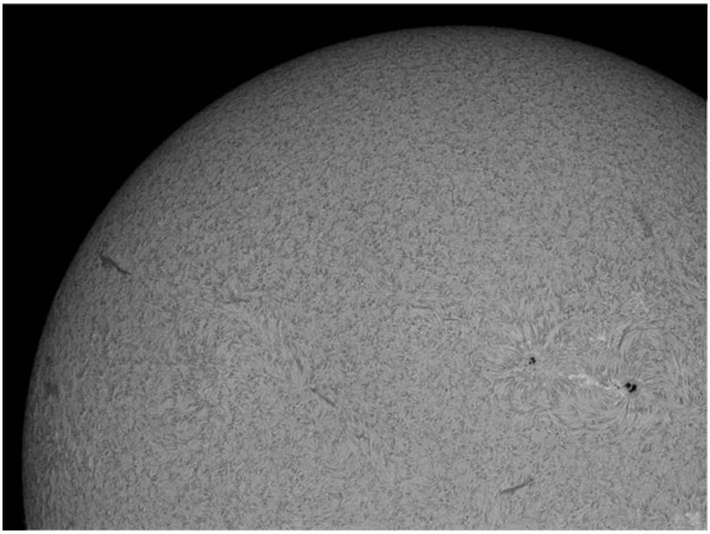

Fig. 12.7 Lucky Image of the Sun taken in Hydrogen alpha showing sunspots and filaments (Image by David Campbell University of Hertfordshire)

An alternative to the hydrogen alpha filter is the 393.4 nm, Calcium K line filter. This isolated light coming slightly lower in the chromosphere and has strong association is magnetic features, such as sunspots and faculae.

Given that you are observing the Sun there are going to be considerable atmospheric turbulence and often solar video files show shimmering or boiling, which reduces resolution. By using a webcam, preferably an astronomical one such as a built DMK or Point Grey (there is no advantage in using a colour camera) one can use lucky imaging to obtain extremely high-resolution images of the Sun such as Fig. 12.7.

12.4 Practical 9: Exploring the Physics of the Sun

The dominance of magnetohydrodynamic forces in the observable layers of the Sun, the Photosphere and the Chromosphere makes the Sun a dynamic and an ever changing target for astronomy. It is also the only star where we can view these forces in any detail. Understanding the nature of surface features and their interaction is a key component of solar physics.

12.4.1 Aims

In this practical you will be observing the Sun over a number of days to determine the evolution of photospheric and chromospheric features, understand limb darkening and determine the Sun's differential rotation rate. In addition, you will learn how to observe the Sun in a safe manner.

12.4.2 Planning

You will need to observe the Sun over a period of approximately 14 days in order to determine the Sun's rotational period. You may have to overcome inherent safety features that prevent the telescope pointing at the Sun. Before observing you should ensure that the correct filters are fitted and that you understand the procedures for solar observing the Sun at the facility you will be using. It would be beneficial if you can observe in both Hydrogen Alpha and White light if at all possible. The authors suggest that you use a webcam as your image device and that you undertake Lucky Imaging in order to improve resolution.

12.4.3 Observing

Ensure that the solar telescope is correctly set up and the required filters are in place and are undamaged. If the solar telescope is co-axially mounted with another telescope ensure that it has its lens cap firmly attached.

Open the dome and point the telescope at the Sun. Many telescope control programs will not let you perform this action. Consequently, you may have to drive the telescope to position manually; please check with your observatory staff if you need to do this, and if so how. Start your webcam and using the shadow of the telescope nudge the telescope until you see the disk of the Sun. At this point, you may need to focus the solar telescope and adjust the webcams exposure time (or frame rate), brightness and gain in order to get a clear image.

Once you are happy with the image, capture approximately 15 s of video. Lucky image this video file using Registax, for example, in order to produce a high-resolution image. Save this image for processing later and then shut down the observatory.

12.4.4 Analysis

Using the images obtained measure the movement rates of sunspots over the face of the Sun over the course of the 14 days.

Compare the images in Hydrogen Alpha to those in white light and monitor how the sunspots and other transient phenomena such as prominences change over the course of the observations.

Most FITS manipulation software can draw a profile, for example, ATP. Use this feature to determine the brightness gradient of the Sun across its surface.

As with previous practicals, you need to write your report up in standard laboratory style. You should discuss the different structure you see and compare how their appearance differs between hydrogen alpha and white light. You should determine the rotation rate of the Sun and comment how, or if, this changes with latitude. The brightness gradient over the surface of the Sun is known as Limb Darkening. You should investigate what is causing this effect and it is time variable.

Chapter 13
Spectrography

Abstract In this chapter, we introduce the concept of spectrography. Building on the previous chapters on the nature of light and astronomical imaging we discuss the spectrograph and show through a simple practical how to take, reduce and interpret our first astronomical spectrum

13.1 Introduction

Spectrography is one of the principle methods that astronomers investigate the Universe. The existence of dark absorption lines and bright emission lines set against the broad continuum emission and their unique association with energy changes; as discussed in Sect. 2.2, has enabled astronomers to determine:-

- From the shift of a line relative to its normal position we can determine velocity and Redshift and cosmological distance as well as inferring the existence of Exoplanets;
- From the width and strength of the lines we can determine relative abundances of elements and molecules as well as the temperature of the body;
- From the shape of the line, we can determine the pressures, density, rotation characteristics of the body and infer the presence and strength of its magnetic field;
- The measurement of the continuum part of the spectrum, that part without lines, enables us to observe non-quantised emission processes and to determine the Spectral Energy Distribution of the body;

Hence, spectrography provides us with a set of tools that allow us to determine a wide range of physical characteristics.

13.2 The Spectrograph

In order to obtain spectra, we must disperse the light from the observed object such that photons of the same wavelength fall in the same location in our imaging device. There are two ways of doing this, by refraction using a Prism or by diffraction using

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a diffraction grating. There is a third device that uses both methods known as a Grism. The Grism is a prism with a grating etched upon its surface.

It is also possible to undertake limit spectrography by using narrow band filters. As these are tuned to a single spectral line, such as [OIII] or [SII], it is possible to measure the intensity of this line, although not the width.

The diffraction grating is a reflective or transmissive surface etched with regular lines, often over a thousand lines a millimetre. The lines in effect turn the source into multiple sources, one per line. These constructively and destructively interfere creating multiple rainbow patterns on the detector.

Equation 13.1 shows the diffraction equation, where d is the line spacing in the grating, θ is the emergence angle, n is an integer known as the order and λ is the wavelength.

$$d(\sin\ \theta) = n\lambda \tag{13.1}$$

As we can see when n = 0 the dispersion angle, θ becomes zero. The 0th order image, therefore, is not spectra per say more the image of the object. Either side of the object is a series of repeated spectra, becoming fainter as we move away from the 0th order. This first of these spectra, either side of the 0th order are the first order spectra that are typically the spectra that would be analysed.

For reflection gratings Eq. 13.1 becomes...

$$d(\sin\ \theta_i + \sin\ \theta_e) = n\lambda \tag{13.2}$$

where the term $\sin\ \theta$ is replaced by the sum of the sines of the Incident angle θ_i and the angle of refraction θ_e.

Figure 13.1 shows the layout of a traditional long-slit spectrometer as found at many small observatories. Light enters the spectrograph through a slit that is designed to both improve the spectral resolution and to isolate the source. Often the size of the slit can be adjusted. Once inside the spectrography the light is collimated so that the incident angle is consistent. The reflection grating disperses the light depending on wavelength. Typically the grating can be rotated so that the area of the spectrum of interest will fall on the centre of the CCD. Likewise, some grating are double backed with the option of a high or low-resolution grating. The dispersed light passes through the camera's optical system and forms a spectrum on the CCD. Most spectrographs will have a slit illuminator to allow the alignment of the source and slit. They may also have a calibrated low-intensity light source such as a small argon or neon bulb. These will produce a strong reference line in the spectra for calibration purposes. Of course, if you are observing in a light polluted area, the sodium doublet lines are ideal for calibration.

The spectral resolution of you spectrograph depends on the level of dispersion and the pixel size of the CCD in addition to the grating and the order being observed. Equation 13.3 determines the spectrograph's spectral resolution $\Delta\lambda$ from the grating width d, the angle of refraction θ_e, the order number n, the pixel scale x and the focal length of the instrument F.

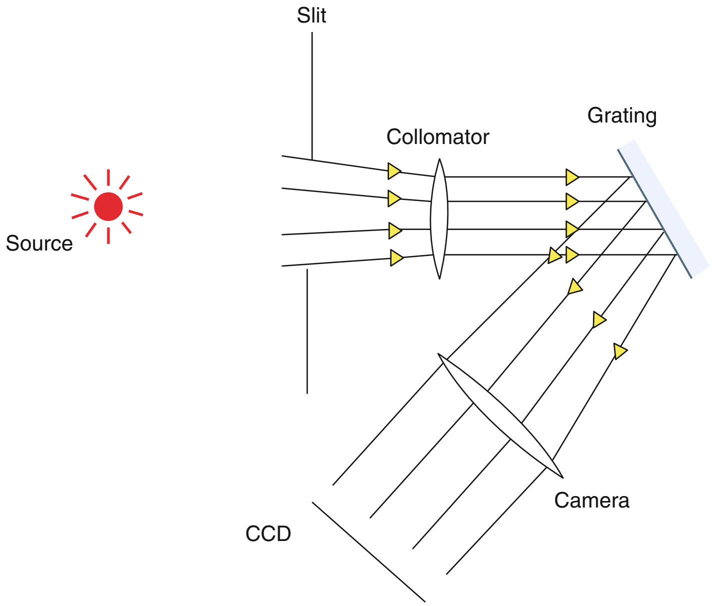

Fig. 13.1 Diagram of a long-slit spectrometer showing the key components of the slit, grating and detector. It is common for there to be a number of additional optical components inside of spectrometers that improve image quality

$$\Delta\lambda = \frac{d \cos \theta_e \Delta x}{nF} \tag{13.3}$$

The ratio between the central wavelength λ and the wavelength covered by each pixel is the spectrograph's resolution, R. Hence, larger R indicates high resolution. For example the commercially available Skelyah Lhires III has an R of 18,000 which equates to a resolution on 0.035 nm near the Hydrogen alpha line, which equates to a Doppler shift of 16 kms^{-1}; good enough to observe the rotation of stars.

$$\frac{\lambda}{\Delta\lambda} = R \tag{13.4}$$

With the light dispersed over multiple orders and light lost in both reflection and refraction off of optical surfaces. It should come as no surprise that a considerable proportion of the light entering the spectrograph does not get detected. A good surface will lose about 2 % of the light reflecting or passing through it while the grating may lose 75 % and the CCD will be of the order of 80 % efficient. Hence, in a typical eight element set up, total throughput will be of the order of $0.98^8 \times 0.75 \times 0.8 = 0.51$. So we lose about half our light within the optics.

It is important, therefore, when undertaking photometry that the exposure time in long enough to get a good signal to noise and that the slit in opened to the optimal aperture.

As with imaging, there is always the option of binning to improve the signal to noise ratio. As individual columns of pixels all represent the same wavelength, assuming that the image is not rotated, we can, and certainly should, vertically bin our spectra to a single row. If we have higher spectral resolution than we need we also have the option to bin linearly after binning the columns. Binning is good practice if you are not particularly interested in the line width, it will also enhance the bright lines when the spectra are plotted and reduce confusion within the plot.

Just as it is possible to undertake field photometry, it is possible to undertake field spectrography but not with a standard long-slit spectrograph as shown above. Multi-slit photometry is possible by using a modified long-slit spectrograph. However, there are problems with this approach, and it can be inefficient.

Fiber-Fed Multi-Object Spectrographs (MOS) use a mounting plate with holes drilled where the image of the stars would fall. Integral Field Units (IFU) divides the field into a grid with a small lens covering each cell. The lenses focus light into fibre optic cables glued on to the back of the lens. The fibres feed into the spectrograph and forms multiple spectra. IFU are powerful tools for determining large-scale movement such as the rotation of galaxies.

It is possible to undertake slit-less spectrography by using an inline transmission grating such as a Paton Hawksley Star Analyser. Inline transmission gratings can be fitted to the nose of a camera, webcam or inserted into a filter wheel. They produce low resolution spectra of all the objects in the field, which may result in confusion and overlapping spectra. However, they are inexpensive, easy to use and are becoming increasingly popular with amateur astronomers.

13.3 Taking and Reducing Spectra

To some degree, the taking of spectra and reducing of spectra is very similar to that of normal imaging. Although the software used for the reduction is different, in general, the camera control software is the same.

For each spectrum, a reference and a calibration spectra are required; this is in addition to bias dark and flat frames.

The calibration will be of a bright source, with a known, standard spectra and at the same air mass as the source; where possible it should be of the spectral type and the same exposure time. If the primary source being imaged has a known standard spectrum, there is no need for a calibration frame.

The Reference Frame is a spectrum of a known source, perhaps an arc lamp or an LED, either within the spectrograph itself or external. The Reference Frame is used to convert pixel position to wavelength.

Once the spectra are taken, they should be turned into science frames by the subtraction of the bias and dark frames and the division by the flat frame. The spectra

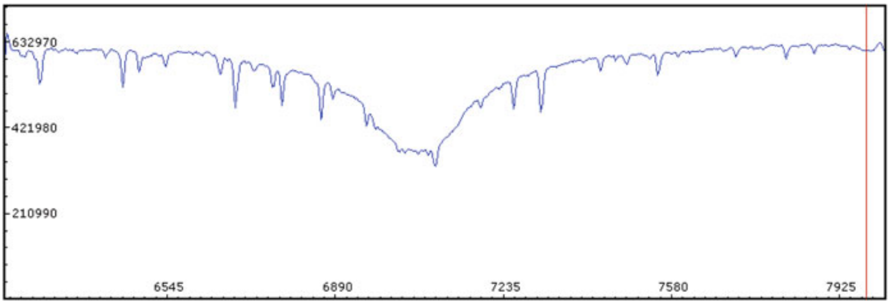

Fig. 13.2 Stellar line spectrum. Referenced but not calibrated for instrument spectral response. X axis is in ångströms, y axis is in counts

should be located within the frames and cropped out. A region of the sky should also be cropped out parallel to the spectra, with the same size and the same starting column. These Sky regions contain unwanted lines from the atmosphere (as do the source spectra), subtracting them from the source removes these unwanted lines.

The Reference spectra (which have lines of known wavelength) are used to correlate pixel position to wavelength and the wavelength scale. When these are applied to the Source and Calibration spectra, they convert to line spectra plotting intensity against wavelength. This should appear somewhat like the stellar spectrum shown in Fig. 13.2.

Accounting for spectral response is more involved. The calibration spectra need to be divided by the standard for that object. This process leaves a spectral response curve for your instrument set up. It will, however, be messy, due to the spectral lines; something smoother is needed, which can be achieved by fitting a polynomial; the instrument response curve. Inverting this curve and multiplying it by the source line spectra we obtain a calibrated spectra.

Normal practice is to normalise the spectra by identifying the highest peak and dividing the entire spectra by peak value.

Once a spectrum is reduced spectral lines may be identified from their position along the curve.

13.4 Practical: 10 Methane in the Atmosphere of Neptune

The Solar system bodies Uranus and Neptune contain a significant amount of methane in their atmosphere. This methane is sufficient to be detected by a small instrument with a spectrograph.

13.4.1 Aims

You will be taking spectra of one either Uranus and Neptune to identify key components in their atmospheres. You will learn how to operate a telescope fitted with a spectrograph and how to take, reduce and interpret spectra.

13.4.2 Preparation

Before arriving on the site, you should have selected a suitable target and calibration star. You should have produced finder charts for the night and have the source the standard spectra. Uranus and Neptune are large enough to have a visible disk that should make identification straight forward.

13.4.3 Imaging

On arrival open the dome and cool the camera(s) down. If needed take flats or obtain suitable library flats.

Once down to operating temperature slew the telescope to a bright star and do a focusing check. If you have a flip mirror or off-axis instrument, use this to point the telescope and confirm that it is aligned and point at the right piece of sky. When you focus up the spectrograph, you may have to open the slit wide.

Once you are aligned and focused, move the telescope to the source. Align the target in the slit, which should now be closed up to a suitable width. Image the target, with binning if needed, making sure that all pixel stay within the linear range of the CCD.

Once you have suitable spectra, take dark and bias frames and then turn on the reference source and take another spectrum. Ensure you can see the lamp lines in the image.

Once you have your reference, slew to your calibration source and align in the slit. Take the spectra, again while remaining in the linear range, follow this with dark and bias frames. Then take spectra with the reference lamp on.

Save all you files to a transportable medium and check that they are in an acceptable condition; for example, they have no plane trails, and once you are happy with your work shut the observatory down.

13.4.4 Reducation

Once you have the spectra, you have to reduce them in order to get science line spectra:-

- Using your favorite reduction package, remove the bias and dark frames from the spectra and divide by the flat.
- Rotate the frames so that the spectra are flat and not tilted. If the spectra are curved, remove this.
- Crop the spectrum and save to file. Perform the same action to the sky area, ensuring that the cropped size is the same as the spectra and that it starts at the some pixel column.
- Subtract the sky from the source to produce a sky removed spectrum.
- Load the reference spectra into a spectral reduction package such as IRIS, Vspec or Rspec and bin them, resulting in a line plot of intensity v pixel number.
- Locate the emission lines in your reference spectra noting the pixel value for each line and its corresponding wavelength. Use this to find the Linear dispersal of the spectrograph.
- Load the calibration spectra and calibrate it using the pixel location of the wavelengths found in the reference spectra and the calculated Linear dispersal. Applying this should create a line plot of intensity v wavelength.
- Load the standard spectra and bin it down to a line spectra if it is not already, and divided the calibration spectra by the standard spectra.
- Fit a polynomial to the resulting spectra. It may be necessary to remove some spectral lines to get a smooth fit. Save this plot.
- Load the source spectra and bin it. Calibrate this spectrum using the data from the reference spectra, as previously.
- Take the calibrated source spectra and divide it by the polynomial you fitted to the calibration spectrum.
- Normalise your spectrum by finding the value of the highest peak and dividing through by that value.

13.4.5 Anaylsis

Using a spectral line catalogue identify and label the spectral line within your spectrum. Pay particular attention to lines redder than 600 nm as these may be molecular lines. Additionally compare the solar spectrum to your spectrum.

Write you observations and reduction up in the standard format. You should discuss how the spectrum from you source compares to the Solar spectrum and what this implies. If you did see molecular lines, you should discuss what this suggests about the atmospheric content of the planet's atmosphere and temperature.

Chapter 14
Radio Astronomy

Abstract Radio astronomy is becoming increasing important in astrophysics, with increasingly complex instruments, sensitive and large instruments coming online. In the chapter, we introduce some of the concepts in radio telescope design. As teaching radio telescopes are still rare, we introduce radio astronomy with three practicals. Mapping the Milky Way, Detecting Meteors and observing Jupiter. The last two of which can be undertaken with low-cost radio set-ups.

14.1 Introduction

In the early 1930s, Karl Jansky (whom the Jansky is named after) was working on improving shortwave radio reception. Using a directional antenna, he noticed a single that appear and disappeared over a 24 h cycle. Using his equipment as a primitive radio telescope he was able to determine that the source of the signal was the not Sun (although the Sun is a bright radio source) as he previously suspected but a source in Sagittarius. We now know that this source is the supermassive black hole and radio source, located at the centre of the Milky Way, known as Sgr A. So started the age of radio astronomy, and the number of known astronomical radio source rapidly increased.

In the early 1960s two physics Arno Penzias and Robert Wilson where working on the Bell Lab's Horn antenna in New Jersey (interestingly Jansky was also working at Bell Labs when he discovered Sgr A). Penzias and Wilson were working on cryogenically cooled microwave receivers and were troubled by a continual background signal being received. The single was coming from all parts of the sky and was uniform in intensity. After a complete examination of the entire instrument, including removing a white dielectric deposit from within the horn (i.e., bird droppings) they concluded the signal was real.

The physicists contacted astrophysicist Robert Dicke thinking he might be able to help with the source of the mysterious signal. Dicke was, as it turns out and inspired choice. He was working on the theory of the evolution of the Universe, the Big Bang, and the prediction that the should be a remnant signal from the age of de-ionisation, seen as an isotropic microwave signal. Dicke even went as far as

M. Gallaway, *An Introduction to Observational Astrophysics*, Undergraduate Lecture Notes in Physics, DOI 10.1007/978-3-319-23377-2_14

suggesting that the Bell Lab Horn antenna mich be the perfect instrument to detect the signal. This was the very signal detected by Penzias and Wilson – The Cosmic Microwave Background (CMB).

Penzias and Wilson were awarded the Nobel Prize in Physics 1978 for the discovery of the CMB. When you look at a de-tuned terrestrial television about 1 % of the signal you see is from the CMB. You are seeing the birth cry of the Universe on your television.

A few years after Penzias and Wilson's discovery a young graduate astronomer, Jocelyn Bell Burnell was working on a dipole radio telescope in Cambridge, known as the hop field, due to its similarity to the supports used for growing hops.

Bell Burnell noticed a strange signal in her data. It was astronomical, strong and varying very rapidly and with a very consistent period. Initial fears that the signal was of alien construct were soon put aside when it was realised that Bell Burnell had detected the first pulsar. A dense rotating neutron star emitting two tight beams of radio waves that briefly become visible to the observer as the neutron star rotates. Very much in the manner of a lighthouse. Neutron stars are formed in supernovae explosions, and this object was located in the heart of the Crab Nebula.

Bell Burnell's doctoral supervisors were awarded the Nobel Prize in Physics for Bell Burnell's discovery, Bell Burnell was not, somewhat controversially. Bell Burnell is on record as stating that she believes that this was the right choice and that her supervisors were fundamentally responsible for the project. However, Bell Burnell has gone on to be a very successful and well-respected astrophysicist and is a strong advocate for the involvement of women in the science.

Radio astronomy has become increasing attractive to astrophysicists as it can operate in daylight and most weathers (depending on frequency), is mostly unaffected by interstellar reddening and can observe both high and low energy physics that are not accessible by conventional optical and infrared observations.

The principle limitations of radio astronomer are its general insensitivity and poor resolution. However, the ability to join an array radio telescopes together to create a single large virtual telescope and to synthesis a telescope with a diameter equal to the distance between telescopes in the array with the highest separation (interferometry) have overcome these obstacles.

In the coming years, three very large interferometers will reach full operations. Both ALMA, very large a submillimeter interferometer and LOFAR very large a long wave interferometer are now operating. Work is about to start in South Africa and Australia, on the Square Kilometer Array, and its completion will usher in a new era of radio astronomy.

14.2 Instrument Background

A radio telescope consists of two components, the antennas that receive the radio signal and the turns it into an oscillating electrical signal, with the frequency of oscillation been the same as that of the signal. This change occurs due to the electromagnetic characteristics of light, the magnetic field inducing a current in the conducting antenna.

This current propagates down co-axial cabling, the transmission line, to the receiver that measures the voltage being received, which is proportional to the signal strength can be taken as how load the signal is. Often within the transmission line there will be one or mode Low Noise Amplifiers, designed to enhance the signal on its way to the receiver.

All antennae have five defining properties; Impedance, Directional characteristics, Foward Gain, Polarisation and Beam Pattern.

The Impedance is the measure of the opposition that the antennae present to the induced current. Impedance should be matched throughout the system as mismatching can cause internal reflections and signal loss.

Some antennae, such as the one on your wi-fi router, will be omnidirectional. Although some radio telescopes use omnidirectional antennae, for example, LOFAR, it is generally not a desirable property of a radio telescope antenna. Hence, most radio telescope antenna are designed only to receive radio waves from one direction – as much as possible.

The forward gain is just how efficient the antenna is when pointed to a source when compared to a standard dipole antenna. Measured in *dB*. A 3 dB forward gain means your antenna is twice as sensitive, 6 dB would be four times as sensitive. Similarly, the backwards gain is the amount of signal that is lost as it leaves the antenna. Basically when talking about forward gain, the bigger, the better.

As we have previously discussed electromagnetic waves consists of an electric wave at right angles to a magnetic wave. If the orientation of the electric field is always the same for all photons from a source, the source is polarised. In effect, most astronomical sources have only limited levels of polarisation, although most antennae are polarised. Hence, half the photons are lost. Although this feature can be used to block unwanted terrestrial sources.

The Beam Pattern is the spread of radiation caused if we used our antenna to transmit rather than receive. Do not worry is this seems an odd concept trust me the maths works. We find there is a main beam extending out from the antenna related to its field of view. However, there will be some other beams, known as sidelobes extending as pairs either side of the main beam. You should be aware that you can detect a source in a side lobe, but the side lobe will have lower sensitivity.

The most common antenna is the dipole, which consists of two lengths of wire or tube, or rods that are linked to the radio receiver via the transmission line. The most common dipole is the half wave, which, as its name suggest has antennae half the length of the wavelength we wish to receive. Dipole antenna has a doughnut shaped Beam Pattern and will therefore in the plane it is mounted.

The Yagi-Uda antenna, now more commonly know as just a Yagi, is a multi-dipole antenna, again normally half wavelength. Yagi antennae can be optimised to balance the competing forces of gain v band width. Hence in domestic applications, such as receiving television, which required wide bandwidth Yagi antennae tend to have reflectors to improve gain. As anybody who has tried to mount a new television antenna will know, a Yagi tends to have a small Beam Pattern, which is good for astronomical applications. Hence, the Yagi antenna is often used in low-cost systems.

At the more professional end of radio astronomy, we come to the parabolic reflector or just dish. These are satellite receiving dishes scaled up. In fact, domestic Sky Television dishes have become popular amateur radio astronomy antenna. Traditional, big dish, radio telescope, such as the 4.5 m RW Forrest teaching telescope at the Bayfordbury Observatory (Fig. 14.1), consist of a large dish designed to focus radio waves to a point. At this point will be a simple, quarter wavelength waveguide (just a conductive tube in effect) designed to direct the signal to the receiver. The waveguide is surrounded by the feedhorn (or can) which is a metal can open at one end that guides the incoming radiation onto the waveguide and protects it from stray radiation.

The receiver attached to you radio telescope will have an operating frequency range and a centre frequency. Your receiver, if it designed for radio astronomy should be able to perform two functions. The first is a continuum plot. In this mode, it will just display the variation in source intensity over time. It should also be capable of spectrography. In this mode, it moves the frequency that it is listening in small steps plotting the intensity if the signal against wavelength or frequency. In many cases, it is possible to adjust the integration time, i.e., how long is spent at each wavelength and the step size, the range over which it will be listening to each point along the scan. In effect increasing the step, size is like binning with CCD. Your receiver will most likely report the signal strength in volts. With some difficulty, this can be turned into instrument flux then standard flux. However, just as with counts with CCD, this conversion is not always required.

Radio sources can be divided into two basic classes. Low and high energy. Lower energy sources are quantised source with low energy levels. Typical of this is the 21 cm line. This is a strong emission line for monoatomic hydrogen caused by the change in quantum spin state of the proton and electron. Monoatomic hydrogen (HI) is the most common form of hydrogen and hence is a vital tool for the identification and mapping of galactic structure. Other examples of radio spectra lines are ammonia, water, methanol and carbon monoxide. All these lines occur at low temperatures and low energies. High-energy sources include the Sun, Jupiter, Black holes, active galaxies, pulsars and supernova remnants. In these cases, strong magnetic fields accelerate charged particles in a spiral causing the emission of breaking radiation with a characteristic spectrum related to the Larmor radius. Such sources are often hot and are also X-ray emitters hence the odd effect that x-ray astronomers also tend to be radio astronomers.

The bandwidth limitations of your receiver and antenna setup will determine what you will be able to observe with the instrument. The 21 cm line is one of several regions of the radio spectrum that are protected. This is a strong emission line for monoatomic hydrogen caused by the change in quantum spin state of the proton and electron. Monoatomic hydrogen (HI) is the most common form of hydrogen and hence is a vital tool in the identification and mapping of galactic structure. However, you will not be able to observe pulsars using a telescope.

Fig. 14.1 The 4.5 m R.W Forrest teaching radio telescope located at the Bayfordbury Observatory UK (Image courtesy of the University of Hertfordshire)

14.3 Radio Practicals

The following radio practicals are dependent on the correct equipment. The first of these, Mapping the Milky Way, requires a telescope capable or receiving 21 cm emission. The second, Counting Meteors, requires a simple set up of a yagi antenna, an FM receiver and a copy of Spectrum Lab would also be very useful.

The last practical requires a RadioJove set up, or something similar that can pick up emissions at 8–40 MHz from a point-like a source. Constructing such an instrument in a good practical in itself.

14.3.1 Mapping the Milky Way at 21 cm

The presence of dust and gas in the plane of the galaxy makes it impossible to map our galaxy globally the optical. The 21 cm line emission from monoatomic hydrogen can, however, be used. From the Doppler shift of the and intensity of the line, we can determine the kinematics and size of the hydrogen clouds.

Figure 14.2 shows our galaxy from above the disc, with the spiral arms, marked. As we observe through the galactic plane any line of sight may intersect hydrogen clouds at different distances from the Galactic Centre. We determine these distances using the cloud velocities form as derived from the shift in their 21 cm line and using model of the rotation of the clouds around the Galactic centre.

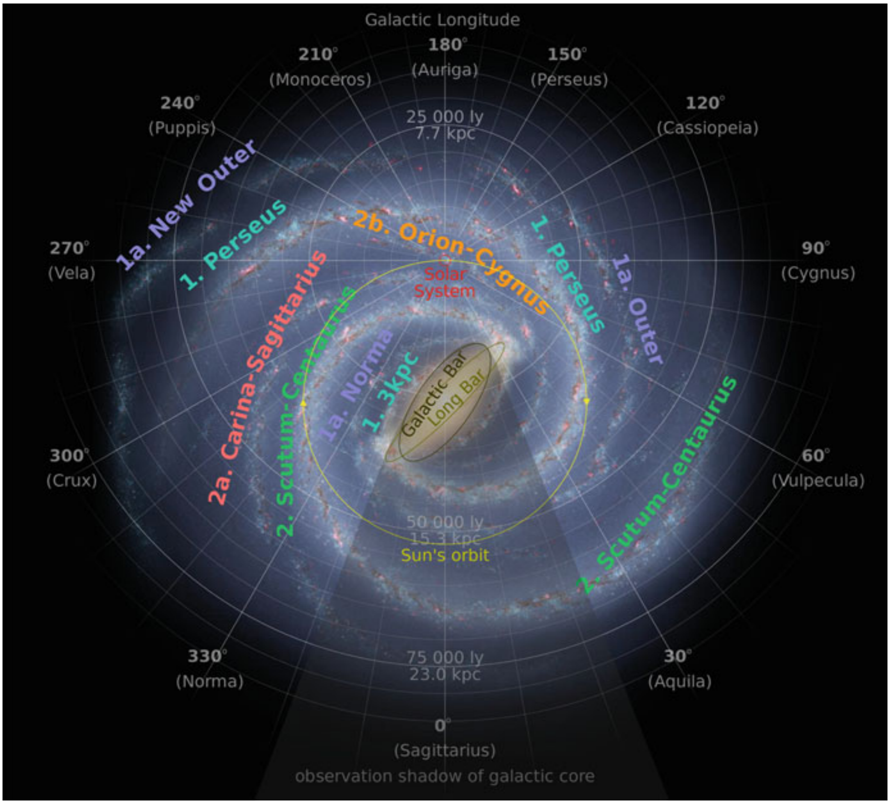

Fig. 14.2 Image of the Milky Ways as derived from observations (Courtesy NASA and R. Hurt)

You should look at your planetarium package with the Milky Way turned on, to determine what parts of the Galaxy you will be able to see and where in the sky it is in galactic coordinates. As the telescope only effectively has one pixel, you will be scanning the Milky Way by taking spectra and them moving the telescope half a beam width before taking the next spectra. This will be needed for the whole visible region, so best to plan how to do this in advance.

14.3.1.1 Aims

The principle aim of the practical to go give the user experience in using a radio telescope and taking data. However, the observer will also be determining the rotation curve of the Milky Way and mapping the HI distribution.

14.3.1.2 Observation

Once at the telescope start it up and slew to a bright radio source, for example, Cas A or the Sun. Once there take spectra with the radio telescope and confirm that you are seeing a signal. You might have to adjust the gain or integration time to see detail in the spectra.

Once you are happy with the operation of the telescope, slew to the first region of the Milky Way. Once on target take the spectra. Note the velocity position of the peak in the curve and its intensity as well as its position in the sky in Galactic Coordinates. If there is more than one peak you might be seeing two or more clouds, at differing distances, in which case note them all. Save the plot to a suitable media.

Slew the telescope half a beam width to the next region of the Milky Way you are going to examine and take other spectra. Do this until you have covered at least a 20° section of the Galaxy.

You might wish to take multiple spectra of each section if you have time as this will help identify your uncertainties.

At the end of the observation, run do not forget to park and shut-down the telescope.

14.3.1.3 Analysis and Write Up

We now wish to map the HI in the Galaxy and find the Galaxy's HI rotation curve. The Local Standard of Rest is the origin of our reference system, i.e. the current position of the Earth compared to the galactic centre at any one time. However, The Earth in moving in relation the Local Standard of Rest as it orbits the Galactic Centre. Hence, any Doppler sifted signal we receive from an HI cloud contains this movement at well as that of the cloud around the Galactic Centre. Figure 14.3 shows the relationship between the Galactic Centre, GC, radial velocity of the Sun relative to the Local Standard of Rest V_o, the distance from the Sun to the Galactic Centre R_0, the distance of the Cloud from the Galactic Centre R_L and the cloud and the cloud's radial velocity relative to the Galactic Centre.

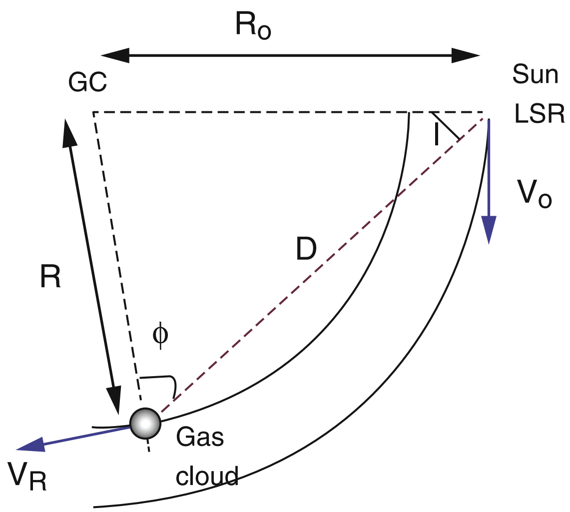

Fig. 14.3 Illustration of the Sun movement around the Galactic Centre, the position of an HI cloud and Local Standard of Rest (Image courtesy of the University of Hertfordshire)

R_0 is known from observation to be 8.5 kpc and V_o to be 232 km^{-1}. The line of sight velocity of the cloud V, as measured by your radio telescope is given by

$$V = \left(V_R \frac{R_o}{R} - V_o \right) \sin(l) \tag{14.1}$$

where l is the galactic longitude. We assume that latitude has little or no influence.

We can see from Eq. 14.1 that for any given line of sight, where l is fixed, tV depends only on $\left(V_R \frac{R_o}{R}\right)$. If the rotation curve is flat or decreases with radius the maximum value for V should be achieved when R is at its smallest value and this occurs at the point where the angle between R and the line of sight is 90°. Hence, the maximum line of sight velocity $V_m(l)$ is given by

$$V_M(l) = V_c(R_L) - V_o \sin(l) \tag{14.2}$$

although this only applies to $-90° < l < +90°$.

We now only need to find R that can be done by using

$$R = \frac{V_R R_o}{\frac{V}{sin(l)} + V_o} \tag{14.3}$$

For each spectra using the largest line of sight velocity calculate R_L/R_o and $V_c(R_L)$ then make a plot of $V_c(R_L)$ against R_L/R_o remembering to use error bars.

Using this rotation curve you can now find the distance to all the clouds you observed by solving the quadratic $R^2 = D^2 + R_o^2 + 2DR_o \cos(l)$.

Using the distance to each cloud and its galactic longitude you need to make a map of the Milky way.

Right your report up in the standard style including as usual all you data and calculations (and errors) as well as the all important rotation curve and map.

If the clouds orbit around the Galactic Centre is Keplerian as we would expect V_c would scale by the square root of R. Is this what you see in your and it isn't what do you think the implications of this are? You should discuss this in your write up.

Your map should show the HI structure of the Milky Way. You should discuss in your write up how your map compares to Fig. 14.2 and do you think are the implications of any difference.

14.3.2 Radio Meteor Counting

There are a vast abundance of meteoroid in the Solar System, most are the size of a grain of salt, some the size of house bricks and a very few are as big as a car of larger.

When they enter the Earth's atmosphere, as about million kilograms do a year, they ionize the atmosphere as they decelerate down from about 20 kms^{-1}. This creates a bright streak of light in the sky at an altitude of about 80 km – a meteor. A few of these objects survive the fiery passage through the atmosphere and are found as meteorites.

There is a period pattern in the appearance of meteors throughout the year with periods with few meteors and periods with many, sometimes hundreds an hour, an event known as a meteor shower. Showers are associated with comets. Comets leave a trail of dust as they orbit the Sun, and the showers occur then the Earth's orbit intersects the orbit of the comet. Meteors not associated with a shower are known as sporadic.

Of course Meteors happen day and night and in all weathers so most go unnoticed from everyday observers. However, there is a relative inexpensive and easy way to detect them at all times and in all weathers.

To key to this is the ionization of the atmosphere by the incoming meteor. Ionized gas is reflective to short wavelength radio waves. So by placing such a radio source over the horizon and pointing and antenna and suitable receiver at the source we find that we can detect the ionization trail. Normally we receiver no signal from the source, it is over the horizon, however when a meteor event occurs between the source and the observer the ionized air acts as a mirror that allows us, very briefly, to see the source. Hence, we hear a meteor in the receiver as a shot blimp, normally Doppler as the meteor moves in the radial direction. This bleep often sounds like a woohoo from a fairground ride ghost train.

Fortunately, we don't have to create our own sources. VHF television masts and military space surveillance radars are very suitable sources, although, of cause, this was not the designers intention (Fig. 14.4).

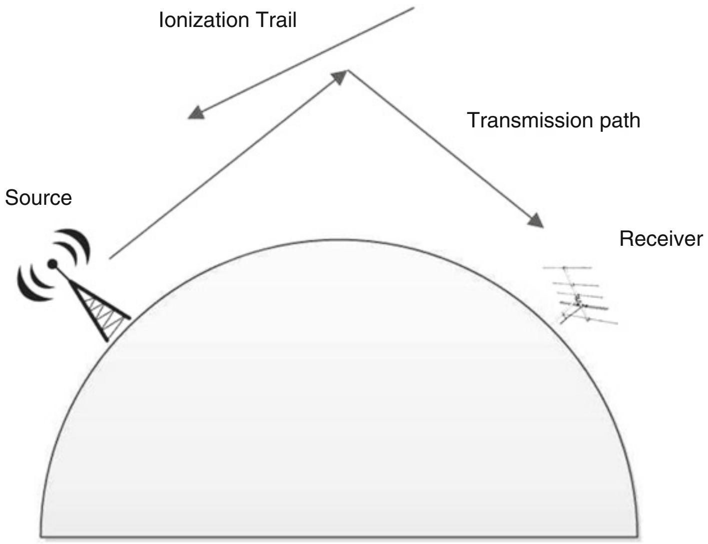

Fig. 14.4 Illustration of the detection of meteors using radio reflection

14.3.2.1 Preparation

You will need to Observe over a period of 24 h. You can either set up the system to record the process or you can takes shifts. You will need to identify the transmitter you are using, its frequency and location relative to the Observatory. You will also have to confirm that there is a meteor shower at the time you are Observing and where the radiant is. If you are using a program like Spectrum Lab to record the data you should ensure you are familiar with its operation.

14.3.2.2 Observations

On arrive you will to set up for observations if this has not be done already. This means pointing the antenna in the correct direction and tuning the receiver to the right frequency. If you are using Spectrum Lab you will need to insure that the sound output of the receiver is plugged into the line-in port sound card of the computer. Check to make sure you are picking up meteor reflection. If you are not using Spectrum Lab or something similar we suggest that the sound is recorded for backup.

Over a period of 24 h note down every meteor trace and the time it occurred. Long notes or streaks then to be satellite traces (or once for the author a bad alternator in security's car!) and can be ignored. You should note ever 15 min the number of detections in that 15 min window.

14.3.2.3 Analysis and Write Up

Plot the 15 min count against time. The curve should not be flat. Discuss what you think this means and the physical process that could cause this. What do you think the implication of the radiant on the completeness of your detections and how might it be improved? What are your sources of error and how significant do you think these are?

Write up your report in standard format including your plot and data.

14.3.3 Radio Emission from Jupiter

In 1955 burst radio, emission was detected from Jupiter at 22 MHz, the first planet to be detected in the radio. Part of this emission is synchrotron radiation from particles trapped in Jupiter's rather powerful magnetic field. This component is at 178–5,000 MHz.

A second radio component at 3.5–39.5 MHz was seen and appears to be strongly related to the orbit of Jupiter's inner moon Io. It is thought that Io and Jupiter are liked by a Flux Tube, a series of magnetic lines. Electrons spiral down the tube, emitting radio waves in the process. The spot at which the electrons descend into Jupiter's atmosphere as even been detected from Earth. Figure 14.5 illustrates this interaction.

NASA has developed RadioJove a school project designed to introduce the public to radio astronomy.[1] The radioJove project includes a receiver kit, software and a simple antenna. This should be suitable to detect radio emission from Jupiter. However, in order to get higher signal to noise it might be worth investing in a suitable Yagi, a tracking mount and an off-the-shelf receiver.

However, independent of the kit you have access to, as long as you have a setup able to receive decimetre emission from Jupiter you should be able to do this practical.

14.3.3.1 Aims

Using a simple radio telescope and receiver you will observing the radio emission from Jupiter and interpret the variation in signal strength.

14.3.3.2 Observations

You need to observe Jupiter for a number of days (or nights) in order to get a complete understanding. If you only have a loop antenna then observing during

[1]The RadioJove project can be found at http://radiojove.gsfc.nasa.gov/

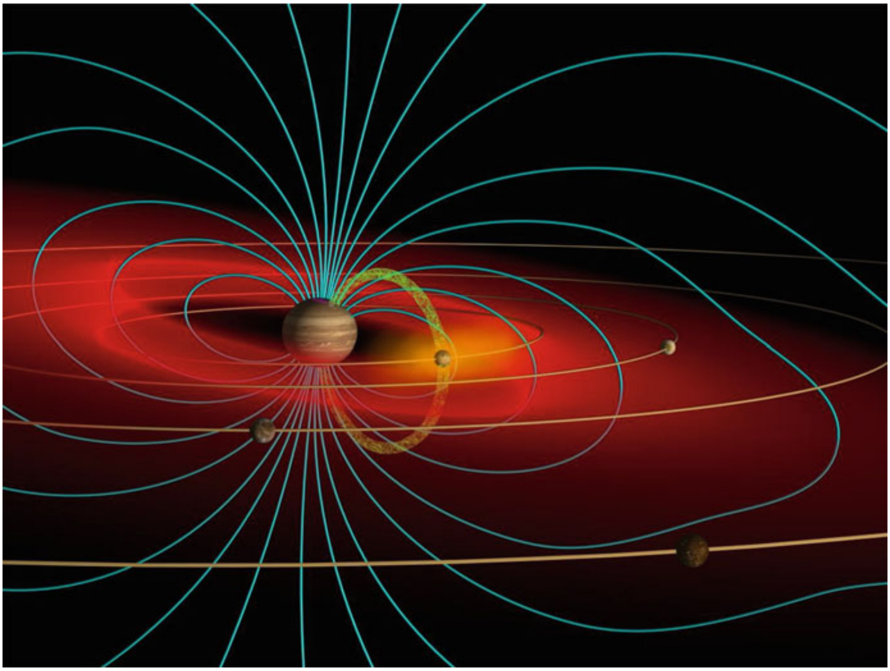

Fig. 14.5 Illustration of the magnetic interaction between Jupiter and Io. The flux tube can be seen in *green* (Image courtesy John Spencer. Copyright John Spencer CC BY-SA 3.0)

daylight may be difficult it the Sun in very active. If you have a directional antenna that can track then daylight observations are possible but better results can obtain after dark.

Turn on the receiver setting it to the correct frequency. If necessary point the antenna at Jupiter and start tracking. There are short-term and transitory, variation in the radio emission from Jupiter. You are looking for more long-term variation. To find this record the signal strength every 5 min then use the average of 12 signal strengths to find the mean hourly signal. Use this to also calculate your uncertainties.

14.3.3.3 Write Up and Analysis

Once you have all you data plot time against intensity, remembering to use error bars. Using your favourite planetarium program, plot the latitude of Io, Europa, Callisto and Ganymede against time. Compare these three plots with your radio data.

Write up your results in the standard format including your data, plots, uncertainties and any calculations. Discuss the nature of the variation in the radio signal and its relationship to the main Jovian moons.

Chapter 15
Astrometry

Abstract Astrometry deals with the position and movement of astronomical objects in the sky. It is one of the most ancient applications of astronomy and in its modern form it is used to determine to position of stars with the Galaxy, the orbits of comets and asteroids and detect the existence of exoplanets. We show how to perform astrometry using SAO DS9 to determine the proper motion of an asteroid and from there its distance from the Sun.

15.1 Introduction

Astrometry deals with the position and the movement of astrophysical bodies like the planets, asteroids, comets and stars. The production of an astronomical catalogue, with the positions of those objects, requires astrometric observations and calculations.

Astrometry is the oldest scientific application of astronomy as it was the astronomic measure of Tycho that allowed Kepler to develop his three laws of motion, Newton to discover his law of gravity. It also gave the first hint that Newton was incorrect which lead to Einstein developing general and special relativity which we first confirmed by astronomic of stars during an eclipse.

In modern astronomy astrometry helped in the discovery of the supermassive black hole in the Galactic Centre as well as, indirectly, the detection of exoplanets.

Astrometry is one of the techniques we can use to determine the distances to the stars. The method of Stellar parallax requires very high precision astrometry in order to measure the angular displacement of a nearby star when referenced to more distant stars, caused by the rotation of the Earth around the Sun. A 1-arcsec displacement is equal to a distance of 1 pc. However, parallax related changes for source beyond 100 pc are difficult to detect. Although there has been some progress with very long baseline radio interferometry that has positional resolutions of sub mill-arcseconds. This pushes direct distance measurements out to kiloparsecs for radio sources, some of which, for example, masers, are associated with other features such as star formation.

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The European Space Agencies (ESA), Gaia space mission, is attempting to measure the position of a billion stars with a positional accuracy of 24 milli-arcseconds. If successful this will be a major step forward in our goal of mapping the Milky Way.

Astrometry is also used to determine **Proper motion**, the movement of a star across the plane of the sky – as opposed to radial motion that is to or from the Earth. For example, a star's position is measured twice 10 years apart. The first position is (α_1, δ_1) and a decade later it is at (α_2, δ_2) The changes in the angle is $\mu_\alpha = \alpha_2 - \alpha_1 and \mu_\gamma = \gamma_2 - \gamma_1$. To find the proper motion we apply $\mu^2 = \mu_\gamma^2 + \mu_\alpha^2 \cdot \cos \gamma_1$.

Although this appears, rightist forward measuring changes is position are challenging over short time scales (i.e., years). The change in position is small, and the resolution of a ground-based telescope is limited by atmospheric seeing (hence Gaia being spaceborne). Finding the centre of a star can be done by finding the centroid of the Gaussian distribution of the star and repeating the processes increase the reliability of the measurement. Likewise using many stars, of known position, to locate the position of the unknown source, also reduces uncertainties. Once the direction position and hence proper nation vector of the star is known, we may have to account for the proper motion of the Sun and Earth depending on the application.

We have already briefly mentioned plate solving in this book. Plate solving is a form of astrometry in that relates pixel location to sky location. Plate solving programs such as Pinpoint determines the plate scale of your image from the pixel size and focal length of the image. It then identifies the stars within the field. At which point it uses one of the several algorithms to match star patterns in its library to those in the image. Once it has a match, it identifies the coordinate to pixel transform and write it into the header in the form of the World Coordinate System WCS, which might include scaling, rotation and field curvature corrections.

15.2 The Proper Motion of Asteroids

The identification of solar system objects, for example, comets and asteroids, is undertaken by their observed orbital properties. These orbital properties are determined from repeated observations over consecutive nights, if possible. Three nights is the minimum the Minor Planet Centre suggests.

In this practical you will observe an asteroid at opposition over success nights in order to determine its proper motion and approximate distance assume a circular Keplerian orbit.

Aims

You will be taking images of a bright asteroid at opposition over a number of evenings to determine its proper motion and its solar distance. You will also learn how to use the DS9 vector and blink function.

Planning

Using the Minor Planet Centre (MPC) website http://www.minorplanetcenter.net/iau/MPEph/MPEph.html you should identify an asteroid on, or near opposition that is bright enough, and in the right position, to be observed. You should download the Ephemeris for this target to be sure that you have the position for the days and time you will be observing. Create finder charts for each night as per normal.

Observing

- On each night of observing, as per usually cool the camera down and start the telescope up. As we are undertaking astrometry we don't need dark or bias frames;
- Move the telescope to a bright star, focus up and ensure the pointing is good. Make any adjustments as need;
- Move to the target (make sure you have the right RA and Dec for that night) and take a short image and using your finder chart confirm that you are in the right location;
- take a number of longer exposure images ensuring your stars in the image have Gaussian PSF;
- Plate Solve the images, using either a local program such as Pinpoint or an online program such as astrometry.net
- Save your images to a portable media and shutdown the observatory.

Reducation and Analysis

You will need to open your images in pairs is SAO DS9 with tiled frames and then blink the frame so you can identify the object.

- Open DS9 and from the frame menu select Tile frame and then New Frame;
- Click on the left-hand frame and from the file menu select open, and open your first image. Select the right-hand tile and load your second image into it;
- You will now have two FITS images displayed. On each tile in turn select Zoom fit to window and then scale 99.5 %;
- From the frame menu select frame match WCS. This will orient the two frames so that they match; You should have some like Fig. 15.1 displayed.
- select frame blink and the two frames will be displayed alternatively every half second. You object should be seen moving against the background stars;
- Make a note of its location in each frame. Repeat the produce for all your frames until you have located the asteroid in all your images;

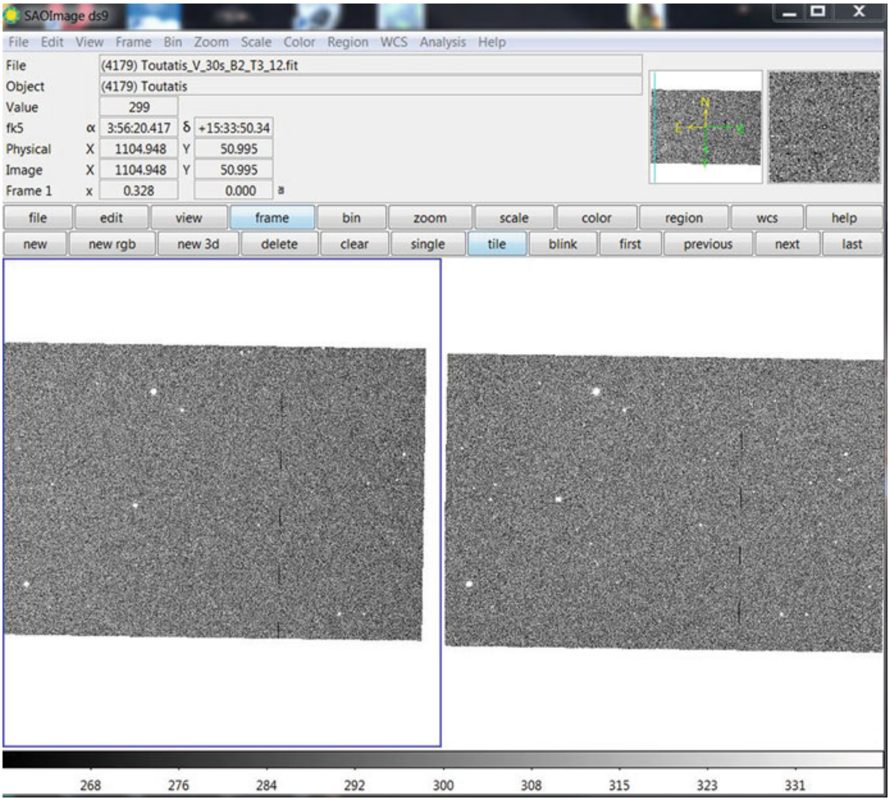

Fig. 15.1 Screen shot of two frames loaded into SAO DS9 prior to blinking. (Images University of Hertfordshire and the Smithsonian Astrophysical Observatory)

You now need to determine the position of the asteroid in each image. You could do this by just moving the pointer over the asteroid and reading off the WCS coordinates. However, this is not very accurate.

- There should be some bright stars in the field that you know the position of. It is best to have this is decimal degrees as working in Sexagesimal makes the math challenging;
- Pick forms the region menu shape vector. The draw a vector from the reference star to the asteroid;
- click on the vector to activate it and select region centroid. This step is very important as it determines the centre of the PDF for both the reference star and the asteroid;
- Double click on the vector to bring up its information window. You should note the vector's length in arcseconds and angle;
- Add vectors from at least three referees stars and use the vectors to determine the position of the asteroid. Use the mean position and the range as your error;

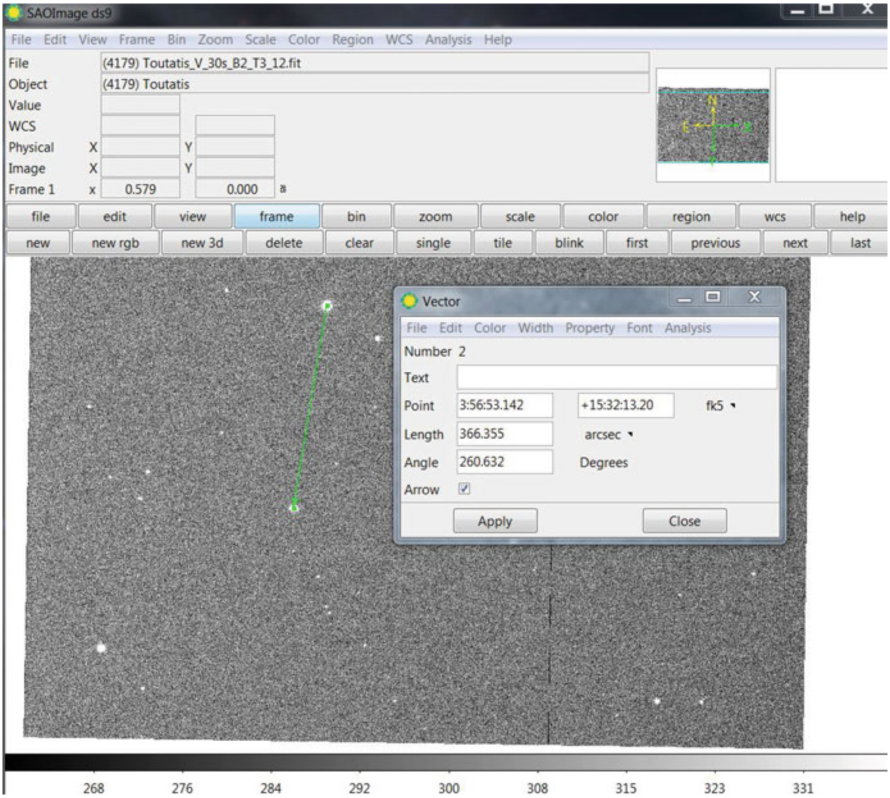

Fig. 15.2 Screen shot of a vector between an asteroid and a reference star and the Vector information window (Images University of Hertfordshire and the Smithsonian Astrophysical Observatory)

- Now you know the start and end position of your asteroid you use $\mu^2 = \mu_\gamma^2 + \mu_\alpha^2 \cdot \cos\gamma_1$ to determine the proper motion.

You should now be able to work out the distance of the asteroid from the Sun. Remember that the asteroid is at opposition and so the velocity vector for the Earth and the asteroid, are parallel. Assume that both orbits are circular, Sun centred and Keplerian (Fig. 15.2).

Write Up

Write up your report in the standard style, including all your calculations and images.

Discuss how you think your could improve your uncertainties and reflect on how you would deal with an asteroid that was not at opposition.

Glossary

AB system A photometric standard based on instrumentation.
Absolute Magnitude The magnitude of a body at a distance of 10 pc for a non-solar system object and a distance of 1 AU for an object inside the Solar System.
Absorption Lines Dark lines in a Spectra caused by the absorption of light by atom and molecules between the emitting source and the observer.
Achromatic A two-component lens designed to reduce chromatic aberration.
ADU Bit Number The number of bits that the processor of an ADU uses. This limits the maximum number it can work with.
Airy Disc The produced pattern by the diffraction of light within the optics of a telescope or other such device.
All-Sky Photometry A photometric technique that does not use reference stars.
Angular Resolution The telescope's ability to separate two objects and is given by the average wavelength of the light being observed divided by the aperture of the telescope
Aperture Either the diameter of an optical system such as a telescope of the size of the region being measured in photometry.
Aperture Photometer Tool A free-ware multi-platform application designed to perform aperture photometer.
Aperture Photometry Photometry using an aperture around an image of a star and the count within the aperture being integrated.
Apochromatic A three component lens designed to reduce chromatic aberration.
Apogee For an object in an orbit, this is the point at which it is furthest from the object it is orbiting.
Apparent Brightness The brightness of a source as a view from the Earth.
Apparent Magnitude The magnitude of a source as viewed from the Earth.
Apparent Solar Time The local solar time based the observation of the Sun
Arcminute A measure of angle. One 60th of a degree
Arcsecond A measure of the angle. One 60th of an arcminute

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Array Size The size of a CCD in pixels

Astigmatism A problem caused by the incorrect alignment of optical components.

AU The mean distance between the Sun and the Earth, 149,597,870,700 m. Used mainly for working with the solar system object.

Background In photometry the unwanted signal that does not come from the source, rather the Sky and internal sources of noise, etc.

Bahtinov Mask A mask that causes a distinctive pattern when the telescope is in focus.

Balmer Series A series of hydrogen lines caused by the transition of an electron down to the n = 2 energy level

Band A range of wavelengths. For example, the R-band is the wavelength over which a Johnson R-band filter is transparent.

Barlow A device to increase the effective focal length of a telescope

Bias The signal within a CCD frame that comes from the read process and other such sources.

Bias Frame A zero-second exposure image used to remove the bias.

Binning The joining of pixels to form one large pixel.

Black Body A perfect emitter and absorber of radiation

Blooming The spike effect caused by reading a saturated pixel

Bolometric Overall wavelengths. For example to the bolometric magnitude is the magnitude integrated over the entire spectrum.

Broadband Filters A filter that covers multispectral lines.

Cassegrain A telescope using a hyperbolic secondary mirror placed parallel to the parabolic primary reflecting the light down the tube and through a hole in the centre of the primary with the focal plane lying just outside of the tube at its rear.

Catadioptric A telescope design that uses a correcting lens (known as the correcting plate) in front of the primary mirror

CCD Charged Couple Device. The light sensitive part of a CCD camera.

Celestial Equator The Earth's equator projected onto the sky.

Charge Diffusion The loss of electrons during the CCD transfer process

Chromatic aberration The appearance of coloured rings in refractors, due to differing wavelengths being brought to differing focus points.

Circularly Polarised Polarised to the rotational orientation of the photons electric field rather than its linear orientation.

Circumpolar An object that is either visible to an observer all year round or never visible

Cold Pixel A pixel reporting low at all times

Collimation The alignment between the optical components in a reflecting telescope. A telescope out of Collimation has elongated point sources.

Colour Index The ratio of two fluxes in differing bands. Colour is a distant independent value.

Colour-Magnitude Diagram (CMD) A plot of magnitude verses colour (for example B v B-R).

Coma A problem with a reflector where light incident to the edge of the mirror to be brought to a slightly different focus to those at the centre, resulting in the image becoming smeared out.

Coordinated Universal Time A standard time based on GMT. The different between GMT and UT is currently minor.

Cosmic Ray A high energy charge particle. The Source of Cosmic Rays are not yet fully understood although high energy events such as supernovae are thought to be likely candidates.

Count A dimensionless instrument specific unit, indirectly related to flux.

Dark Adaption A combination of physiological responses that increase the ability to see in low light levels. Dark Adoption takes about 20 min to reach its optimal level. It is lost almost instantly when exposed to bright light.

Dark Current The noise caused by thermal electrons.

Dark Frame A image of the same exposure time as the light frame but with the shutter closed. Used to remove the dark current.

Diagonal An optical device that changes the pathway of light as it leaves the telescope.

Diffraction Halo A halo that appears around bright objects due to diffraction

Diffraction Spike A cross like spike appears around bright stars in images. Often added by amateurs in Photoshop

Digitisation noise Noise caused by an ADU being a digital device.

Dithering Differing. Moving the frame a few pixels between exposures so that the some sources are not exposed to the same pixels

Dome Flats Flat frames that use the inside of the dome as the illumination source

Ecliptic The path along which the Sun appears to travel.

Elliptic Coordinates A coordinate system using the ecliptic as the fundamental plane

Emission Line Bright lines in the spectra caused by the presence of atoms and molecules in the emitting source.

Emission Nebula Any nebula that emits light, normally due to embedded stars

Equation of Time The translation between Mean Solar Time and Apparent Solar Time

Equator The great circle on a sphere perpendicular to the meridian and equal distance from the poles

Exoplanets Planets orbiting other stars

Extended Objects A resolved object, for example, a planet or a galaxy.

Extinction Coefficient The level of extinction

Eye Relief A measure of the comfort of an eyepiece when being used.

Field Curvature A problem associated by the projection of a curved surface (the sky) onto a flat one (the CCD).

Field of View The amount of sky visible through the instrument, measured in degrees or arcminutes.

Field Stop The aperture of an eyepiece.

Filters A device designed to allow only a range of specific wavelength to pass through to the detector.

FITS Flexible Image Transport System. The standard format for astronomical images.
FITS Header The part of a FITS image contain image data but not imaging data.
Flat Fielding A process to deal with non-uniformity of a CCD.
Flat Frame An image of the twilight sky or inside of the dome. Used for removing the flat field noise
Flip Mirror A flat mirror which can be moved from a position where the mirror is parallel to the telescope to one where it is 45° with the telescope
Flux The amount of energy transferred through an area in a set time.
Focal Length The distance between their first optical component and the first point of focus. A main characteristic of a telescope.
Focal Reducer A device to decrease the effective focal length of a telescope
Forbidden Line A spectral line that cannot be produced on Earth due to the low gas densities needed.
Free-Bound Emission Emission caused by the capture of a free electron. This contributes to continuum emission.
Full Well Capacity The number of electrons a pixel can hold
FWHM The Full Width Half Maximum, a measure of a spread of a Gaussian profile such as seen with a star.
Gain The ratio of photon to count in a pixel in a CCD
Gravitational Reddening The general relativity effect that causes the light from high gravity environments to have an increased wavelength.
Great Circle A circle intercepting the centre of a sphere so that this circle with the same circumference as the sphere
Greenwich Mean Time The mean solar time at the Royal Observatory in Greenwich
Hertzsprung-Russell (HR) diagram A plot of absolute magnitude against colour. The HR diagrams describe stellar evolution and is one of the foundations of modern astrophysics.
Home The zeroing position of a telescope
Hot Pixel A pixel reporting high at all times
Hour angle Local Sidereal time minus the RA of the object.
hour circle Any great circle celestial sphere that passes through the celestial poles perpendicular to the equator.
Hubble Palette The guidelines for how the narrow band Hubble filter set is displayed in colour images
Hydrogen Alpha A Balmer series emission line at 656.28 nm, caused by a $n = 3$ to $n = 2$ hydrogen transition.
Image Processing The process of producing a high-quality image by manipulation of the data.
Image Reduction The process of turning a light frame into a science frame
Integrated magnitude The total amount of radiation in a set band coming from an object, expressed as a magnitude. Most often used for extended sources.
International Astronomical Union (IAU) The governing body for astronomy made up of professional astronomers from around the world

Interstellar reddening The reddening of light from object caused by the preferential scattering of short wavelength over long ones.

IRAF A professional Linux data reduction package for astronomy.

Isotropically In every direction. Most stars radiate Isotropically.

Jansky A unit of flux used in radio astronomy.

Julian date A standard dating system using the day as the base unit and the origin being noon on 1st January 4713 BC

Lab Book A journal containing a complete description of all observations and experiments including calculations and observations. A good Lab Book should enable other investigate to reproduce the lab book's authors experiments.

Light Frame The unprocessed digital image of an astronomical object

Light Pollution Any source of man-made stray light.

Limb Darkening The effect whereby the edge of gaseous bodies such as stars appears to darken at the edge.

Linearly Polarised Polarised to the linear orientation of the photons electric field rather than its rotational orientation.

Local Sidereal Time The sidereal time local to an observer.

Lucky Imaging A image technique using a high-speed video camera in order to produce high-resolution images.

Luminance Frame A image normal in Clear, used to enhance detail in a colour image

Luminosity The total amount of energy radiated by a body over a specific band width. If calculated over the entire spectrum its is known as the Bolometric Luminosity.

Lyman series A series of hydrogen lines caused by the transition of an electron to a lower energy level.

Magnification The ratio of the angular size when using the telescope and when not using it.

Main Sequence The area of the HR diagram that represents stars that are core hydrogen burning.

Maksutov-Cassegrain A Catadioptric which used the correcting plates as the reverse of the secondary.

Master Frames A calibration frame made by combine multiple frames

Mean Solar Time The mean time as measured by using the Sun, the principle unit of which is the day.

Meridian The great circle on a sphere perpendicular to the planes of the celestial equator and horizon

Meridian Flip A problem with equatorial mounted telescopes that require them to 'flip' over when a source crossed the meridian.

Metallicity The ratio of Hydrogen v another heavy element, normally iron, in as star

Narrow Band Filter A filter that covers one spectral line.

Neutral Density Filter Very broad filter that removes light equally over a very large range of wavelengths. These are normally used to cut down the light from a bright source such as the Sun or Moon.

Neutrino A subatomic particle with zero charges and an extremely low mass. The Neutrino interacts weakly with matter and is produced in vast numbers during nuclear synthesis.

Newtonian A reflecting telescope where a flat mirror set at 45° to the primary mirror and just inside the focal point so that the point of focus is outside the tube, often close to the top.

Nodding Going off target. For example when you cannot get a representative background in the current image.

Parsec The distance at which an object would show a parallax displacement of one second of arc. It equates to $3.08567758 \times 10^{16}$ m.

Pedestal The charged added to a pixel to prevent wrap around during mathematical processes.

Perigee For an object in an orbit, this is the point at which it is closest to the object it is orbiting.

Photoelectric effect An effect by which certain materials produce a current only exposed to light of a certain wavelength.

Photometry The science of measuring the brightness of an object.

Pipeline A series of programmes used to process automatically an image.

Pixel Size The size of an individual pixel

Planck's Law A physics law that describes the relationship of emission to the temperature of a blackbody

Planetary Nebulae A nebula formed by the ejection of the outer layers of a red giant.

Plate solving A process by which the WCS transform for an image is found.

Point Spread Function In astronomy how the light from a source spreads out over the CCD. For a star, this should be a Gaussian

Polar Aligned In terms of telescopes mounts, a mount where one axis is aligned with the celestial poles.

Polarisation Filters A filter that only lets pass photons whose electric field is orientation in a specific direction.

Population I Young stars like the Sun

Population II Older Stars such as found in globular clusters

Population III The very first stars. Yet to be detected

Propagation of Uncertainties The methods by which mathematical operations effect errors.

Proper Motion The motion of an object against the background stars.

Proton-Proton Chain The nuclear process chain that converts hydrogen to helium resulting in the production of energy. There are four such chains, with a varying degree of importance within a star dependent on its mass.

Quantum Efficiency In relation to CCD, the efficiency of the pixels at converting photons to electrons.

Quantum mechanics The fundamental branch of physics which describes the universe at scales the atomic and subatomic length scale.

Rayleigh Criterion The non-Airy resolution of a telescope.

Read Noise Noise created by the process of reading the CCD.

Readout noise Noise associated with the ADU.

Residual Bulk Image Ghost images caused my saturation in a previous image

SALSA J A free-ware multi-platform application designed to perform time series photometer.

SAO DS9 A FITS manipulation package built by the Smithsonian Astronomical Observatory.

Saturation The point at which a pixel is full and can detect no further photons during this exposure

Schmidt-Cassegrain A class of reflecting telescope based on the Ritchey and Chret́ien design but with a correcting plate.

Science Frame A light frame that has had dark, bias and flat frames applied.

Seeing The angular dispersion of stars caused by atmospheric turbulence.

Sexidecimal notation The description of angles as hours minutes and seconds (of degree minutes and seconds) as opposed to decimal degrees.

Shot Noise Noise that originates from the discrete nature of the source. For example the photon or electron.

Sidereal Time Time as based on the Earth's rotation measured relative to the fixed stars" rather than the Sun.

Signal to Noise Ratio Simple the radio between the source signal and the noise. High signal to noise data has less uncertainty

Sky Brightness The intensity of the light being emitted and reflection of the sky. The sky brightness is the main limiting factor to deep-sky observations.

Sky Flats Flat frames that use the twilight sky as the illumination source

Small Circle A circle on the surface of a sphere that does not intercept the centre of the sphere

Spectra The rainbow-like distribution of light made by a Spectrograph.

Spectrograph A device that split light regions of similar wavelength.

Spherical aberration A optical problem associated by a mirror being spherical. Light incident at differing point in the mirror come to focus at different points. The Hubble Space Telescope famously suffered from Spherical aberration.

Stopping Down Reducing the aperture of a telescope

Summing Co-adding images

Surface brightness The magnitude per unit area.

Transit telescopes A telescope designed just to observe objects as the cross the meridian.

Transmission window A region in the electromagnetic spectrum that a body is transparent to. Typically refers to the Earth's atmosphere.

Vega System A photometric standard based on the apparent magnitude of Vega in that band.

World Coordinate System A system to convert pixel position to sky position

Zero Point A calibration number to convert an instrument magnitude of flux to a standard one. Normally changes between nights.

Further Reading

Budding E, Demircan O, (2007) Introduction to Astronomical Photometry. Cambridge University Press

Burke B.F, Graham-Smith F. (2010) An Introduction to Radio Astronomy. Cambridge University Press

Howell S.B, (2006) Handbook of CCD Astronomy. Cambridge University Press

Warner B.D (2006) A Practical Guide to Light curve Photometry and Analysis. Springer

Kitchin C.R. (2013) Astrophysical Techniques, Sixth Edition. CRC Press

Roy A.E.,Clarke D. (2003) Astronomy: Principles and Practice, Fourth Edition. Institute of Physics Publishing

Arnold S., (2003) Getting Started in Radio Astronomy: Beginner Projects for the Amateur (The Patrick Moore Practical Astronomy Series) Springer New York

Lashley J. (2010) The Radio Sky and How to Observe It (Astronomers' Observing Guides). Springer New York

Harrison K.M.(2011) Astronomical Spectroscopy for Amateurs (The Patrick Moore Practical Astronomy Series). Springer

Bracken C. (2013) The Deep-sky Imaging Prime. CreateSpace Independent Publishing Platform

Berry,R (2005) The Handbook of Astronomical Image Processing. Willmann-Bell

Ireland R.S (2009) Photoshop Astronomy. Willmann-Bell

Birney D.S, Gonzalez G., Oesper D. (2006) Observational Astronomy. Cambridge University Press

Fleish D. kregenow J. (2013) A Student's Guide to the Mathematics of Astronomy. Cambridge University Press

Northedge A., Thomas J., Lane A., Peasgood A., (1997) The Sciences Good Study Guide. The Open University

Chromey F.R., (2010) To Measure the Sky: An Introduction to Observational Astronomy. Cambridge University Press

Taylor J. R., (1997) An Introduction to Error Analysis: The Study of Uncertainties in Physical Measurements.University Science Books,U.S

Wall J.V, Jenkins C.R, (2003) Practical Statistics for Astronomers. Cambridge University Press

Kitchen C.R, (2000) Solar Observing Techniques (The Patrick Moore Practical Astronomy Series) Springer

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