Advanced Lab Course

Photometry of Star Clusters

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1 Motivation & Overview

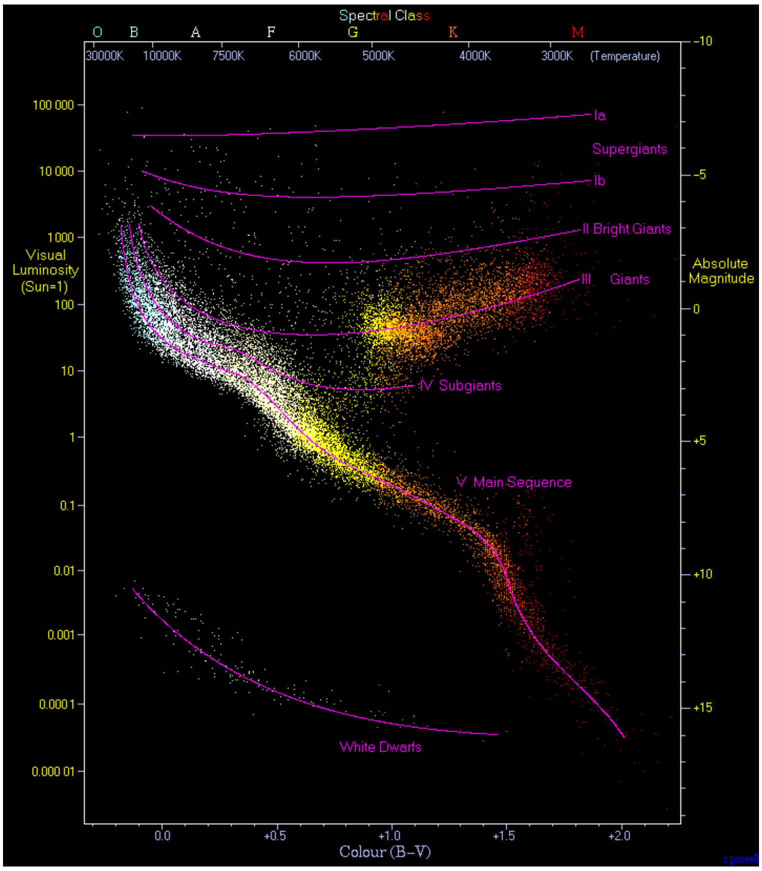


Figure 1: A typical Hertzsprung-Russell/temperature-luminosity (HRD) or color-magnitude diagram (CMD) showing 22000 stars from the Hipparcos Catalogue together with 1000 low-luminosity stars (red and white dwarfs) from the Gliese Catalogue of Nearby Stars. Valuable information on a star's evolutionary state can be derived from its position within the diagram (source: http://www.atlasoftheuniverse.com/hr.html).

Star clusters belong to the most important objects in the Universe. First of all, they are the fundamental building blocks of galaxies, because it is nowadays believed that most, if not all, stars are born in such groups of a few dozen up to several million stars. Hence, nearly all stars in galaxies have once been member of a cluster. That makes it crucial for our understanding of galaxy formation and galactic stellar populations to investigate these objects in detail.

Moreover, star clusters are unique test beds for stellar-evolution theories. Their most important property, that all member stars were born in a single star-burst out of one giant molecular cloud, enables us with a few simple observations to get a snap-shot of stellar evolution of a whole population of stars at a certain age, distance and metallicity. Without star clusters our knowledge on stellar evolution wouldn't be nearly as detailed as it is today.

The other way round, observations of star clusters offer the possibility to easily derive their ages, distances and metallicites by comparing the observations to theoretical stellar-population models of certain compositions and evolutionary stages. This latter application is the main objective of the underlying lab-course project.

For this purpose, the Hertzsprung-Russel diagram (HRD, Fig. 1) can be used which during the last century proofed to be the optimal tool for studying stellar populations and stellar evolution. Main objective of this lab-course project is therefore the understanding, preparation and analysis of a color-magnitude diagram (CMD), which is a direct derivative of the HRD.

This script is organized as follows:

- Section 2 gives an introduction to the subject of this project, i.e. star clusters and the color-magnitude diagram.
- Section 3 introduces the basic concepts of astronomical observations as well as the techniques and instruments you will use in this project.
- Section 4 covers the observations you will take out, how you prepare them properly and how to carry them out such that you can get valueable data.
- Section 5 is about the data reduction of your images and how you extract the information you need from the images.
- Section 6 is the main scientific part where you will be guided through the analysis of your data.
- The appendix will help you with the THELI software package which you will use during this project.

Please read all sections (except the appendices) carefully **before** you start. Make sure you answered all questions in the text, since you are not allowed to start the project before answering all questions.

2 Basic Knowledge of Star Clusters and CMDs

2.1 Star Clusters

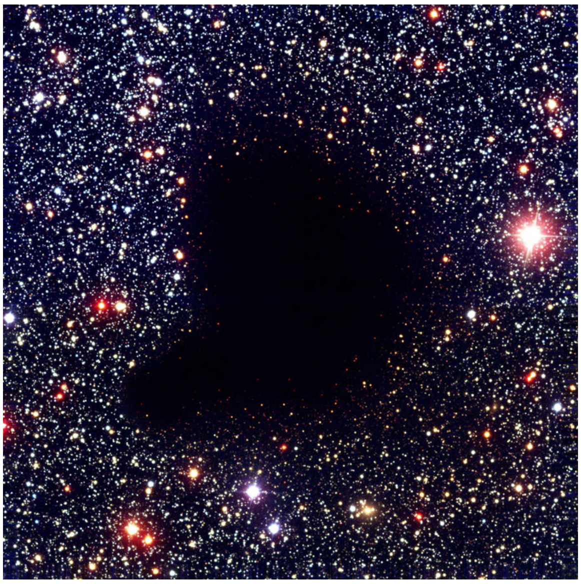


Figure 2: Picture of the giant molecular cloud Barnard 68 which is relatively nearby, with a distance of about 200 pc and a diameter of about 0.2 pc. The cloud can only be seen indirectly in optical wavelengths as it hardly emits light whereas it absorbs all light coming from background sources. It is not known exactly how molecular clouds like Barnard 68 form, but it is known that these clouds are themselves likely places for new stars to form. In fact, Barnard 68 itself has recently been found likely to collapse and form a new star system (picture from the FORS Team with the 8.2-meter VLT, ESO).

Star clusters play a key role in the development of our understanding of the Universe. But what exactly is a star cluster? In principle any agglomeration of more than a few stars may be called like this, where "a few" is not well specified and may be taken to be around ten. In terms of star-cluster dynamics, since the constituent stars mutually attract each other by the force of gravity, a star cluster may also be defined as a dynamically bound system of a number of stars.

The fundamental property of a star cluster is its origin, which is assumed to be a single giant molecular cloud for all members of one cluster (Fig. 2). The according formation scenario of star clusters is quite well understood nowadays: a collapsing cloud fragments into small clumps which form the progenitors of the cluster stars. Depending on the size of the molecular cloud and on the conditions of its collapse, a fraction of about 10-30% of the gas is consumed by star formation (Adams & Myers, 2001; Lada & Lada, 2003; Allen et al., 2007). The masses of the stars produced thereby follow a more or less universal distribution function, which is a quite simple power-law: the so-called initial mass function (e.g. Salpeter 1955; Kroupa 2001). An important feature of the initial mass function (IMF) is that it predicts a large number of low-mass and just a few very massive stars (Fig. 3).

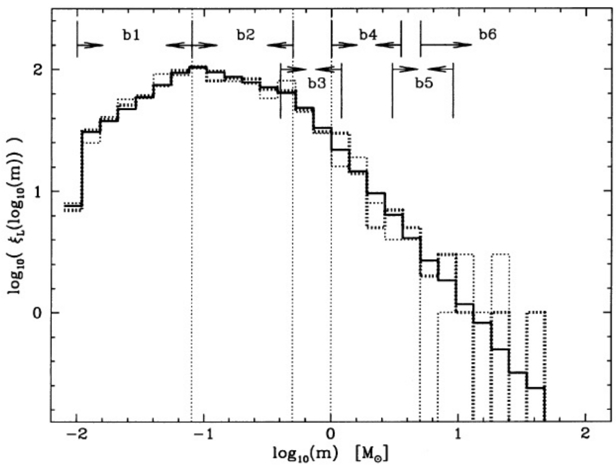


Figure 3: Logarithmic representation of the canonical IMF, ξ (solid histogram). The histogram gives the probability ξ of forming a star of mass m in a star cluster. Between $0.08\,\mathrm{M}_\odot$ and $0.5\,\mathrm{M}_\odot$ the IMF has the slope -1.35. For higher stellar masses it has a slope of -2.35. From this figure it is obvious that it is much more likely to form low-mass stars than high-mass stars. Figure taken from Kroupa (2001).

T2.1: *The canonical IMF, $\xi(m)$, has the form:*

$$\xi(m) = 0.237\, m^{-1.35} \quad \textit{for} \quad m \leq 0.5\,\mathrm{M}_\odot, \tag{1}$$
$$\xi(m) = 0.114\, m^{-2.35} \quad \textit{for} \quad m > 0.5\,\mathrm{M}_\odot, \tag{2}$$

which is normalized such that the integral over ξ from $m = 0.08\,\mathrm{M}_\odot$ to $m = 150\,\mathrm{M}_\odot$ is equal to 1. If you draw 1000000 stars from this IMF, say for the simulation of a globular cluster, how many will be below $0.5\,\mathrm{M}_\odot$? How much mass will be in stars below $0.5\,\mathrm{M}_\odot$?

When the first newly formed stars ignite, the cluster will still be embedded in its birth gas cloud. If the initial cloud was rich there will be a couple of large, so-called, O- and B-stars with masses of up to $150\,\mathrm{M}_\odot$, where this upper mass limit is still wildly discussed (Kroupa, 2005). As can be seen in Fig. 4, these luminous stars will soon blow out the left-over gas and free the cluster from its birth cradle with their enormous radiation pressure and as a cause of ongoing supernovae.

What is left is an ensemble of stars which is more or less tightly bound, strongly depending on the initial conditions and the birth parameters. If the initial gas loss is too violent the cluster will dissolve within a short time, otherwise it will virialise within a few million years and from then on dissolve slowly (Boily & Kroupa, 2003a,b).

Moreover, based on observations of pre-main-sequence stars, a primordial (i.e. initial) binary fraction of about 100% has been found (Kroupa, 1995) which means that almost every clump in the collapsing gas cloud splits into two subclumps and in the end yields two distinct stars which form a binary system. The actual value of the binary fraction of observed clusters is still a debated topic, though. Due to insufficient spa-

Figure 4: Two young star clusters. On the left NGC602, a star forming region, where the massive stars in the centre already started to blow out the gas. On the right the centre of the Pleiades cluster, which is about one hundred million years old and exhibits almost no more gas (pictures taken from NASA and the STScI).

tial resolution this question cannot be answered directly by observations yet, hence uncertainties are still quite large. The binary fraction of all stars in the Milky Way is believed to be about 50%, i.e., every second star on the sky has a companion which in most cases cannot be seen with the naked eye, while relatively young clusters like the Pleiades (Fig. 4) and Hyades show fractions of about 60-70% (Kroupa, 1995).

T2.2: *The binary fraction of a star cluster, f_{bin}, is defined as*

$$f_{bin} = \frac{N_{bin}}{N_s + N_{bin}}, \tag{3}$$

where N_{bin} is the number of binary systems whereas N_s is the number of single stars. Imagine you observe a star cluster and you detect 1000 point sources but you know the cluster has a binary fraction of 0.7, how many stars are in this cluster?

2.2 Open Clusters vs. Globular Clusters

Historically, star clusters are split up into two distinct populations: open clusters and globular clusters. Although these two families of stellar groups exhibit significant differences (see Table 2.2), the definitions (as many other things in astronomy) are not strict and there is even a number of objects (like the cluster Westerlund I) that cannot be clearly assigned to one of the two.

Figure 5: The colliding galaxies NGC4038 and NGC4039, also known as the Antennae galaxies. During this merger of two gas-rich spiral galaxies thousands of star clusters have formed and are still being formed. Actually, the bright blue points in this Hubble view are not single stars or star clusters but clusters of star clusters. Credit: Brad Whitmore (STScI) and NASA.

Open clusters present the lower mass range of star clusters. They assemble from a dozen up to several thousand stars in a region with a diameter of 1-10 pc. Hence densities vary significantly from cluster to cluster and reach from about 0.1 $M_{\odot}pc^{-3}$ (which may rather be called associations than clusters) up to 10^3 $M_{\odot}pc^{-3}$.

Globular clusters on the other hand are rich clusters with 10^4 to 10^7 stars and diameters of 20 to 150 pc. Unlike open clusters, which are often asymmetric and less centrally concentrated, these systems are quite smooth and spherical. They furthermore show a very high concentration in the core which extends from 0.3 up to 10 pc. Typical densities of these core regions are 10^4 $M_{\odot}pc^{-3}$, thus lie clearly above open clusters and even represent one of the densest stellar environments in the Universe.

In addition to their different dimensions and shapes the two species of clusters show other fundamental differences. Taking a closer look at pictures of open and globular clusters in the Milky Way immediately shows that the former in many cases have diffuse emission from gas while the later do not. Also, open clusters show bright blue stars which are young and massive, while they are completely missing in globular clusters.

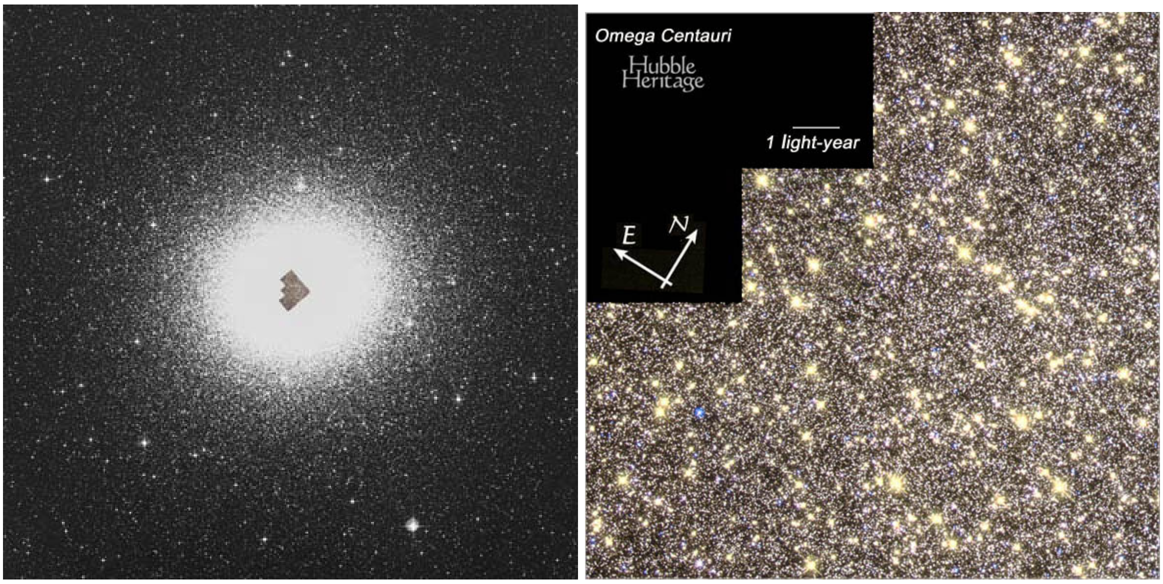


Figure 6: The globular cluster ω Centauri is the largest known cluster in the Milky Way. It is about 50 pc in diameter and consists of about 10 million stars. On the right is a close-up of the innermost region of ω Centauri, observed with the Hubble Space Telescope (pictures taken from the Digitized Sky Survey, NASA and the STScI).

From the location of the turn-off point in the Hertzsprung-Russell diagram the age of a cluster can be derived quite accurately. The ages of open clusters found in this way range from a few million years up to about 10 billion years while globular clusters may have solely formed about 11-13 billion years ago - at least this is true for most of the known globular clusters of the Milky Way and other galaxies. However, as mentioned above, there are some young luminous clusters that resemble in size and mass a globular cluster.

Despite some exceptions, the two groups therefore seemed to be of completely different origin. Open clusters form continuously, while globular clusters came into existence when the Universe was much younger and thus denser. Actually, globular clusters are even some of the oldest objects in the Universe and therefore put a strong limit on the age of the Universe, but also on the understanding of stellar evolution and on structure formation after the Big Bang.

Another hint for a fundamental difference of their origin was seen in the distribution of the two types of clusters in the Milky Way and other spiral galaxies as for example Andromeda. Globular clusters are randomly distributed in the halo of the host galaxy and follow random orbits while open clusters are solely found in the galactic disk.

All this was explained by taking their different formation processes into account:

- As mentioned above open clusters form in the disk, as this is the only place where enough material can assemble to induce star formation. After birth the newly formed cluster still moves around the galactic centre in the plane of the disk, just as the birth cloud has done before.

- In contrast, the majority of globular clusters may have formed while larger structures as the Milky Way have not existed yet, hence do not have to share the rotation of the final galactic disk of their host galaxy. Some of the globular clusters may have even been captured by their host galaxy during merger events (for further details see for example Binney & Merrifield 1998).

The modern picture interprets both families of clusters as the low-mass and high-mass part of the same initial cluster mass function (ICMF), which, similar to the IMF, is a power-law. Observations of star-burst galaxies like the Antennae galaxies (Fig. 5) show that in one star-formation event clusters of all masses are formed, of course, depending on the available material and the local star formation rate (SFR). That is, globular-clusters-like objects can only form in rich molecular clouds with a high SFR - conditions which especially existed in the very early Universe or during the merging of two spiral galaxies (Kroupa & Boily, 2002). Today in the Milky Way, as in most spiral galaxies, star formation is restricted to the Galactic disk since this is the only place where enough gas can accumulate to induce gravitational collapse.

The Milky Way exhibits 150 to 200 globular clusters with typically 10^5 stars. Open clusters are more frequent as they are produced continuously, so there are about 1000 known in the Milky Way. But since they all lie in the Galactic plane most of them are heavily obscured by dust. The total number of open clusters in the Milky Way therefore is supposed to be about 20000.

Probably the most famous open cluster of the Milky Way is the Pleiades Cluster, which is a typical galactic cluster with about 1000 stars, a diameter of approximately 10 pc and an age of about 125 million years. The Pleiades can be well seen with the naked eye due its small distance of roughly 135 pc and the bright heavy stars in its centre (see Fig. 4).

By far the most massive cluster of the Milky Way is ω Centauri (Fig. 6) with more than 10 million stars and a diameter of about 50 pc. With its enormous mass, ω Centauri represents the upper mass limit of a star cluster, therefore may also be classified as an ultra compact dwarf galaxy or the nucleus of a stripped dwarf spheroidal galaxy (e.g. Fellhauer & Kroupa 2003). In addition, recent observations show that it is indeed much more complex than a regular star cluster since its temperature-luminosity diagram shows pronounced substructure, i.e., ω Centauri may consist of more than one stellar population (Hilker et al., 2004).

T2.3: *Calculate the angular size of:*

- *a typical globular cluster in the halo (50 pc diameter at a distance of 10 kpc),*

	Abundance	No. of stars	Diameter [pc]	Age [yr]
Open clusters	10^4	$10^1 - 10^4$	$10^0 - 10^1$	$10^6 - 10^{10}$
Globular clusters	10^2	$10^4 - 10^6$	$10^1 - 10^2$	10^{10}

Table 1: Rough overview of the basic paramters of open and globular clusters. Abundance gives the estimated number of clusters in the Milky Way.

- *a typical open cluster in the disk (5 pc diameter at a distance of 1 kpc),*
- *the full Moon (10^{-10} pc diameter at a distance of 10^{-11} kpc).*

2.3 HRD/CMD

In 1913 Henry Norris Russell presented his recent work at a meeting of the Royal Astronomical Society. There he showed a diagram that represented a relation between the spectral classes and the absolute magnitudes of all stars for which fairly reliable distances had been obtained so far (Russell, 1913). Soon the importance of this discovery became clear to the astrophysical community and since then the so-called Hertzsprung-Russell diagram (Ejnar Hertzsprung was the first to anticipate the existence of a relation between the two quantities) has given a great contribution to the understanding of stellar evolution.

Hitherto, much effort has been put into this specific type of diagram and it was found that it is possible to replace spectral class on the abscissa by temperature and absolute magnitude on the ordinate by the star's luminosity to obtain a similar diagram: the temperature-luminosity, or color-magnitude diagram (Fig. 1). The advantage of the latter lies in the way the corresponding quantities can be obtained. While it is rather hard to determine the spectral class of a star, it is much easier to obtain its temperature in means of a color index. So, just by taking pictures of a star cluster with two different filters, the temperature/color and the apparent luminosity/magnitude can be derived and a color-magnitude diagram (CMD) can be drawn. The CMD will show on its y-axis the absolute magnitude of the stars, while on the x-axis there will be the *color*, defined as the difference between the values of the magnitudes in the two used filters (e.g., B-V if V is on the y-axis). How this is done will be subject of this lab-course project.

The evolution of stars within a CMD has been widely studied and is a major subject of every basic astronomy training. The most common classifications of stellar objects like main-sequence star, red giant, white dwarf, etc., are derived from this type of diagram and are taken to be well-known by the students carrying out this lab-course project. An advanced overview on stellar evolution can be found in Binney & Merrifield (1998), p. 258, or, more detailed, in de Boer & Seggewiss (2008).

2.4 Colors & Magnitudes

To understand the various flux measurements we make with our telescope setting, we have to clearify first which quantities exactly we measure and how we can derive physical quantities out of them. Here we introduce you to the basic principles you need for carrying out this project, a more complete picture is given in Binney & Merrifield (1998), p. 26.

2.4.1 Apparent Magnitude

The apparent magnitude, m, is a measure of a star's brightness as seen by an observer on Earth. In other words, it is the integrated radiation flux, f, measured in $\mathrm{W/m^2}$ contained in a particular frequency range, $\Delta\nu$:

$$m = \int_{\nu_1}^{\nu_2} f d\nu. \tag{4}$$

Measuring fluxes of celestial bodies in different frequency ranges is called astronomical photometry which is the main objective of this lab-course project.

Based on a magnitude system first introduced by ancient Greek astronomers, the apparent magnitude is a comparative scale in which brighter stars are given smaller magnitudes than fainter stars, such that

$$m_1 - m_2 = -2.5 \log_{10}\left(\frac{f_1}{f_2}\right), \tag{5}$$

where the indices denote the magnitudes and fluxes of two distinct stars.

But there is extinction of radiation through intergalactic gas, interstellar gas and, of course, through the earth's atmosphere. In addition, there's also flux getting lost in our telescope. Thus, the flux which reaches the solar system is not the flux which was emitted by the source and the flux received by our detector, f, is not the flux f_ν that reaches the solar system. While former varies from source to source, the latter is fixed and can be quantified for any instrument. We measure

$$f \equiv \int_0^\infty f_\nu T_\nu F_\nu R_\nu \mathrm{d}\nu, \tag{6}$$

where T_ν is the transmission of the atmosphere, F_ν is the transmission of any applied filter and R_ν is the efficiency of the telescope system.

Depending on the frequency range, there is significant extinction due to the Earth's atmosphere, T_ν. The amount of extinction furthermore depends on the column density of air along the line of sight and increases with lower observing angles. Therefore, the best possible observing conditions are achieved near the zenith, where the length of the light path through the atmosphere is minimal. Extinction and disturbing seeing effects get worse for observations at lower elevations. These effects are often expressed

in terms of the objects airmass a, which tells you through how much atmosphere (column density) the light travels compared to vertical in-fall. For an angular distance z from the zenith it can in good approximation be computed as $a = 1/\cos z$ such that $a = 1$ for an object at the zenith and formally $a = \inf$ at the horizon.
The filter transmission F_ν is readily determined for well defined filters like the Johnson filters which will be used in this lab course (Sec. 2.4.5). The instrumental efficiency R_ν is a composite of the efficiency of the telescope's optical system and the sensitivity of the CCD.

In order to get well defined magnitudes of star-cluster members for a color-magnitude diagram we can correct the observed fluxes of the stars by observing a reference star which is nearby the cluster and whose magnitudes in different frequency ranges/filters are well known, since the observing conditions (air mass, filter, telescope optics, CCD) for such a reference star are approximately the same as for the cluster. By measuring f for the reference star as well as for the cluster stars and by knowing the magnitude of the reference star, we can obtain the magnitudes of the cluster stars using equation 5.

Note: the flux of a single star is mostly determined by fitting a two-dimensional Gaussian distribution to the CCD image. This Gaussian results from the point-like appearance of the star which gets convolved with the telescope point-spread function (PSF). By integrating over this distribution function the total flux, f, of the star can be determined and converted to magnitudes as described above. In contrast to stars, galaxies and star clusters often appear as extended sources on the CCD. Their magnitude is determined by fitting appropriate distribution functions to the CCD image and integrating over these functions out to a pre-defined cut-off radius. In this way it is possible to define magnitudes for extended sources, but these values have to be handled with care as there are always underlying models which have been assumed and which may differ significantly from case to case.

T2.4: *You observe a reference star and measure* 16000 *counts, from the literature you know that this star has an apparent magnitude of 15.0. For a second object you measure* 4000 *counts. What is the apparent magnitude of this object?*

2.4.2 Distance Modulus

The observed flux f of an object depends not only on its intrinsic brightness but also on its distance, d. In fact, the flux decreases with d^{-2}. If we want to calculate the flux F of a certain object assuming that it was at distance D we can therefore use

$$f = \left(\frac{D}{d}\right)^2 F. \tag{7}$$

In this context, the absolute magnitude M is defined as the apparent magnitude an object would have if it was located at some standard distance D, where this distance is always taken to be 10 pc. Using equation 5 we then get

$$m - M = -2.5 \log_{10} \left(\frac{f}{F} \right) = 5 \log_{10} \left(\frac{d}{D} \right) = 5 \log_{10} d - 5, \tag{8}$$

where the quantity $(m - M)$ is called the *distance modulus* of the specific object. Hence, by knowing m and d we can correct the apparent magnitude for the non-standard distance. On the other hand, if we know m and M we can infer the distance d.

T2.5: *From the literature you know that the abolute magnitude of your object from T2.4 is* $M = 10$ *mag. What is the distance of the object, and what is the distance modulus?*

2.4.3 Interstellar Extinction

Absorption and scattering of photons in the interstellar medium can cause stars to appear dimmer than they actually are. This effect is called interstellar extinction and has to be handled with care. If we have A magnitudes of extinction then equation 8 has to be rewritten as

$$m - M = 5 \log_{10} d - 5 + A. \tag{9}$$

Fortunately, interstellar extinction is, unlike the distance effect, strongly wavelength dependent such that it can be measured by taking images in multiple color filters. For many objects reliable extinction measurements are therfore available

T2.6: *How does the distance of your object from T2.4 change in the case you have 0.2 mag of extinction?*

2.4.4 Metallicities

In astronomy all chemical elements heavier than He are called metals. In the first few minutes after the Big Bang just a very low percentage of all baryons was synthesized to metals while most baryons synthesized to hydrogen and helium. In fact, accurate primordial abundances have already been predicted by Alpher, Bethe & Gamow (1948) out of theoretical Big-Bang nucleosynthesis considerations before first accurate measurements were made. The standard values which are assumed for primordial abundances in most stellar-evolution calculations are $X = 0.765$ for hydrogen, $Y = 0.235$ for helium and $Z = 0$ for metals.

Through stellar evolution, the metallicity in the Universe increases as hydrogen and helium are processed into heavier elements. Nuclear fusion in stars in combination with stellar winds and supernova explosions permanently enriches the interstellar material. The solar metallicity, for instance, is about $Z = 0.02$, hence it has formed out of already enriched gas. Stars which form out of enriched material evolve differently from metal-poor stars as the metal content has a large influence on stellar evolution and the stellar structure. For instance, higher metalicities cause stars to become dimmer and cooler below a stellar mass of about 4 $M_\odot$. Above this mass, an increase in Z only causes a decrease in a star's temperature.

For a single star, this metallicity effect can only be taken into account by taking deep spectra and fitting theoretical models of stellar atmospheres to the observed spectra. Since all stars in a star cluster have the same Z, the metallicity of a cluster can be determined from a CMD by fitting various stellar-evolution models with a range of metallicities to the data and using the fact of the mass-dependent reddening and dimming of stars with increasing metallicity. Unfortunately, the precission of the underlying data has to be very high for this method. Due to the rather bad seeing conditions in Bonn we will refer in this experiment to literature values.

T2.7: *Draw two schematic CMDs in one plot, one of a globular cluster with an age of 11 Gyr and a metallicity of* $Z = 0.001$, *and the other of an open cluster with an age of 100 Myr and* $Z = 0.02$. *Include the main sequence, the giant branch and the horizontal branch (if applicable) in your sketch.*

2.4.5 Johnson-Filter System

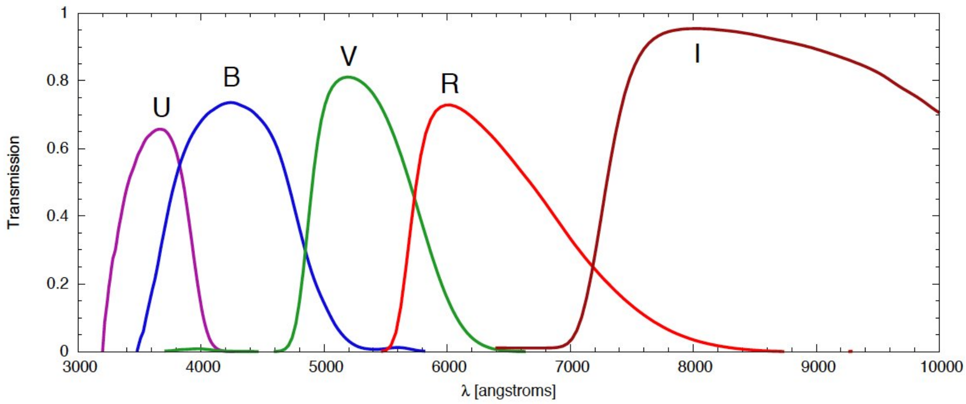


Figure 7: The transmission functions of the five available Johnson filters (picture taken from the University of Göttingen).

The most commonly used photometric system is the UBV-system based on the work by Johnson & Morgan (1953). The acronym stands for ultraviolet, blue and visual and denotes the wavelength coverage of the three most-often used filters. Meanwhile, this system has been extended to the infrared with the filters R (red), I (infrared) and

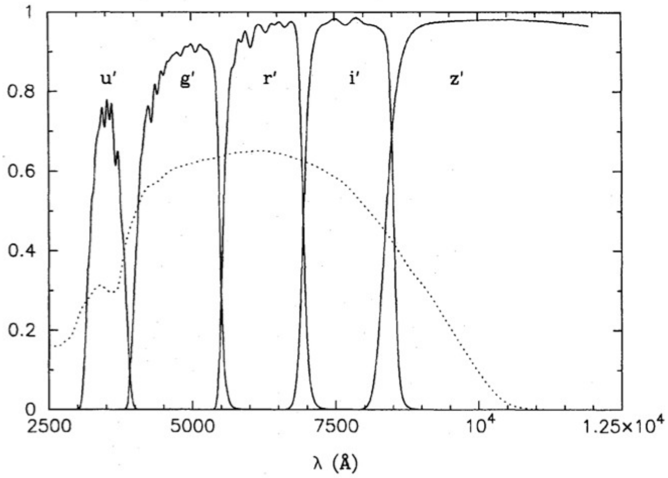


Figure 8: The transmission functions of the five Sloan filters (picture taken from Fukugita et al. 1996).

then J, H, K, L and M (for an overview of the corresponding wavelengths and filter characteristics see for example Binney & Merrifield 1998, p. 53). For the lab-course project the Johnson filters U, B, V, I and R are available (Fig. 7).

Note: Nowadays, the Johnson-filter system is being replaced in many applications by the Sloan-filter set (Fukugita et al., 1996) which has a similar wavelength coverage (u', g', r', i', z') but allows a higher transmission in each filter (Fig. 8) and therefore requires shorter exposure times. Due to this varity of available filters, conversions of observed fluxes into apparent magnitudes have to be done carefully. It is crucial that the measured fluxes of the reference stars in a given filter are compared to the listed magnitudes of the reference stars in the specific filter, i.e. if you observe with Johnson filters make sure the magnitudes of the reference stars you use are listed in the Johnson system and not in Sloan or anything else. The same holds for the theoretical isochrones which you fit to your data.

3 Basic Knowledge of Astronomical Observations

Almost all of our astronomical knowledge has been devised from the measurement of electromagnetic waves emitted from various forms of matter (e.g. stars, molecular clouds, etc.) in the universe. Due to the large distances from these emitters to us, the intensity of radiation we can measure is very small, thus sensitive equipments are needed. Nowadays for optical observations like the one you will do, radiation (photons) is collected and imaged by an optical telescope and detected (converted into electronic signal) by a CCD detector. This section tends to give you a basic introduction to the equipments you will use, and what you need to do to produce scientific data.

3.1 Telescope optics

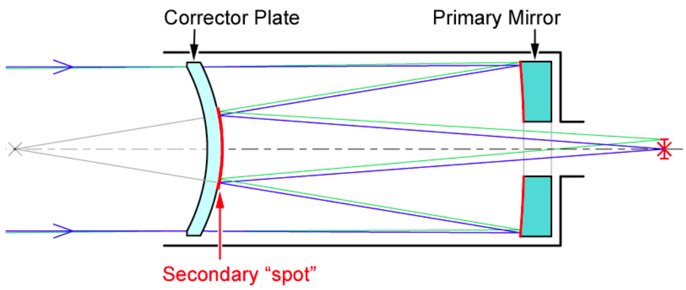


Figure 9: Light path in a Cassegrain reflector telescope. The red star indicates the Cassegrain focus where the CCD camera should be placed.

The telescope used for this lab course is a 50 cm Cassegrain reflector telescope placed in the dome on top of Argelander-Institut für Astronomie, Bonn. It has an f-number (f-ratio) of f/9 at its Cassegrain focus and f/3 at its primary focus. We will use the telescope here only in the Cassegrain focus.

T3.1: *Calculate the focal length of the telescope in both foci.*

3.2 CCD detector

CCDs (Charged coupled devices) are now widely used as photon detectors, both in the largest astronomical optical telescopes and in your pocket digital camera. A CCD detector consists of a two-dimensional array of picture elements (pixels), which are produced as a light-sensitive metal oxide semiconductor (MOS) capacitor on a silicon substrate. CCDs make use of the inner photoelectric effect to convert the distribution of photons to the distribution of electrons, which are then collected in capacitors and sequently read out: after an exposure is terminated, the collected charge is shifted column by column to a readout column by an alternating voltage impressed on the picture elements. The readout column is finally read out pixel-wise and the resulting signal amplified and converted to a digital signal by an analogue digital converter.

As a detector, a CCD has the advantages of:

- high sensitivity (high quantum efficiency of up to 90%),
- high dynamical range (the limits of luminance range that a detector can capture),
- linearity over almost the entire dynamical range,
- large spectral range (mid-infrared to ultraviolet for optical-optimised detectors),
- direct availability for further computer aided data analysis.

A less advantageous property of CCDs is the so-called dark current. At room temperature dark current brings CCD pixels to their saturation level within a minute or even less. Therefore astronomical CCDs are always cooled. Cooling is done thermo-electrically (temperature difference to ambient temperature: about 30°C), with closed cycle systems or by liquid nitrogen (CCD temperature can reach -100°C).

The CCD camera in use for the lab course is of type SBIG STL-6303E, which has 3072×2048 pixels, where the size of each pixel is $9\ \mu m \times 9\ \mu m$. It has a Full Well Capacity of 100000 e^-, i.e. in each pixel it can store 100000 electrons, and an ADU (Analog To Digital Converter Unit) gain of 1.4 e^-/ADU. It can be thermo-electrically cooled to about 30° below room temperature.

T3.2: *Calculate the field-of-view (FOV) of the telescope and the theoretical angular resolution. What limits the angular resolution during your observations?*

3.3 Observational conditions and requirements

There are some undesirable yet existing effects which will hinder you from getting good data if you don't take care of them correctly. Here we list those of them which will be encountered in this lab course. The treatment of these effects during data reduction will be done rather automatically using sophisticated software.

3.3.1 Seeing

The light distribution of a point source (e.g. a star) on the image plane is called point spread function (PSF). In the idealised case, a PSF is determined by the telescope aperture. With a normal circular aperture with diameter D, it is an Airy disc with an angular resolution of

$$\Delta\theta = 1.22\frac{\lambda}{D} \tag{10}$$

where λ is the wave-length of incoming light. However, when the blurring effect of earth's atomosphere is taken into account, the actual resolution is much worse for ground-based telescopes. And the PSF can be better described by a two-dimensional Gaussian. The size of such a stellar image can be described by the full width at half maximum (FWHM) of the Gaussian, which is called the seeing of the image. Naturally seeing measures the actual resolution in a particular observation, and it reflects the stability of Earth's atmosphere on a given night at a given location as seen through the telescope. The best seeing on the surface of earth is about 0.4". It's found at high-altitude observatories on small islands such as Mauna Kea or La Palma. For the condition above AIfA Bonn, a 2" seeing is already good.

To measure the seeing on your image, you can choose a non-saturated star and measure its FWHM in number of pixels using the telescope software, and then convert it to arcseconds using the relation:

$$\frac{\text{pixel size}/\mu m}{\text{focal length}/nm} \times 206 = {''}/\text{pixel} \tag{11}$$

The pixel size of the CCD and focal length of the telescope can be found in the above subsections.

T3.3: *Derive the factor 206 in equation 11.*

3.3.2 Focusing

If the CCD camera is out of focus you will get a big blob for each star on your image. So, before taking science frames, make sure you have the right focus by adjusting the focus until you get the sharpest image.

3.3.3 Linearity, saturation, dynamical range and exposure time

When speaking about linearity, "linear" means the increase in measured signal is propotional to the increase in the incoming photon flux. It is a desired property of the detector since it enables a direct measurement of the incoming photon flux. A CCD detector has good linearity over almost the entire dynamical range, but only over the dynamical range. When the detector reaches the saturation level, it is no longer possible to derive the exact number of photons that reached the detector originally. Therefore it is very important not to saturate the objects of interest by a too long exposure time. For bright stars a few seconds' exposure suffices to reach saturation. On the other hand one needs a long exposure to detect weak sources. And the longer the exposure time, T, is, the larger the signal to noise ratio (remember that S/N is proportional to $\sqrt{T}$). The selection of exposure time is thus dependent on the specific scientific goal of the observation. Only if a large number of objects of interest is not saturated, photometric analysis and calibration are possible. Thus, better make many shorter exposures of your object and co-add the images (see next section).

T3.4: *For a standard star of magnitude* $m_1 = 3.0$ *you get* 30000 *counts after an exposure time of* $t_1 = 10s$*. How long do you have to make an exposure for a star of magnitude* $m_2 = 5.0$ *to get the same number of counts?*

3.3.4 Cosmics and dithering

Two more artifacts will influence the quality of your data: "dead" pixels on the detector and cosmic rays. Dead pixels on images only represent noise, without any signal. Cosmic rays, on the other hand, lead to the saturation of several pixels around the impact position. While the number and positions of the dead pixels are fixed, the number of cosmic rays depends on the exposure time and their positions are random. These artifacts will hinder you from getting information on those pixels which correspond to certain positions on the sky. To avoid any loss of information at those positions, several exposures (6 is suggested here) for each scientific object are taken and the telescope is slightly moved of few arcseconds for each exposure. This method is called "*dithering*". The multiple frames are then aligned and the median of each pixel of the combined image is determined, with the consequence that all extreme values are discarded. This method also has the advantage that the signal to noise ratio improves without any saturation due to brighter sources.

3.4 Image Reduction Steps

3.4.1 BIAS subtraction

The signal of the CCD is first converted from an analog count signal (electrons in the pixel, i.e. a voltage value) into a digital number by an analog-digital converter. For example, with a 16 bit analog-digital converter you would transform your analog voltage values from the pixels into $2^{16} = 65536$ discrete levels, i.e. each pixel would get a value between 0 and 65535. For a better coverage of the available range of counts the logarithm of the analog signal is taken first. This is because you want to be very accurate for pixels with only a few counts but do not need to be that accurate for pixels with many counts. But, since the analog-digital converter can only handle positive values and fails for a value of zero, an offset has to be added electronically to every pixel value during the read-out process before the logarithm is taken. Otherwise, small voltages would lead to negative numbers thus would make the converter give out high positive values as the next value below 0 is 65535 in the example of a 16 bit converter. This so-called BIAS has to be recorded by a zero-time exposure called the BIAS frame, which then has to be subtracted from every image as the first step of data reduction.

3.4.2 Dark current subtraction

Even if the CCD chip is NOT exposed to optical light, there will still be a current flowing in it due to thermal fluctuations, which is called the dark current. Dark current is one of the main sources for noise in image sensors such as a CCD. The pattern of different dark currents in the pixels across the CCD can result in a fixed-pattern noise. Taking DARK frames and subtract them from the science (and FLAT) frames can remove an estimate of the mean fixed pattern, but there still remains a temporal noise, due to the fact that the dark current itself has a shot noise.

Note that the level of dark current is strongly dependent on the temperature of the CCD chip and the length of the exposure. For a liquid nitrogen cooled CCD camera, the dark current can be neglected for many observational purposes. But in this lab course it has to be taken into account. Thus, you should also make sure that the CCD temperature and exposure time for the DARK frames match those of the light (science) frames.

3.4.3 Flat fielding

You also need to take some FLAT frames by making exposures towards a uniformly illuminated background (e.g. twilight sky or a carefully constructed and illuminated dome flat field screen, currently we use the former). However, your obtained image will be far from uniform. On your FLAT frame you will probably see "donuts" which represent dust grains somewhere within the light path, and other variations of light due to the telescope optics. Also, the response of the CCD is not exactly the same from pixel to pixel, such pixel-wise variation is also recorded in a FLAT frame. The exposure time for the FLAT frames should be determined such that the peak brightness level of a FLAT frame is 1/2 or 2/3 of the saturation level (saturation = full well in e^-/ADU gain). And for each filter used for science exposures, a new FLAT frame should be made. This is because the pixel response to incoming light is wave-length dependent. During data reduction a FLAT frame will be normalized to an average value of 1 and used to divide the science image.

T3.5: *What can you infer from the sizes of the donuts?*

T3.6: *How many counts do you expect your FLAT frames to have?*

3.4.4 Masking/weighting

Since one pixel in your CCD might be more sensitive with respect to another one (which always happens), you would like to trust the sensitive one more since it gives you data with higher S/N. This can be done by assign individual weights for every pixel, a process called weighting. The weighting factors can directly be taken from the normalised FLAT. Weighting is also neccessary when you try to co-add frames (see co-adding): your object on each frame may be at different positions (usually the case, see dithering), when co-adding them, a proper weighting ensures maximised S/N in the final image.

For those bad pixels/columns on the CCD chip and cosmic rays, a simple way of treatment is to "mask" them using softwares and assign them less weight. There are two types of masks: global and individual. One global mask is made for a particular CCD detector and can be applied to all images produced by it. One can also create an individual mask for each image, counting also for cosmic rays.

3.4.5 Astrometric calibration

Your obtained image is a telescope-configuration-dependent projection of a curved sky. Thus the pixel/detector coordinates and the sky coordinates do not have a simple relation. To project your image back onto the sky coordinates is called astrometric calibration. This can also be done by softwares, with the help of a reference catalogue.

3.4.6 Sky subtraction

During exposures the CCD not only collects light from your target of interest, but also receives radiation from the background sky. In addition there will be ADU counts from unresolved objects and glow from objects not even in your field of view (e.g. a bright star nearby, or the moon!). By determining this background level and removing it from your image, only the source flux will remain. The usual way of modelling the background is to first remove all objects in your frame and then smooth the image with a specific kernel width. Then this background image can be subtracted from the original frame.

3.4.7 Co-adding

By stacking all science images into one and making sure that each object falls onto the same pixel, the final image will have a higher S/N value than each image alone. This process is called co-adding.

3.5 Photometric calibration, reference stars

For the same star on the sky, observations through different telescopes under different weather conditions will yield different fluxes. After converting to magnitudes, one obtains a particular instrumental magnitude which does not directly reflect the true magnitude of the star. The most simplistic solution is to assume that you only have an offset Z between your observed magnitudes and the true ones.

$$m_{calib} = m_{instr} + Z \tag{12}$$

Usually you can calibrate your instrumental magnitudes by observing some reference stars or "standard stars" whose calibrated magnitudes are well known and stable. Observe such stars using the same configurations as your scientific objects. For this lab course project it is sufficient to look up the magnitudes in the different filters for some reference stars in your chosen clusters from a catalogue, or compare it to a well calibrated CMD of this cluster from the literature and add appropriate offsets.

4 Observations

After reading this section you are supposed to know what you need to do during your night-time observations, step by step. Furthermore, you should choose your objects to be qualified for the observations. Feel free to discuss with your tutors if you have questions.

The data files you should obtain during your observations are:

- 15 FLATs in each of the 2 filters (B and V),
- 15 BIAS frames,
- 10 DARK frames,
- 10 SCIENCE exposures in each of the 2 filters for each object,
- *optional:* observe FLATS and SCIENCE for the R band too, if you would like to have a nice RGB (color) image of your star cluster!

4.1 Choosing Your Objects

As a first exercise, choose yourselves the star cluster to observe. Your choice should base on the "visibility" of the objects during the night of observations, i.e. their tracks across the sky. It would be desirable to have the object as high (in altitude) as possible during the time of observations. One other factor to take into account is the position of the moon. You wouldn't like to have your object to be close to a full moon since a bright moon would significantly increase the level of your sky background and thus decrease the signal-to-noise ratio (S/N) of your object.

One convenient way to visualize the tracks of the objects during one night is to produce a visibility plot. Make one yourself according to the instructions in the task below, figure out its meaning, and choose your objects basing on it.

T4.1: *Generate a visibility plot for the date of your observation. Procede as follows:*

- *choose an object at a fairly high altitude during the time of your scheduled observation,*
- *go to http://catserver.ing.iac.es/staralt/ ,*
- *select mode: Staralt; Date: your expected date for observation,*
- *specify the coordinate of AIFA, Bonn: 07 04 01 50 43 46 [75],*
- *include "Moon distance" in "Options",*
- *generate one plot for open clusters and one for globular clusters.*

Which direction does the peak of one object track correspond to?

T4.2: *Pick two objects (one OC and one GC) according to your visibility plots.*

4.2 Observing Schedule

Since time is short during the night every step has to be planned in advance. Therefore, an observing schedule has to be prepared beforehand. For details on the tasks which have to be performed during the night see the observer's manual at the telescope.

1. Telescope and CCD camera setup (your tutor will help you with this step).

   open the dome

   remove the mirror covers

   attach the CCD camera to the telescope; the power and data lines

   launch the necessary software for telescope and camera control

   synchronize telescope/camera and computer

2. Cool down the CCD to about 25°C below the ambient temperature (set, e.g., -30°C in CCDsoft).

3. FLAT frames (15 FLATs/filter). Note that there's very limited time in which a SKY FLAT (FLAT frame pointing the evening/morning twilight sky) can be taken. So prepare everything beforehand and hurry up!

4. Focusing by means of the *Bahtinov mask* (remember to focus every time you start to observe your target with a different filter)

5. Science frames of your object: 10 science frames/filter, apply dithering in between (only a few arcseconds). Make sure that your binning (for all frames is set on *1x1*)

6. BIAS and DARK frames (15 and 10 respectively). They can be taken at any time during the night. The exposure time of DARK frames should be the same with the science frames. **Note:** depending on the telescope software it may be possible to automatically take DARK and BIAS frames with each exposure and automatically subtract them from the science frames. This may be convenient but is not always recommendable.

4.3 Data storage

Since you will take a number of frames during the night it is recommended to create an appropriate tree folder on the local hard drive of the telescope beforehand. The data reduction software THELI requires a certain folder structure which you should stick to from the very beginning to avoid confusion. This structure looks as follows:

- The top folder name should have the date of observations in it, e.g.
  `.../2012.02.29/`

- There should be a separate folder for any BIAS frames, e.g.
  `.../2012.02.29/BIAS/`
- If you take DARK frames, then you also need a separate folder, e.g.
  `.../2012.02.29/DARK/`
- Each set of FLAT frames for a specific filter should have a separate folder, e.g.
  `.../2012.02.29/FLAT_B/`
  `.../2012.02.29/FLAT_V/`
- Each set of science frames for a specific object and filter gets its own folder, e.g.
  `.../2012.02.29/M92_B/`
  `.../2012.02.29/M92_V/`

- *Suggestion:* if for any reason you realize that one of your frames has some problem and will not be used in your analysis (e.g., saturated FLAT frame, moved SCIENCE image, etc...), note down the frame number so that you can discard it (move it in a "discarded" folder, for example) as soon as you have time or right before your data reduction. Of course: take a new frame in order to substitute the bad one.

Also name your files properly, otherwise you will get confused on the next day or whenever you work with the data again.

5 Data Reduction

For reducing the data and extracting a catalogue of sources out of the pictures, a number of steps are necessary and a variety of scripts and programmes will be used. In this section we will introduce the tools you need for creating a color-magnitude diagram, this will be done in the order in which you will have to use them.

5.1 SKYCAT/DS9

You can check your own images and compare them with the virtual-telescope data by using Skycat or DS9. For the first one type
`> skycat`
in a terminal window. Open any image with the menu File/Open and select the the file you want to open. To adjust the view of the image push the button Auto Set Cut Levels first and then go to the View/Colors menu. In the new window you can choose various color scalings, color maps and intensity functions. Pick for example Linear, heat and gamma for a good representation of your star cluster (Fig. 10). With Skycat you can also access on-line catalogues like DSS. Simply choose an appropriate catalogue from the Data-Servers menu.

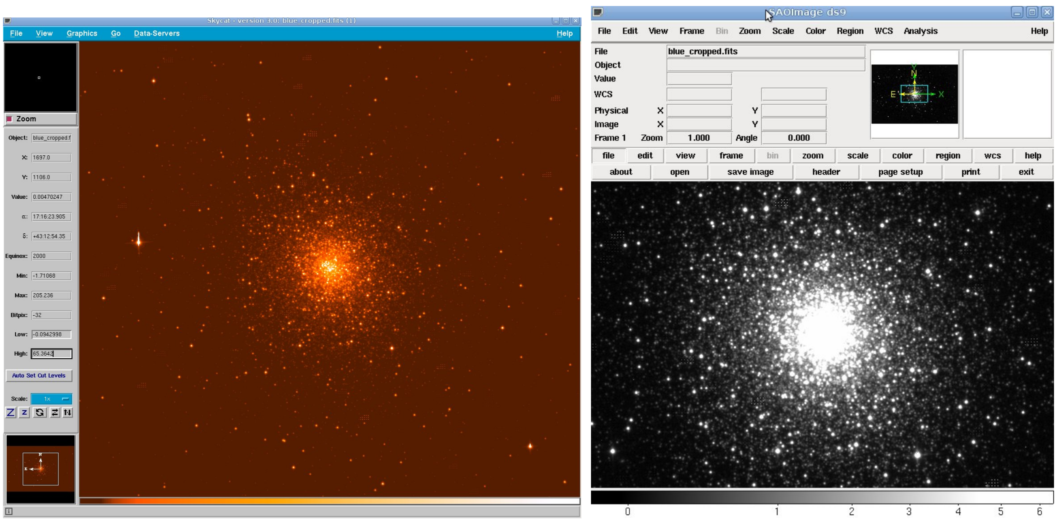


Figure 10: Main windows of Skycat (left) and DS9 (right).

DS9 is similar to Skycat but offers slightly different functionality. To start DS9 type
`> ds9`
in a terminal window and the DS9 window will show up (Fig. 10). As in Skycat, open any file with the File/Open menu. The scaling of the view can be adjusted via the Scale and the Color menu. Choose, for example, Zscale and Grey. You can now easily adjust the scaling by holding the right mouse button and moving the cursor up and down or left and right. If you want to access on-line catalogues, you can find them in menu at `Analysis > Image servers`.

5.2 THELI

The data reduction itself will be done with the graphical-user-interface (GUI) version of THELI, a freely-available software package for the reduction of astronomical imaging data which was in part developed at AIfA[1]. An introduction to THELI is given in the appendix of this document, therefore we only give the necessary steps here but strongly recommend going through the appendix first.

Before you start working on your data make sure you have made a backup of the whole data set. Also ensure that all files are stored in appropriate folders as described in Sec. 4. Throughout the reduction we recommend to follow the processing of your files by checking the specific folder content after each reduction step and following the renaming of the files. This might be very helpful in case a reduction step fails or has to be redone or undone. Note that THELI is not fail safe and needs the caution of the user. Note also that *not all reduction steps have to be repeated for each filter*, e.g. BIAS processing, and might even lead to errors when repeated.

Start THELI by typing
`> theli`
in a terminal window and the main window will show up which has seven panels which have to be gone through from left to right. Thus, in the following sections we will go through each of the seven panels in the correct order. Do not try to launch more then one THELI session at once.

5.2.1 Initialise

In the initialise window of THELI enter an appropriate name for the data set you are reducing into the "Current LOG file" line, a new log file will be created. Start with one filter and don't start reducing a second filter set before you finish the first. Just to make sure that no old preferences are set, hit the "Reset" button next to the LineEdit afterwards.

For the given filter enter the directory names into the appropriate lines. Make sure the "Main path" is an absolute path and not a relative one, i.e. that it gives the full path starting from root.

Finally, before proceeding to the next panel, choose the camera you used from the instrument list. The camera that you used for this lab course is the ST6303: you can find it clicking on the "user defined" box, above the list. Then select *STL6303@AIfA*. **Note:** There is another existing *STL6303* instrument file in the "Commercial" list, but pay attention to do NOT select that one!

[1]`http://www.astro.uni-bonn.de/~theli/`

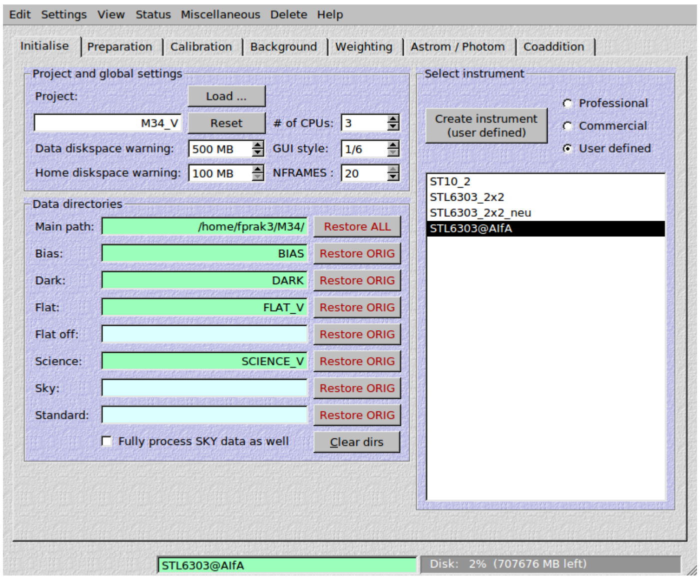


Figure 11: The initialise window of THELI.

5.2.2 Preparation

In the preparation window you have to check the “Split FITS / correct header” box and hit the “Start” button. Note that this task has to be done for the BIAS and DARK frames just once as they are used for all filters. If you go through this task for a second set of pictures for a different filter, THELI will automatically detect your “master BIAS” and “master DARK”, so it will skip the splitting for these frames.

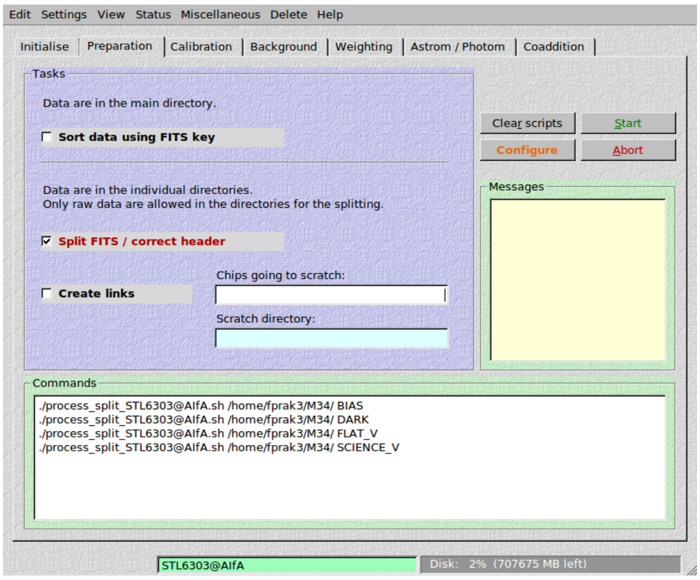


Figure 12: The preparation window of THELI.

5.2.3 Calibration

In the calibration window first check the "Process biases", "Process darks" and "Process flats" boxes and then hit "Start". Note again, if you already processed BIAS and DARK frames, don't check again the first two boxes.

Finally, you have to calibrate your data. Therefore, check also the "Use DARK" box (for those cases where you need a dark current subtraction, i.e. long exposures).

5.2.4 Background modelling

There is no need for background modelling in this lab course project, but if you think the background variations in the final coadded image are too large, you may go back to this step (and redo the follow up tasks).

In that case you may start with parameters similar to the ones in figure 26, but select "Subtract model" as application method in group (4).

5.2.5 Weighting

In the weighting window you have to run the "Create global weights" and "Create WEIGHTs" process.

5.2.6 Astrom / Photom

The Astrom / Photom window of THELI contains essential steps of the data reduction process of your data which have to be done with care. First of all, find the coordinates of your object, e.g. by DS9, and enter them into the RA and DEC fields. (This may be necessary, if the pointing of the telescope is significantly offset from the true pointing.) Doing this may require you to confirm in a dialogue box that you want to overwrite the existing header information. *Make sure that you use the same coordinates every time.*

Then you have to download an astrometric reference catalog. Therefore, select "Web (France)" and "2MASS". (In many cases "SDSS-DR9" (Sload Digital Sky Survey, data release 9) can be the better choice, but the galactic plane is hardly covered by this survey). Make sure the radius of your requested catalog covers a few arcsec more than your field-of-view, choose an appropriate magnitude limit (e.g. 15) and hit "Get catalog". Sometimes the query returns an error message. This either means that the server is temporarily not available (just wait a bit and try it again) or that your target is not within the coverage of this catalogue (try a different one, e.g. "USNO-B1"). Also note, that the larger your radius and the fainter the limiting magnitude, the more sources you will get in your catalogue. Don't have too many sources in this catalogue as the astrometric solution will get less accurate with more such degrees of freedom.

Check the "Create source cat" box and push the "Configure" button. In the create

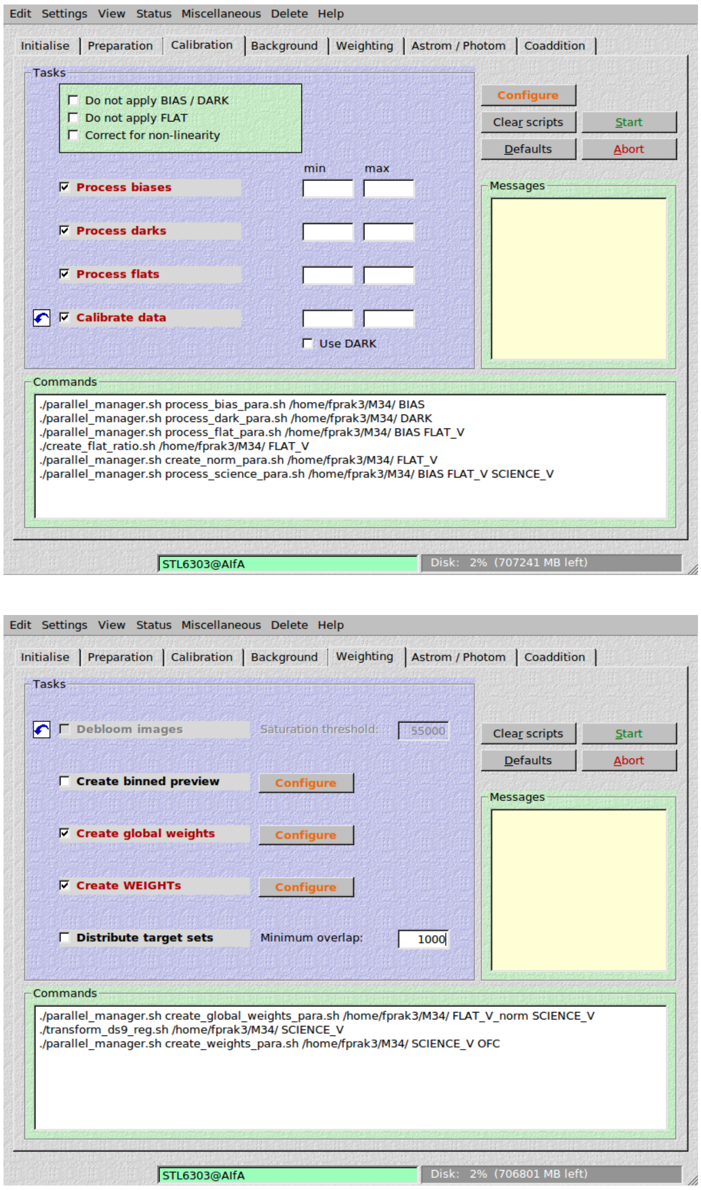


Figure 13: The calibration and weighting windows of THELI.

source catalog window (Fig. 15) set the DETECT_THRESH and the DETECT_MINAREA parameters both to 5. If your field contains a huge number of objects, you can rise DETECT_THRESH to 10 or to an higher value. Starting with (5,5), you can switch to a higher combination, such as (10,10), (40,15) or (100,20), in case THELI finds no astrometric solution in the following reduction step. For very noisy data with a low S/N ratio you can try (1,5) first. Then configure the astrometric and photometric solution. Therefore, check "Astro + photometry", select "Scamp" from the drop-down menu and enter the "Configure" menu. In the astrometry configuration window (Fig. 15) and try the following values: POSANGLE_MAXERR 5, POSITION_MAXERR 5, DISTORT_DEGREES 1. Also, check the box "Match Flipped Images". Finally start the calibration processes. If THELI finds no astrometric solution, try creating a new source catalogue with different parameters as stated above or change the two MAX-ERR parameters to about 10.

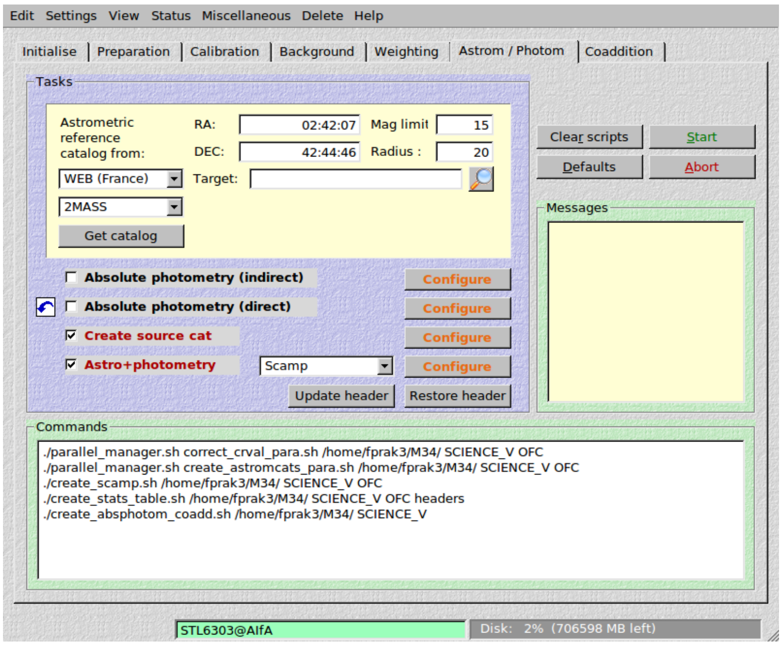
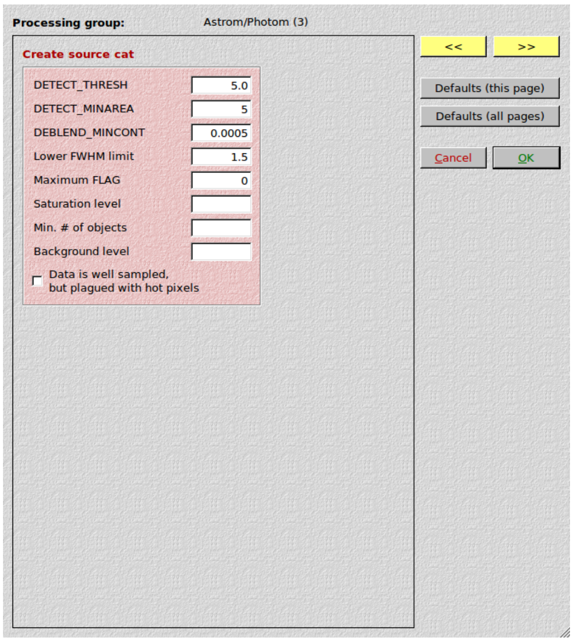


Figure 14: The astrometry/photometry window of THELI.



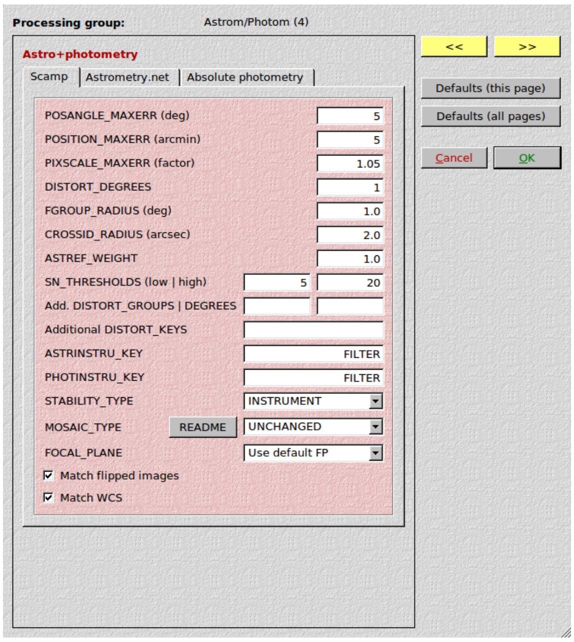


Figure 15: The create source catalogue window (left) and the astrometry configuration window (right) of THELI.

5.2.7 Sky subtraction

In order to create a nice RGB color picture in the end (optional), it is useful to subtract the remaining sky background. Most likely you can run the task in the coaddition window with default parameters. Check the parameters and make sure that DT (detection threshold) is low (1.5-3) and the kernel width large (¿ 100) and tick the “Sky subtraction” box.

5.2.8 Coaddition

In the coaddition window (Fig. 16) first start the sky subtraction by checking the corresponding box and push the “Start” button. Then check the “Coaddition” box and enter the “Configure” menu. In the configuration window (Fig. 17) enter the coordinates which you also entered in the previous reduction steps. Furthermore, enter an identification string into the specifc box. The name should be clear, e.g. blue, green, red or B, V, R. Also, check that the pixel scale (of the CCD), is set to 0.4 for the ST6303 camera. Then, fill the “Sky postition angle” field clicking on the “From Image” button. If a weird number (as -999) appears, it means that something went wrong in the astrometric calibration...check it out! If everything is fine, exit the configuration menu clicking on “OK” and start the coaddition. After this step you are done with the given filter. Proceed with the reduction of the next filter before you go to the next section.

After reducing all available filters you should make a new folder, e.g. `[...]/ALL_COADDS/` and copy all the *coadd_XXX* folders into this folder.

5.2.9 Prepare color picture

If you have taken exposures in three different filters (B, V, R), you can proceed with this step. After the reduction of the images you have to make sure that all pictures cover the same field. This is usually not the case. Therefore we crop the borders of the images such that all images have the same size and show the same part of the sky. This can be easily done with the “Prepare color picture” routine of Theli which you find in the top bar under “Miscellaneous”.

In the “Create color picture” window (Fig. 18) write the location of your folder with the coadded images (created in the previous reduction step) and hit “Get coadded images”. You can then find the cropped coadd and weights image in a subfolder in the same folder, where the input images are located.

If this step was successful your are done with the data reduction with Theli. The next step will be to extract the final source catalogue from the coadded and cropped images. This will be done with SExtractor in the next section.

But before proceeding with the next step you can create a color picture of your object if you have managed to take images in three filters. Therefore go to the “Color calibration” tab of the “Create color picture” window and select the three images in the drop

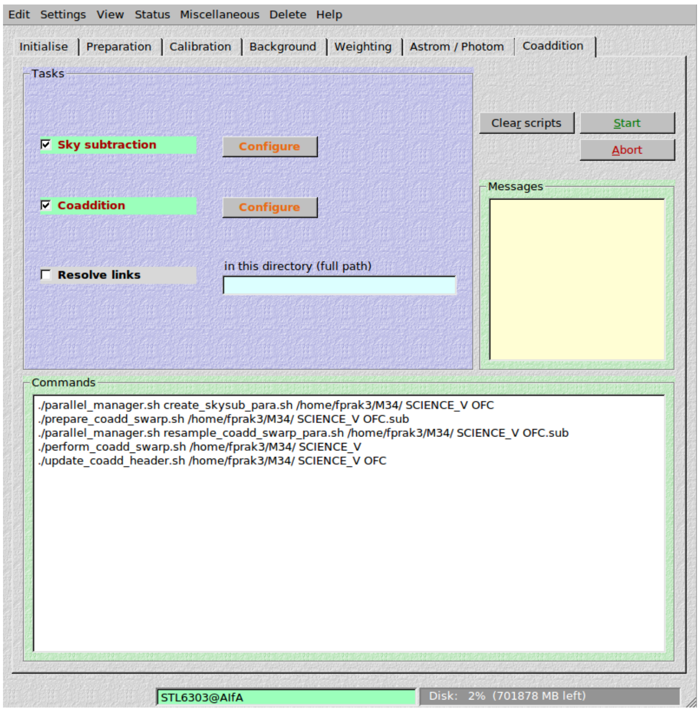


Figure 16: The coaddition window of THELI.

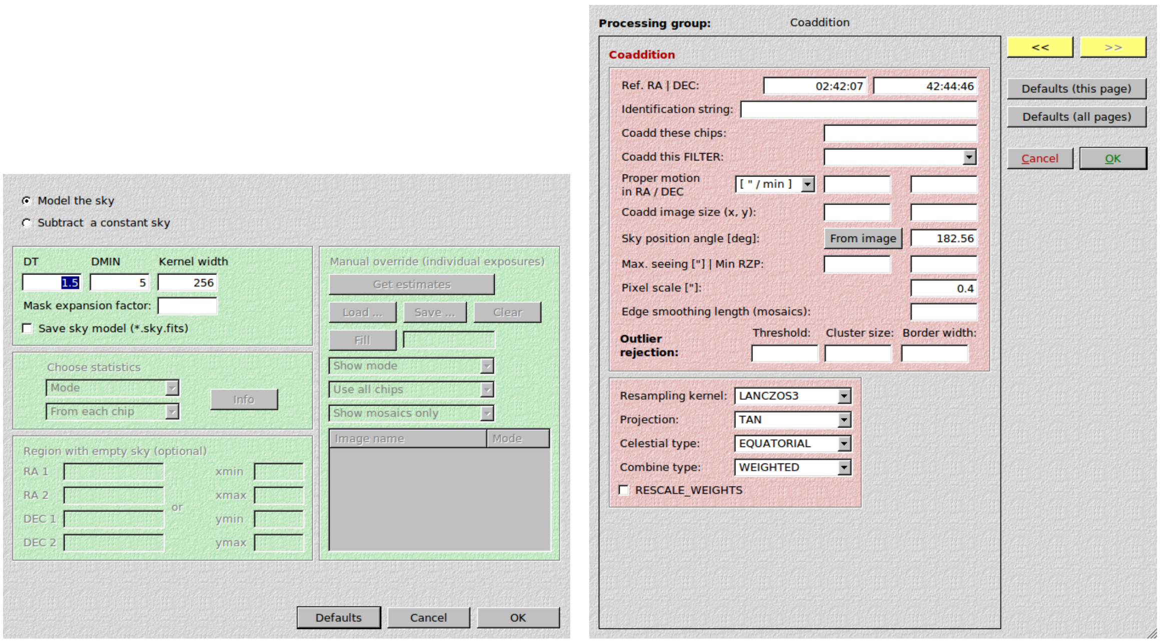


Figure 17: The sky subtraction configuration (left) and coaddition setup window (right) of THELI.

down menus (matching them in a meaningful way to the RGB color channels) at the top (Fig. 19). First try to let THELI create the color picture automatically. Therefore select one of the four calibration options and hit "Calibrate" and then "Preview". If the colors of the final picture are not according to your taste hit the “Reset” button and try the other approaches or edit the weight factors manually. The result will be stored in `preview.tif` in your current working directory.

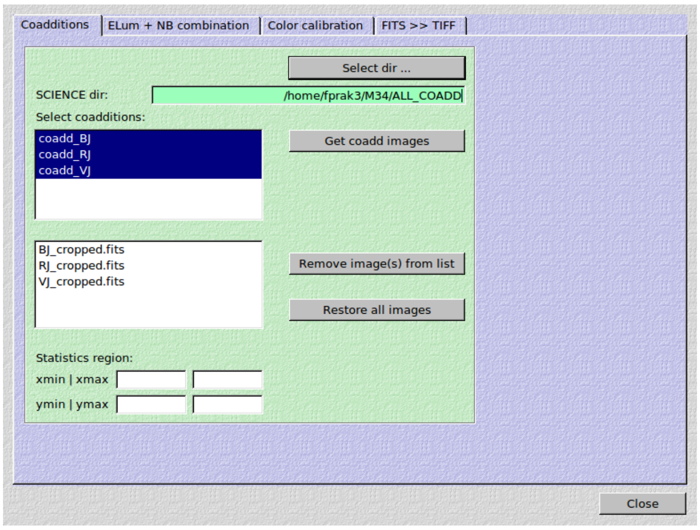


Figure 18: The create color picture window of THELI.

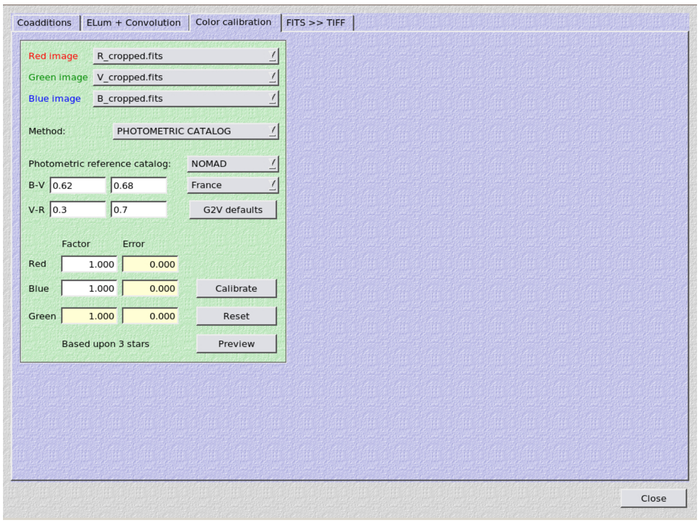


Figure 19: The create color preview window of THELI.

5.3 SExtractor

The final extraction of the sources will be done using SExtractor[2], a program that builds a catalogue of objects from an astronomical image. Although it is particularly oriented towards reduction of large scale galaxy-survey data, it can perform reasonably well on moderately crowded star fields. Thus, it may lead to bad results in the centres of globular clusters, but will suffice for our purposes.

Here, we will extract the sources with a small home-made shell script that makes use of the SExtractor routines. Therefore, create a folder, e.g.
`../2014.02.29/M92_CATALOGUE/`
and copy the following files into this directory:

- `get_catalogue.sh`,
- `get_catalogue.assoc`,
- `get_catalogue.makessc`,
- the cropped images,
- the cropped weights.

Run the script by typing the following command in a terminal window:
`> get_catalogue $PATH $B $V $R`
where $PATH is the full path of the current folder, $B is the name of the blue cropped image, $V is the green cropped image and $R is the third, in case you have made observations in a third filter. Otherwise simply repeat argument $V, in this way you will get an "R" column in the catalogue which is redundant, since it repeats the magnitude of the V-band.

The script will create source catalogues, just as was done in Theli, for each filter and cross reference the three catalogues. Only sources which get detected in all three filters will be written to the resulting ascii catalogue which will be located in
`../2014.02.29/M92_CATALOGUE/result/result_ascii.cat`
where the columns are:

1. right ascension (J2000) in degrees,
2. declination (J2000) in degrees,
3. B magnitude,
4. V magnitude,
5. R magnitude.

[2] `http://www.astromatic.net/software/sextractor`

You additionally get catalogues for each filter and the combined catalogue in an enhanced file format (LDAC) comparable to the .fits format for images, which you don't need here any more, but which you should keep in mind in case you ever need to deal with more sophisticated catalogues.

You can quickly count the number of detected sources by typing
`> wc -l result_ascii.cat`
within the result directory. The number of detections depends on the parameters DETECT_MINAREA and DETECT_THRESH just like we used them in THELI. You can set those parameters by editing the script with a text editor or by simply adding the two numbers to the command line argument:
`> ./get_catalogue.sh $PATH $B $V $R $DETECT_MINAREA $DETECT_THRESH`
where you should choose one of the pairs (5,5), (10,10), (15,40) or (20,100). You can play with the values and compare the resulting numbers of detections. The larger values will give you less false detections but also much less sources. You have to decide, which detections you will base your analysis on.

6 Analysis

After extracting the catalogue from the observational data you can begin the analysis. Aim of the analysis will be to determine the most basic parameters of the clusters, the distance, the age and metallicity. We recommend to do the analysis with GNUPLOT since it is a very convenient tool for such tasks as plotting data from many different files, however if you have any other preference, feel free to use your favourite plotting tool. For help on the functionality of GNUPLOT we recommend some on-line help pages[3][4].

6.1 Calibration

Begin your analysis by drawing a first color-magnitude diagram. Start GNUPLOT in a terminal window by typing
```
> gnuplot
```
Plot the color index B-V versus the V magnitude with the following command
```
> plot 'result_ascii.cat' using ($3-$4):($4)
```
where the $3 and $4 are the columns in the file. Note that the magnitudes in your catalogue are not calibrated yet. This has to be done by hand, by comparing the measured magnitudes of a few stars with reference magnitudes of on-line archive data and applying a correction constant to each magnitude. You can do that, for example, with
```
> plot 'result_ascii.cat' using (($3+0.3)-($4+0.1)):($4+0.1)
```
where the new numbers (+0.3 and +0.1) are the corrections. For the calibration of the magnitudes you will either use the magnitudes of the reference STANDARD stars which you have eventually observed seperately or your tutor will provide you with the reference magnitudes of some (non-variable) stars in your field-of-view. For the latter, identify the reference stars in your catalogue by their right ascension and declination and calculate the correction constant for each filter. If the quality of your data does not allow for this method your tutor will provide you with a literature CMD of this specific cluster such that you can shift your magnitudes accordingly. Now you can draw a calibrated CMD. What is the limiting magnitude of your observations?

6.2 Extinction

Extinction by interstellar dust makes stars appear dimmer than they are. This can be corrected for by adding a constant factor to each magnitude. In principle these factors can be determined through a color-color diagram in which two color indices are plotted versus each other. But therefore you need reliable images in three colors and sophisticated analyses. This is hardly possible with the data taken from Bonn due to light pollution, thus we will stick to values listed in the literature. Often the extinction factor is smaller than the achievable accuracy in magnitudes which can be achieved from Bonn. Thus, assume an extinction factor E(B-V) of zero, if not stated otherwise by your tutor.

[3] http://www.gnuplot.info/faq/faq.html
[4] http://t16web.lanl.gov/Kawano/gnuplot/index-e.html

6.3 Distance Determination

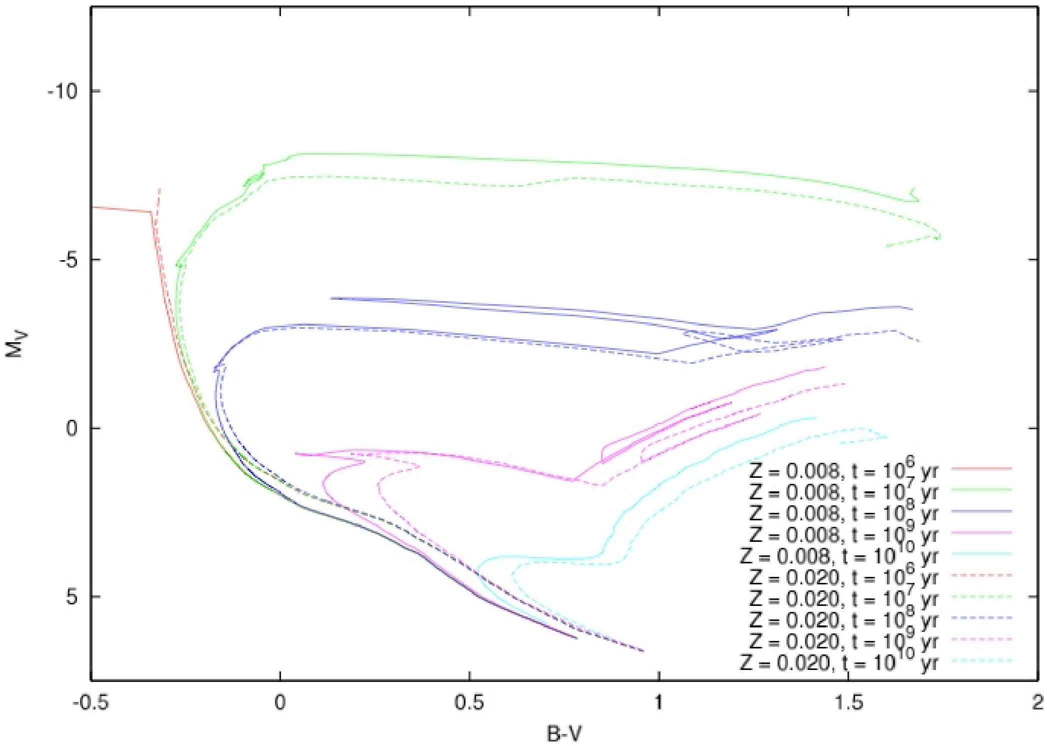


Figure 20: Isochrones for high and low metallicity stellar populations. The main sequence of young clusters is almost independent of metallicity for $B-V < 0$. For old clusters (older than 10^9 yr) the horizontal branch is independently of metallicity located at +0.5.

The distance to star clusters can be determined in different ways, e.g. with variable stars or supernovae. With our CMD we can also derive the distance since there are stars in each cluster with well defined absolute magnitudes. By comparing the observed magnitudes of those stars with their theoretical absolute magnitudes we can evaluate the distance modulus (eq. 8) of the cluster. But the stellar populations of open clusters and globular clusters can differ significantly, such that not every method can be applied to every cluster.

6.3.1 Main-Sequence Fitting

From Fig. 20 we can see that open clusters of young age have a well defined main sequence for high mass stars, which does not vary significantly with metallicity. Based on this part of the main sequence we can estimate the distance without regard to the age and metallicity of the cluster. Therefore take the youngest isochrone of one of the isochrone directories, e.g.

```
isochrones/geneva/008/iso_c008_0600.UBVRIJHKLM
```

and plot column 9 (B-V) versus column 7 (absolute V magnitude) into your open cluster CMD, i.e.

```
> plot 'result_ascii.cat' using (($3+0.3)-($4+0.1)):($4+0.1),
```

`'~/SCPHOT/geneva/008/iso_c008_0600.UBVRIJHKLM' using ($9):($7) with line`
Apply a constant shift to the V magnitudes such that the high mass part of the main sequence fits the observations. From this shift you can calculate the distance d by using

$$d = 10^{0.2(m-M+5)}\text{pc}. \tag{13}$$

6.3.2 Position of the Horizontal Branch

Old globular clusters do not show an extended main sequence (Fig. 20). But, as stellar populations evolve, massive stars move onto the horizontal branch where they burn helium in the core. The absolute magnitude at which they shine in this stage is well known and can be used to determine the distance to old star clusters.

M_V of horizontal branch stars is approximately +0.5 mag. Thus, identify the horizontal branch in your CMD, shift your V magnitudes accordingly and apply equation 13.

6.4 Isochrone Fitting

After we have produced a well calibrated CMD we can fit a theoretical isochrone to the data and see which parameters reproduce the cluster best. For this task we use the Geneva database of isochrones (Lejeune & Schaerer, 2001), named after the university of origin. Note that these isochrones are based on simplified assumptions and that they obey an intrinsic uncertainty. There are other databases of isochrones, coming, e.g., from Padova (Nasi et al., 2008), which in some cases differ substantially from the Geneva models. For simplicity we stick to the Geneva database here.

The available set of isochrones covers metallicities (in solar abundances) of $Z = \{0.001, 0.004, 0.008, 0.020, 0.040, 0.1\}$ and ages of 10^3 yr to 15 Gyr in a reasonable step width. The data is stored in the folder `geneva`. This folder contains subfolders for each metallicity and in every subfolder are separate files for each age. For example the file
`geneva/001/iso_c001_0950.UBVRIJHKLM`
contains the isochrone for $Z = 0.001$ and for an age of $10^{9.5}$ yr. The columns in the files are always the same. Use column 9 for B-V and column 7 for the absolute V magnitude. For each cluster go through the data set and compare the theoretical isochrones with your data. For young clusters you should stick to high metallicities whereas for old clusters you should apply low metallicities.

When you have a good fit, you can change the output to a file instead of the screen and give an output name. Afterwards you must change to the screen again, though:

```
> set terminal eps postscript enhanced
> set output 'figure1.eps'
> plot 'result_ascii.cat' using ($3-$4):($4),
'~/SCPHOT/geneva/008/...'  using ($9):($7) with line
> set terminal x11
```

Our aim is it to understand anything in our CMD, since there should be an explaination for any position of any data point in the graph. Therefore, answer the following questions:

- What age has the specific cluster?
- How large are the uncertainties in this determination? How can you estimate reasonable uncertainties?
- What influences the width of the main sequence, i.e. why is the main sequence not as thin as the isochrone predicts? Think about how the isochrone is produced, and what is assumed in this process.
- Can you identify stars that lie beyond the turn-off point, i.e. are bluer than the turn-off point but still appear to lie on the main sequence? How can this happen?
- How many stars can you identify in this way to appear abnormal in age/color and what does it tell you about the stellar population of this cluster?

Do this for all clusters for which you have taken data and for each additional catalogue which your tutor hands out to you.

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A THELI user manual – lab-course short version

General description of the GUI and its elements

Main elements and nomenclature

The main window consists of two main parts:

- a *menu bar* at the top of the window and
- a dialogue with a *tabbed menu*.

We will refer to the seven tab menues of the latter as *processing groups* (*PG* for short). These are

- *Initialise*
- *Preparation*
- *Calibration*
- *Background*
- *Weighting*
- *Astrom / Photom*
- *Coaddition*

and contain the various *reduction steps* or other settings. The data is reduced by making some initial settings in the first processing group (tell the software where the data is, which instrument was used, etc.), and then one works his way through the remaining six PGs.

Integrated Help system

A very extensive help layer has been integrated that comes in various active and passive forms.

- On the lowest level there are simple *tool tips* that are displayed when you hover with your mouse button over a specific PushButton or other element unless the meaning of those is obvious (see. Fig. 21).
- More extensive help for the various PGs and reduction steps is available under *Help→What's this?*. The cursor will change to a question mark. Move it anywhere into a GroupBox with the reduction steps to obtain general information of what is happening in this particular PG. Or click on the CheckBox of a particular reduction step to obtain more detailed information for this task, such as if this step is mandatory or optional, and if you have to provide any parameters.
- The *Help* menu provides you with further support, such as a dialogue containing an overview of the functionality of the various GUI elements. Furthermore, you can access *this* document as well, the general pipeline documentation or a rather technical paper analysing the performance of *THELI* .

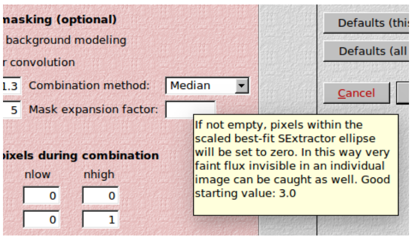

Figure 21: A tooltip appears whenever you hover with the cursor over certain GUI elements.

- If the GUI expects some parameters as input it will highlight the according fields for you with a red background colour etc. The colour coding is explained in detail in the following subsection.
- Each reduction step you run dumps all programme and script output into the SCRIPTLOG. You can access these, for each PG separately, under *View* in the menu bar.

A.1 The INITIALISE processing group

A.1.1 Pipeline Settings

Project and global settings It is useful to enter a meaningfull project name, which will be used to name the LOG files. Imagine you start working on a new R-band data set of NGC 1234, then NGC1234_R would qualify as a nice project/LOG name. Enter it in the corresponding field, then click on the *Reset* PushButton.

This will flush the GUI from all settings that might be left over from a previous reduction run. Parameters in the GUI and the LOG are set to meaningful default values. If the LOG entered does not yet exist, it is created at this moment. It is automatically updated (or created, if not yet existing) if you switch to any other PG, or shut down the GUI. The LOG will contain all reduction steps, parameters and GUI settings you have done or chosen for a particular data set. Choose a new LOG name if you reduce a different data set. If you leave this field empty and start processing anyway, the LOG will be named `noname`.

Whenever you launch *THELI* , it will read the LOG that was used last, updates all GUI elements and internal variables correspondingly and switches to the PG that was active when you closed your last *THELI* session. You can continue with your reduction at the point where you left it the last time.

LOGs are usually stored in ∼`/.qt/` and linked to ∼`/.theli/reduction_logs/`. The previous path may vary depending on your *Qt* installation. You will never have to touch a LOG file, apart from loading an old one into the GUI.

A.1.2 Data Settings

Select instrument Select the correct instrument to tell *THELI* how to handle the input files (`ENZIAN_CAS@HOLI_1M` for the lensing data, switch to *user defined instruments*, `STL6303@AIfA`, if you are reducing data from the AIfA rooftop telescope.)

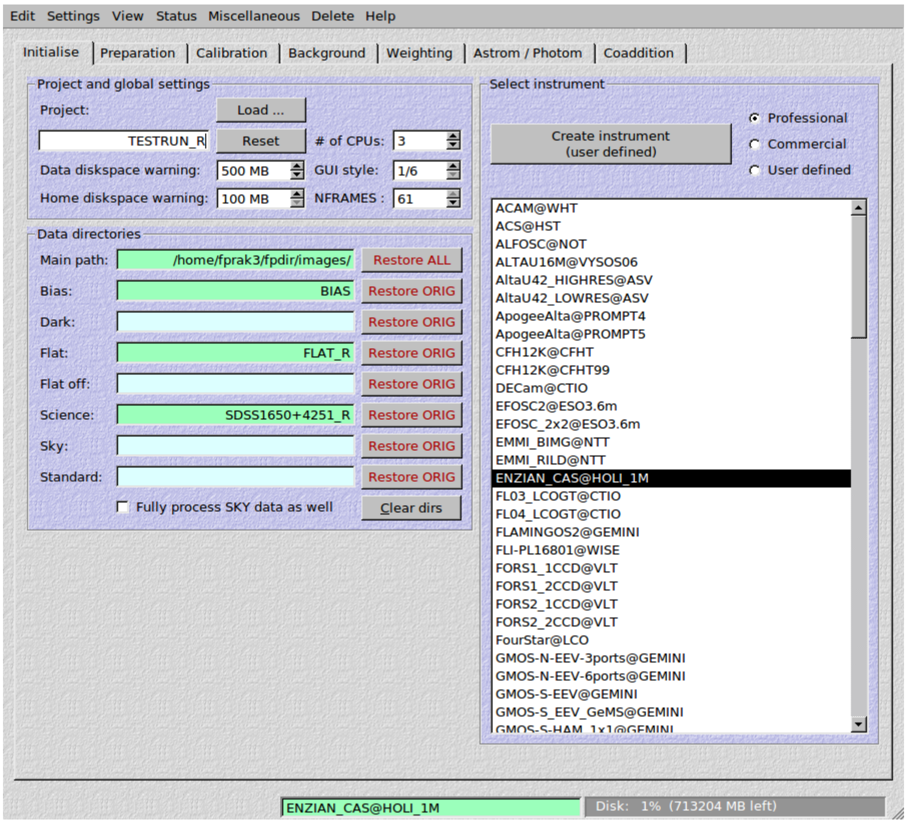

Figure 22: The *Initialise* PG

Data Directories The LineEdit fields collected in this GroupBox tell *THELI* where the data is that you want to process. You specify the main directory path (hereafter: *maindir*) that contains everything, followed by the names of the subdirectories that contain the BIASes, FLATs etc. If the folders exist they are highlighted in dark green. All fields do not accept blank characters as input, in addition the subdirectories do not accept a slash (/). You only need to specify those subdirectories which you actually need.

Restore ORIG This **deletes all data** in the corresponding directory, apart from the very raw data that has been moved into the ORIGINALS subdirectory. The ORIGINALS data is played back, thus restoring the original state before any processing has been launched. If no ORIGINALS subdirectory is present, nothing will be deleted.

Clear dirs This PushButton clears all LineEdits.

A.2 The PREPARATION processing group

A.2.1 Split FITS / correct header

The main job of this task is to split multi-extension FITS files into single chips, thus allowing for parallel processing. It also writes a new FITS header conformed with the *THELI* pipeline. If single-chip images are given, only the FITS header will be updated.

This reduction step will be applied automatically to all subdirectories that are specified in the *Initialise* PG.

How to redo the task Delete all split images in the corresponding directories, and move back the images from the `ORIGINALS` directory.

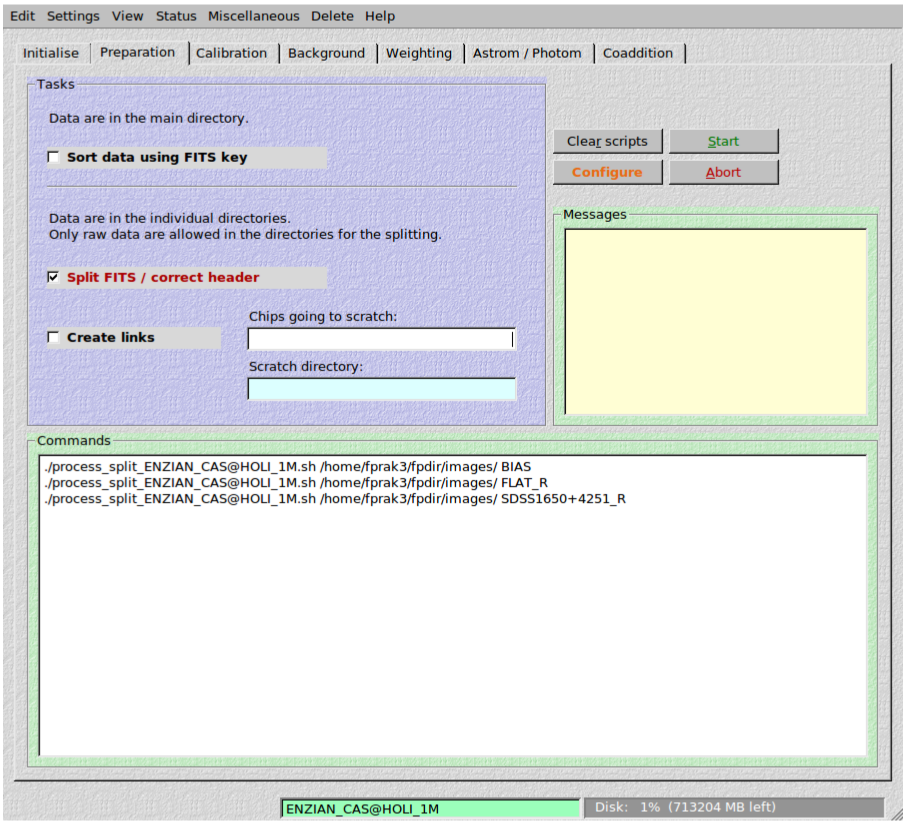

Figure 23: The *Preparation* PG.

A.3 The CALIBRATION processing group

A.3.1 Process biases / darks

Combines all BIASes in the BIAS subdirectory set in the *Initialise* PG. The more BIASes you have the better your master BIAS will be and the smaller the calibration noise that is introduced into your SCIENCE images. We recommend to use at least 10 BIASes. Overscans are corrected and trimmed.

A.3.2 Process flats (MANDATORY)

Combines all FLATs in the FLAT subdirectory. A FLAT correction is very useful even if the camera appears to be illuminated very homogeneously. This is because the FLAT does not only correct for vignetting effects, but also for different sensitivities on a pixel-to-pixel basis. The more FLATs you have, the better the master FLAT will be and the smaller the calibration noise you introduce into your SCIENCE images. We recommend at least 10 FLAT exposures. The FLAT exposures are debiased, overscan corrected and scaled to the highest mode in the stack before combination.

How to redo these tasks Just re-run it a second time.

A.3.3 Calibrate data

Images are overscan corrected, debiased and flat fielded based on the master bias- and master flat frames.

How to redo the task Click on the little arrow next to the `Calibrate data` task. Then activate the task again and re-run.

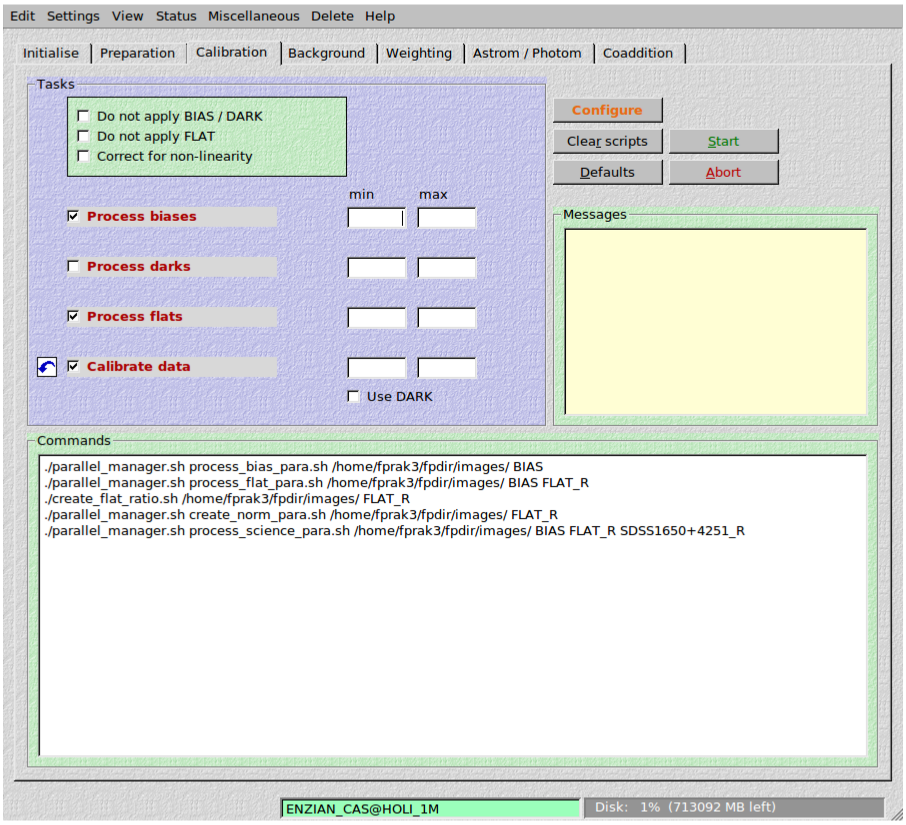

Figure 24: The *Calibration* PG.

A.4 The BACKGROUND processing group

A.4.1 Background model correction

The flat fielding is in general not sufficient to obtain a uniform sky background. Therefore *THELI* provides a set strategies to create and apply a background model:

- subtraction
- division (classical superflat)
- fringe subtraction
- fringe subtraction and division (illumination correction)

The configuration window consists of 5 groups, of which 1 and 4 are the most important ones.

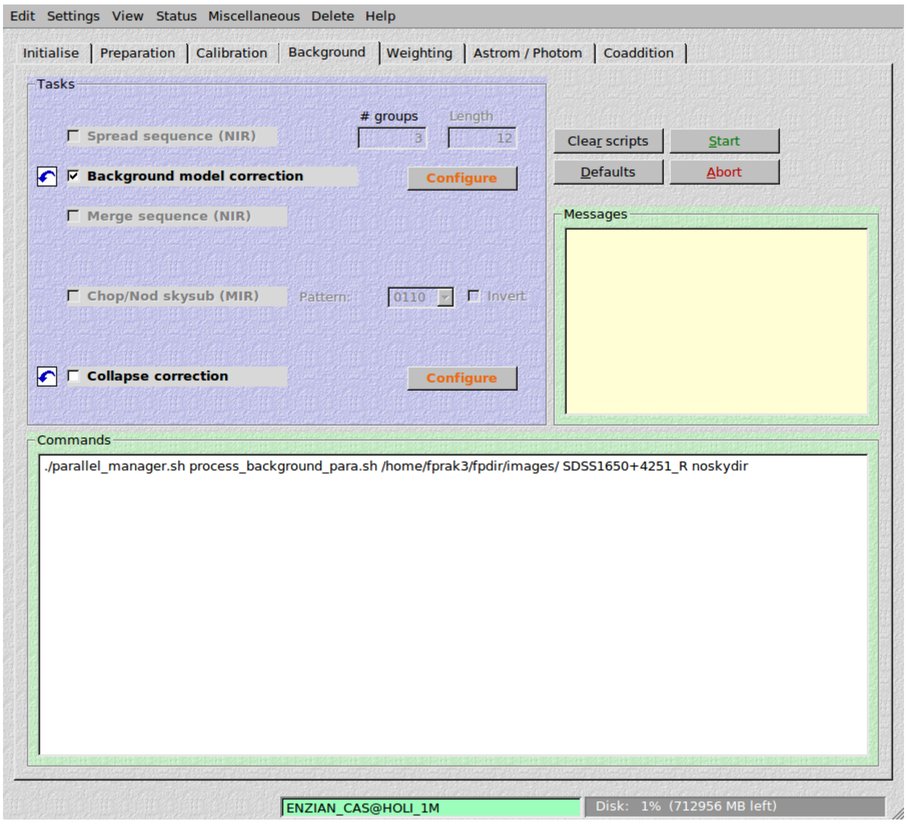

Figure 25: The *Background* PG.

Object masking Here you determine how objects in the images are detected, and which method is used for image combination. The left field takes the detection threshold (DT) per pixel, given in units of sigma of the sky background noise. The middle field takes the minimum number of connected pixels (DMIN) above DT which make up an object. The smaller both values are, the fainter the objects you mask. A good starting point for optical data is (1.0|5), depending on the flatness of the image and the detector size.

If your images exhibit strong fringing, then you can no longer use very low detection thresholds, since then the fringes themselves are detected as objects and thus removed from the background model. In this case one can no longer calculate a fringing model from the background model. We recommend to use a high S/N threshold if strong fringing is present. Try starting with (5.0|5) in this case. With near-IR detectors DT and DMIN often must be increased to 10 in order to not mask features in the very inhomogeneous sky background. If one or more of those three LineEdits is left empty, then the default values will be used without warning.

You can choose between a median and a mean combination for the background model. The median delivers a more stable result for a small number of stacked images, whereas the mean has lower noise when more images are stacked.

How to apply the background model At this step you decide how to apply the background model to the data. You may specify if and how to smooth the background model (and fringing model, if any). These model can be found in the `BACKGROUND` folder.

The right method to choose depends on the kind of background contamination you find in your data. If it is dominated by pure sky background, just use the model subtraction. If you are affected by other effects (e.g. scattered light from the surrounding or moon light), try to divide the data by your model. If you see a fringing pattern, take on of the defringing methods (depending on if you need additional illumination correction or not).

How to redo the task Just rerun. Previous results will be overwritten.

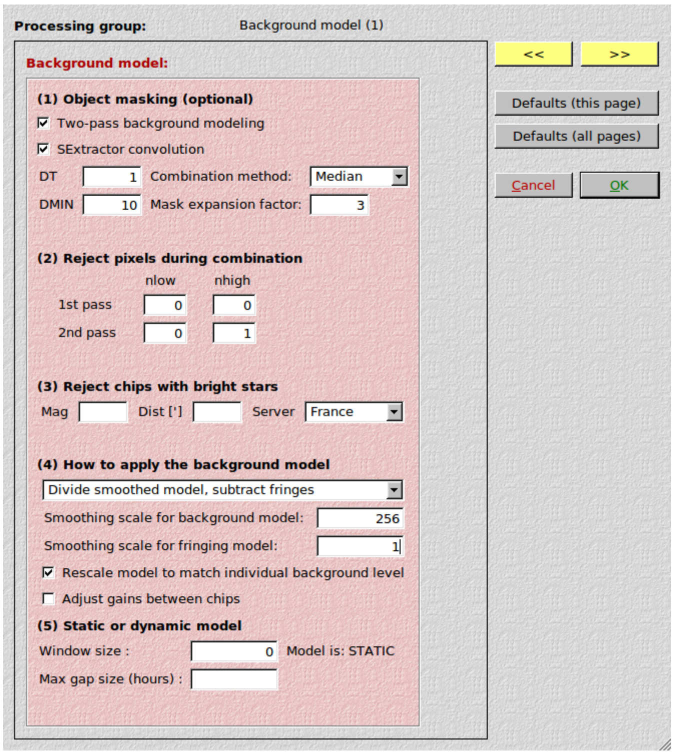

Figure 26: Configuration for the background model creation and application.

A.5 The WEIGHTING processing group

A.5.1 Create globalweights

The normalised FLAT is taken and has bad pixels replaced by zero values. Whether a pixel is bad is determined by one or more threshold pairs which refer to the normalised FLAT itself.

How to redo the task Just re-run.

A.5.2 Create WEIGHTS

All cosmics, hot pixels and other chip defects are detected on an image by image basis in this step.

How to redo the task Just re-run.

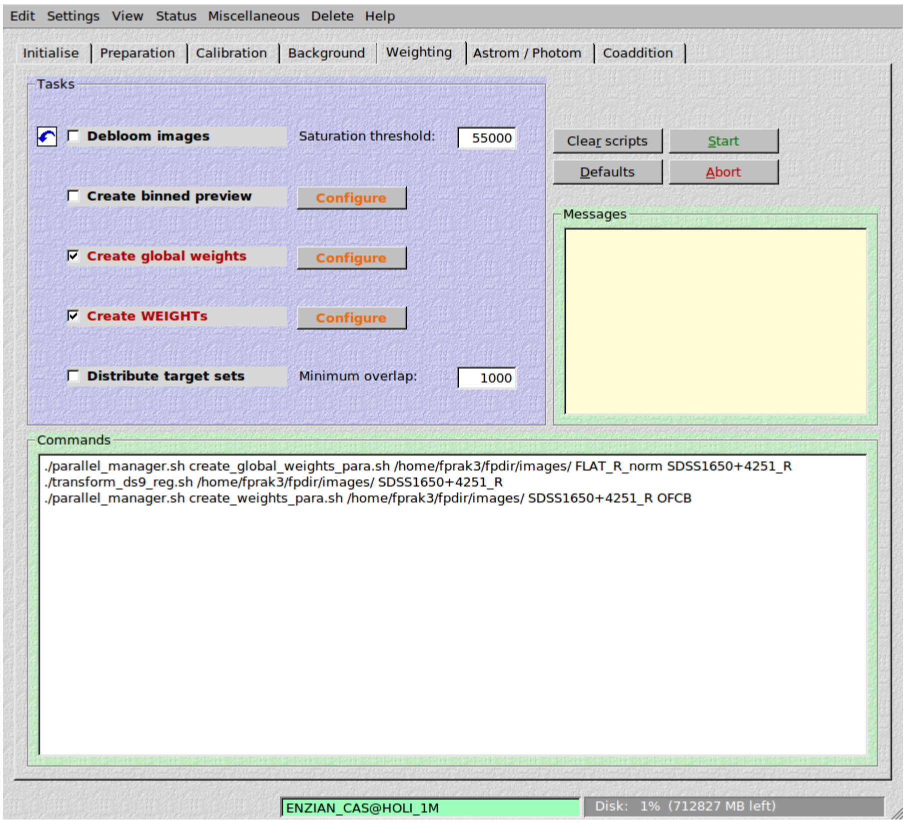

Figure 27: The *Weighting* PG.

A.6 The ASTROM / PHOTOM processing group

A.6.1 Astrometric reference catalogue

Retrieving a catalogue from the web This is the default state. Select the nearest web server and your catalogue of choice (SDSS if possible).

You can control the limiting magnitude of the objects retrieved (in the catalogue's magnitude system), and therefore their number density. The catalogue will be retrieved around the reference coordinates taken from the fits header (`from header`) within the radius specified. If the telescope pointing is biased significantly during an observation run, this should be entered manually.

The catalogue will be downloaded automatically once you start source detection. After you have inserted the correct values, click on [Get catalog].

How to redo the task Old results will be overwritten when a new reference catalogue is retrieved.

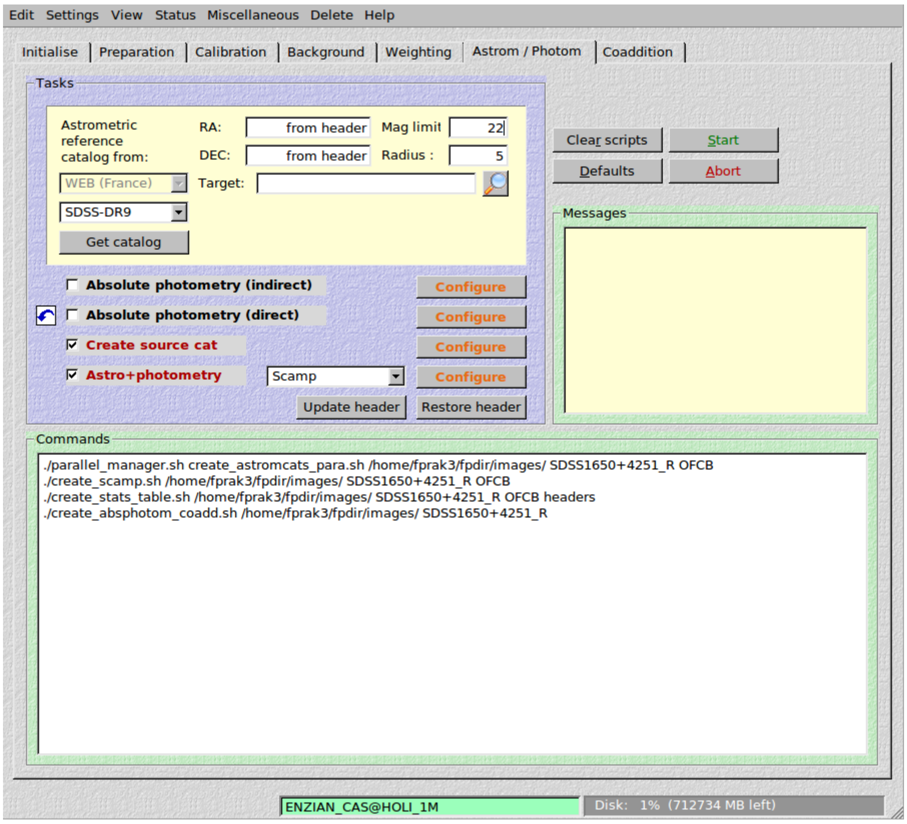

Figure 28: The *Astrom / Photom* PG.

A.6.2 Create source cat

Here we extract source catalogues from all images. The detection characteristics can be fine-tuned by means of the configuration dialogue (Fig. 29).

Parameter configuration *DETECT_THRESH* is the detection threshold (in sigma of the sky background noise) and *DETECT_MINAREA* is the minimum number of connected pixels above that threshold. The latter depends on the seeing and the pixel scale of your instrument. Three more parameters that usually do not need to be modified, are available, too (Fig. 29).

If you leave a field empty, then the default value will be used. No warning message will be printed.

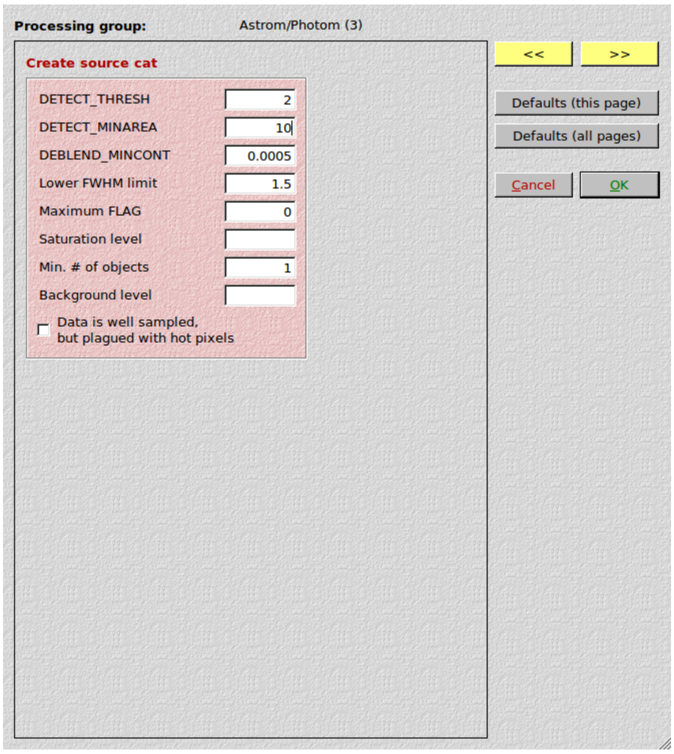

Figure 29: Configuration for the catalogue creation.

General tips for the catalogue creation The WEIGHT maps created previously are taken into account in the catalogue creation process, guaranteeing a clean catalogue that is largely free from spurious detections.

If the image quality is good you can have both DETECT_THRESH and DETECT_MINAREA as low as (5|5). If you have many hundreds or thousands of sources in an image, then choose a higher pair of thresholds, such as (10|10) or (40|15), or even (100|20). *THELI* has been used successfully with the latter parameters even for extremely crowded fields such as the Magellanic Clouds or exposures taken near the galactic centre.

The object catalogues created are saved in `SCIENCE/cat/`.

How to redo the task Just rerun. Old catalogues in `SCIENCE/cat` will be overwritten.

A.6.3 Astro+photometry

Here you can choose between three methods: *Scamp*, *Astrometrix* and *Shift only*. The latter determines only relative astrometric offsets and relative photometric zeropoints. It does not care for sky position angles or absolute sky coordinates. It can not handle mosaicing and only works for single-chip data. It requires a reasonable overlap of sources between exposures. This method is most useful if you work with images that show just one or a few objects. It serves as a fall-back solution if the other two solutions fail (for whatever reason).

Scamp (usually fast) and *Astrometrix* (usually slower) on the other hand are the entire opposite. They calculate the linear offsets between the reference pixel (CRPIX) and the reference coordinates (CRVAL), as well as two-dimensional distortion polynomials of higher order. There is no difference in running them on a single-chip or on a multi-chip camera. In the latter case, the solution is for the entire mosaic.

The results of the astrometry step, regardless of the method chosen, will be written to a `headers` subdirectory inside `SCIENCE`.

Configuration We recommend to use *Scamp* for this experiment. You can run it in its default configuration, with one small modification: the degree of the distortion polynomial should be set to 1 (instead of 3) as there are not many sources in our data that can be used for distortion correction. All other parameters can be left unchanged.

How to redo the task Just rerun. Old results in `SCIENCE/headers` will be overwritten.

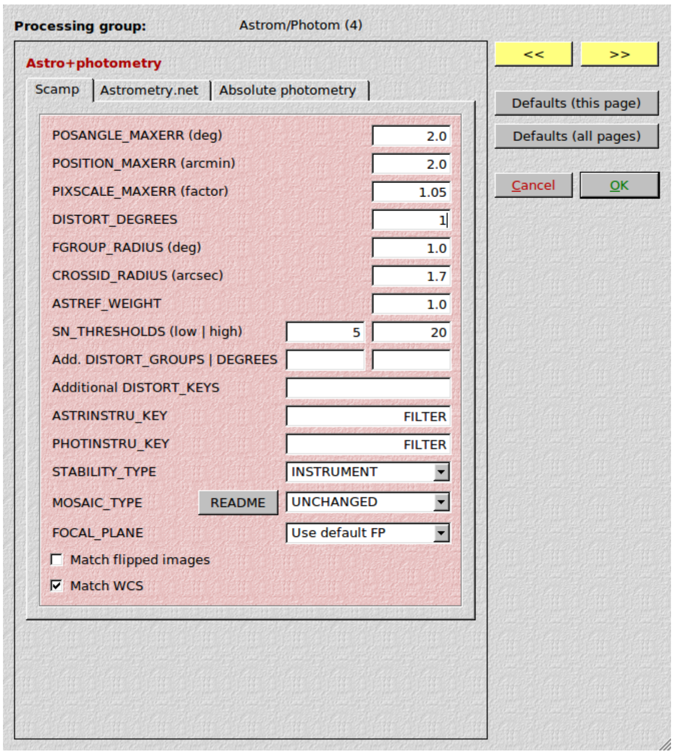

Figure 30: Configuration dialogue for Scamp

A.7 The COADDITION processing group

A.7.1 Sky subtraction

Configuration This step subtracts objects above certain user-provided thresholds from the image. From the remaining sky an estimate is determined. The *Configure* dialog (Fig. 32) presents you with the following options [*lab-course:* here we present only the option we are going to use]

Automatic sky modelling In a first pass, *SExtractor* is run to remove all objects from the image. The result is then smoothed and subtracted. To this end you must provide the usual detection threshold, minimum number of connected pixels, and the extent (pixels) of the smoothing kernel. This is the default method and useful for all exposures where the largest object is significantly smaller than the field of view of the detector.

How to redo the task Just rerun. Old results will be overwritten.

A.7.2 Coaddition

The coaddition goes in three steps. First, global information about the data set is obtained, and the reduction settings are made accordingly. Then, the SCIENCE images and their associated WEIGHTs are resampled. Lastly, the resampled images are combined.

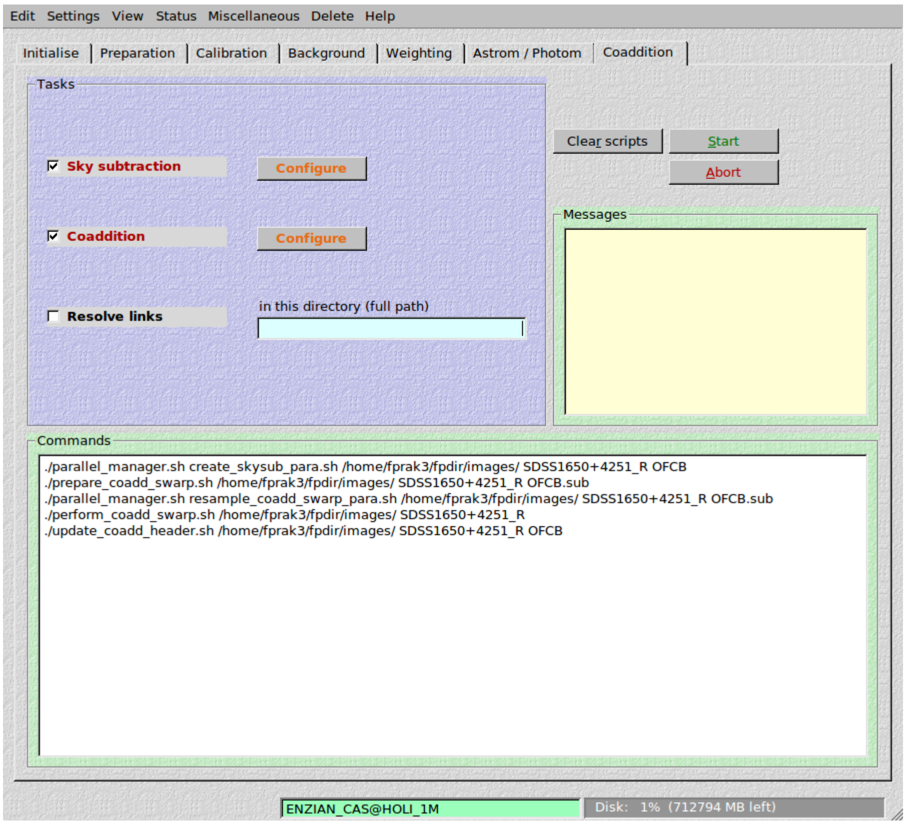

Figure 31: The *Coaddition* PG.

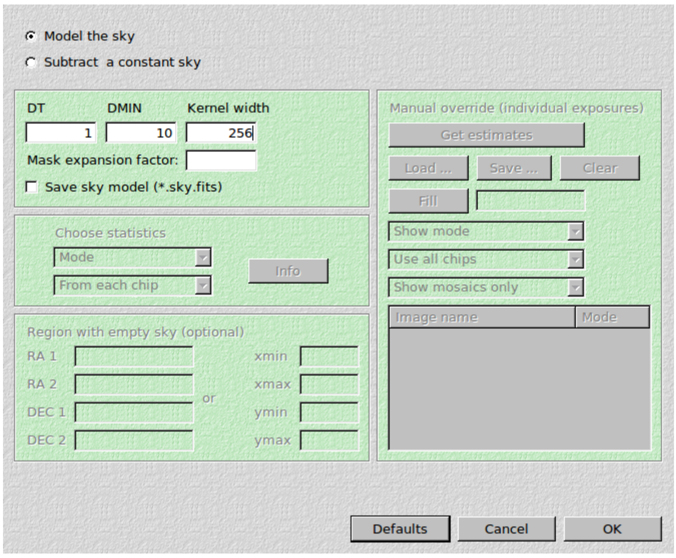

Figure 32: The configuration window for manual sky subtraction

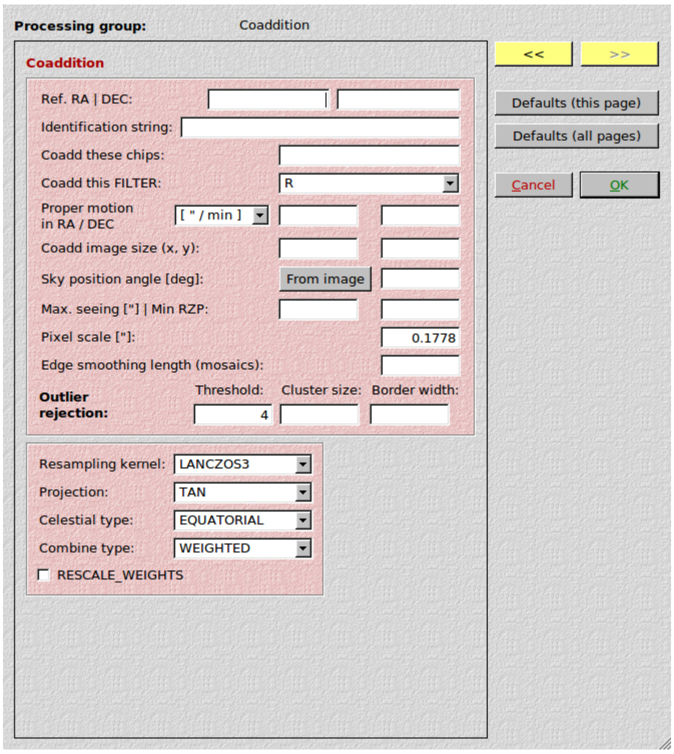

Figure 33: The configuration window for the coaddition

A.8 Image statistics

Calling *Image statistics* from the *Miscellaneous* pull-down menu, presents you with the dialog shown in Fig. 34.

Image directory Specify here the path to the images for which you want to obtain some statistics.

Name filter This is a string for filtering a subset of images out of all files in the specified directory. For example, it can simply be something like `A*OFCS.fits`. If left empty, all images (`*.fits`) in the directory will be considered.

How it works Clicking on the *Get statistics* PushButton will retrieve the statistics. This can take a while, and the GUI will not allow any other action during this time (may change in a future release). The table obtained will automatically be stored in the directory you specified. The name of this file will include the name filter, if such a filter was put. If the table is obtained a second time with an identical filter, the old file will be overwritten without warning. You can manually save the table to a different file name, or load a previously created table.

If the *Create source cat* and *Astrometry* processing steps were done as well, seeing and relative photometric zeropoints will be shown as well, respectively. That requires the presence of a `cat` and / or a `headers` directory in the specified path.

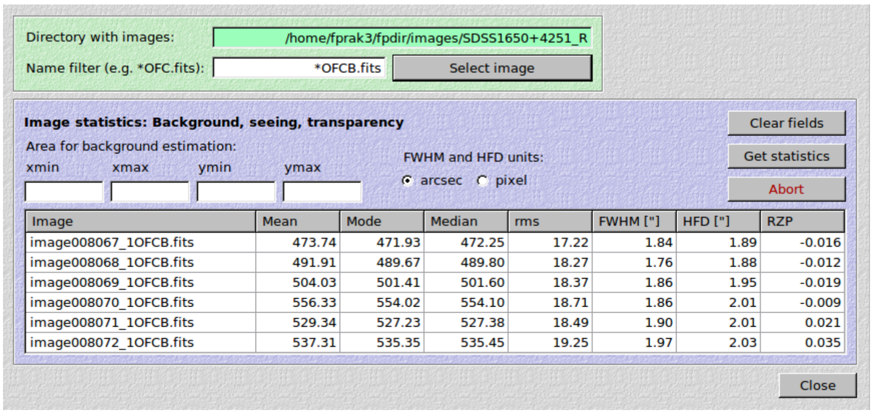

Figure 34: The dialogue for obtaining image statistics.

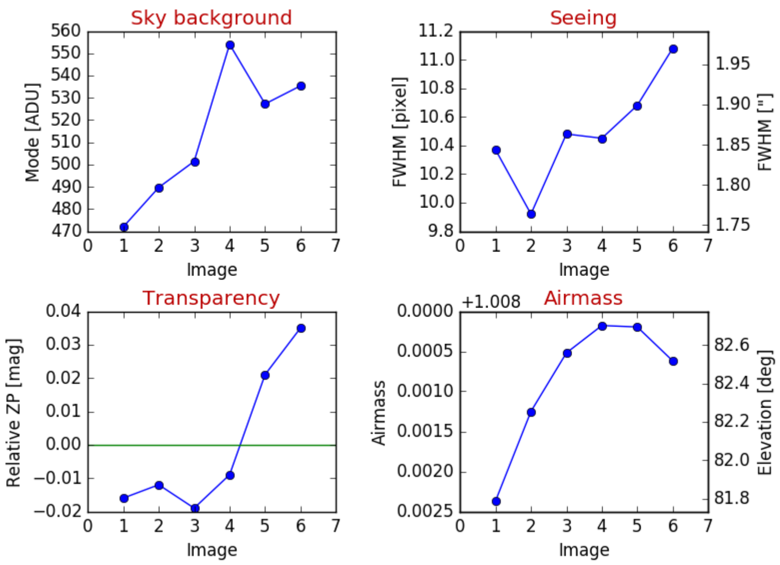

Figure 35: Example for statistics plot.