FP 30

CCD photometry in modern astronomy

Responsible MPIA staff contact:
Jörg-Uwe Pott (jpott@mpia.de)

MPI for Astronomy
Königstuhl 17
69117 Heidelberg

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Disclaimer for handling liquid nitrogen (LN2)

Students are advised to read Section 4 carefully before proceeding with the handling of liquid nitrogen. The emergency contact numbers provided there should be called immediately in case of any mishaps.

1 Introduction

The results of astronomical research contribute significantly to the picture we have of our universe and are of great relevance to many areas in physics as they allow the study of matter under extreme conditions not reproducible on Earth. However, the analysis of these distant celestial objects requires special, complex methods. Large telescopes equipped with highly sensitive detectors are needed to measure the electromagnetic waves originating from these distant objects. In the optical wavelength range (wavelengths from 300 to 700 nm) most commonly semiconductor detectors (so called $\underline{c}$harged $\underline{c}$oupled $\underline{d}$evices, CCDs) are used for this purpose. The high quantum efficiency of these detectors enables to detect up to 9 of 10 incident photons. A detailed understanding of the instrumentation, the elaborate measurement techniques, and the complex data analysis procedures is therefore essential for the astrophysical interpretation of the data.

The Max-Planck-Institut für Astronomie (MPIA) provides a powerful CCD camera for the use in one of the laboratory courses of the physics department of the Ruprechts-Karls-Universität Heidelberg. The camera is attached to the institute's 70 cm KING (Königstuhl Instrument for Night sky Gazing) telescope. In the first part of the experiment fundamental properties of the detector will be analysed, such as the behaviour of the dark current, linearity, dynamical range, and quantum efficiency. Important characterising quantities will be determined and compared to other detection devices.

For the second part of the experiment observations of a star cluster in two optical spectral bands will be analysed and interpreted. The programming language Python will be used to minimize artefacts in the raw data and to perform a photometrical calibration of the data to produce a colour magnitude diagram. This diagram, very similar to the "Hertzsprung-Russell" diagram, is one of the most important tools in stellar astronomy as it provides the means to determine basic parameters of a stellar population. For the star cluster used in this experiment the distance and the age have to be determined.

The experiment "CCD Photometry" demonstrates basic methods of astronomical observation and analysis. It gives an insight into the telescope and detector technology used in modern astronomy, with special focus on optics, semiconductor physics, cryogenics, and vacuum technology. The work with the telescope provides an authentic experience of astronomical observations and research. The analysis of the telescope data is based on Poisson statistics and shows the fundamental limits in detecting stochastic processes. An understanding of these processes and a critical evaluation of the data are of large importance to signal processing in all fields of science.

The experiment is scheduled to take place on four afternoons (or two full days) possibly including one evening for the night-time observations if weather permits. Since clear weather is needed for the observations the planning is flexible and students are encouraged to contact the supervisors in advance to agree on a time schedule. Unfortunately, the climatic conditions on the Königstuhl mountain are not very favourable to sensitive astronomical observations and only a few nights

of the year can be used for astronomical observations. The lab course can be completed without the night time experience.

An understanding of fundamental astronomical concepts (content of the lectures "Introduction into Astronomy and Astrophysics") is required for the participation in this experiment. Expertise in electronic data analysis and especially Linux (e.g. scripting) are of great advantage[1]. For the preparation of the experiment students are advised to read the following book chapters in addition to these instructions:

- Chapter 6 and 7 of Léna et al.: "Observational Astrophysics"(Léna et al., 1998)

- Chapter 5 and 6 in Unsöld, Baschek: "Der neue Kosmos" (Unsöld & Baschek, 2002).

The knowledge of this information (especially on the functionality of the CCD detectors, on the data reduction process and on the significance of the Hertzsprung-Russel-Diagram) will be verified in an oral discussion at the beginning of the experiment. Due to the high complexity of the experiment and the tight schedule, a poor preparation can prolong the execution of the experiment significantly and can lead to the rejection of the students. A careful preparation is therefore absolutely imperative!

The experiment takes place at the Max-Planck-Institut für Astronomie (MPIA), Königstuhl 17, in the eastern dome on the Elsässer-Labor situated to the south of the main institute. Directions can be found on the homepage of the institute: `http://www.mpia.de/anfahrt`.

The following instructions are divided into three main parts: First (section 2) a few fundamental principles are reviewed to ensure the basic comprehension for the experiment. The delineation in the context of this manuscript is, however, not to be considered exhaustive. The second part (section 3) contains a description of the instrumental setup. The third and main part (sections 4 and 5) comprises the problems to be approached as well as the instructions for the execution of the tasks. The instructions conclude with requirements and tips for the concluding report and an appendix with more in-depth information on the software used for this experiment.

Further web resources are

- MPIA house telescope:
  `https://svn.mpia.de/trac/gulli/king/wiki/WikiStart`

- FP30 at MPIA:
  `https://svn.mpia.de/trac/gulli/king/wiki/WikiStart/FP30Doc`

- Official info on the lab course and its grading system (only in german but web translate works):
  `http://www.physi.uni-heidelberg.de/Einrichtungen/FP/info/betreuerhinweise.php`

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[1]You will be faster and require less help from your supervisor. Successfully completing the experiment is also possible with little Linux or terminal/konsole experience.

2 Fundamental principles

This section contains all the basic background information which is needed to successfully master this experiment. Please read everything carefully to be well prepared.

2.1 CCD Detectors

In astronomy it is not possible to investigate celestial objects in controlled laboratory conditions. Almost all of our knowledge about the universe comes from measurements of the electromagnetic waves emitted from various forms of matter in space. The intensity of this radiation is very small due to the large distances in the universe[2]. Typical radiation fluxes have intensities of $10^{-29}\,\mathrm{W\,m^{-2}\,Hz^{-1}} = 1\,\mathrm{mJy}$[3] in radio astronomy or a few photons per m^2 and second in the optical wavelength range.

It goes without saying that the instrumentation used for the detection of radiation is of great significance for astronomy. In optical astronomy (wavelengths between $300\,\mathrm{nm}$ und $700\,\mathrm{nm}$) photographic plates played an important role since their invention in the early 19th century. This completely changed about 35 years ago with the introduction of semiconductor detectors. These devices make use of the photoelectric effect to convert the arrival of photons on the detectors into electrons that are then collected in condensators and read out. These so called charged coupled devices (CCD) are the standard tool for the observing astronomer and it is nearly impossible to imagine modern astronomy without them[4]. Almost all earthbound observatories and space telescopes are equipped with CCDs. They enable the detection of radiation ranging from the far infrared ($\lambda \lesssim 300\,\mu\mathrm{m}$), e.g. the Infrared Space Telescope ISO or the Spitzer Space Telescope) over the optical spectral range to the soft X-ray spectrum (with photon energies of $\mathrm{E} \sim 5\,\mathrm{keV}$, for example on the x-ray satellites Chandra and XMM-Newton). The widespread use of CCD detectors is due to several fundamental advantages compared to other light detection devices such as the photographic plate (photographic emulsion), photomultipiers, digicons, tv cameras, etc. The main advantages are:

- high sensitivity (high quantum efficiency of up to 90%)

- high dynamical range (16 bit)

- linearity over almost the entire dynamical range

- large spectral range (mid-infrared to ultraviolet for "optical" detectors)

- direct availability for further computer-aided data analysis

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[2]An exception is the radiation coming from our Sun.

[3]In radio astronomy, Jansky (Jy) is a commonly used unit to measure the monochromatic flux.

[4]CCD detectors are now also widely used in the industry, consumer electronics, and many other fields of science.
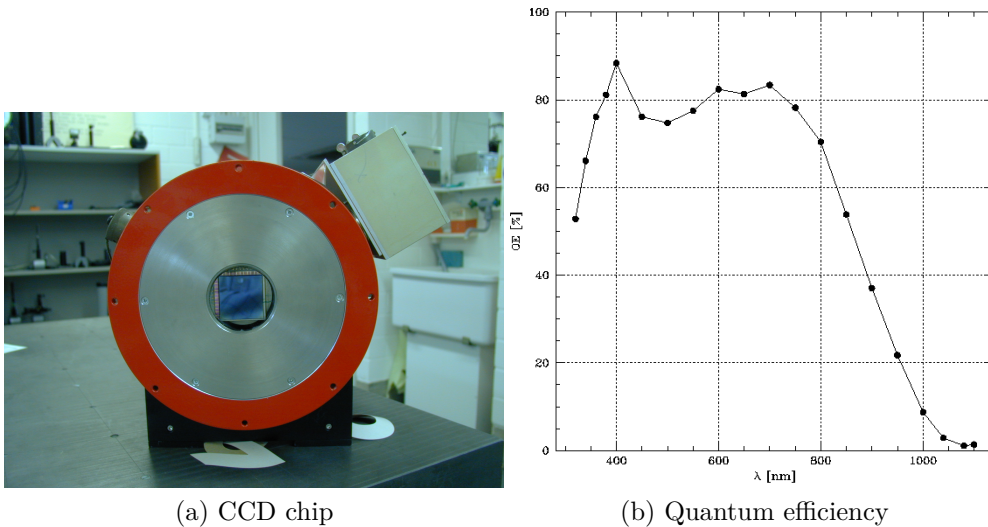
(a) CCD chip   (b) Quantum efficiency

Figure 1: CCD detector Loral/Lesser SITe 2048×2048 [6]: (a) Picture of the detector chip inside the vakuum tank. (b) Quantum efficiency curve of the detector as determined by the European Southern Observatory (ESO). In the wavelength range of 400 to 800 nm the efficiency is about 80%.

A CCD detector consists of a two-dimensional array of picture elements (pixels), which are light-sensitive metal oxide semiconductor (MOS) capacitors on a silicon substrate. Incident photons excite electrons into the conduction band, after which they get collected in the potential wells of the capacitors. After the integration of light is finished, an alternating voltage is used to shift the collected charge column by column to a readout column. The readout column is finally read out pixel-wise and the resulting signal gets amplified and converted to a digital signal by an analogue digital converter (ADU). Since the first CCD (that was designed as a data storage device by Boyle and Smith at Bell Laboratories in 1970(Boyle & Smith, 1970)), the size of the arrays has constantly grown to current sizes of up to a few thousand times a few thousand picture elements[5]. The chip used in this experiment, a 2048×2048 "Loral/Lesser SITe", exhibits impressive characteristics concerning read out noise and quantum efficiency. For the purpose of the experiment, the images are binned by a factor of two, such that the final images have a size of 1024×1024 pixel.. The detector is thus well suited for astronomical measurements, where we deal with low light intensities.

2.2 Observations

What does it mean when a detector is characterised as "linear"? Physically, this means that an increase in the measured detector signal is proportional to an in-

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[5]For example, Philips produces a 12 micron 9216×7186 pixel CCD.

[6]details on the detector: `https://svn.mpia.de/trac/gulli/king/wiki/WikiStart/CameraDoc`

crease in the incoming photon flux. In the linear regime, it is possible to directly compare the fluxes of different objects. The intensity of the flux coming from the celestial object is directly proportional to the intensity of the measured signal. When the detector reaches its saturation level, it is no longer possible to derive the exact number of photons that reached the detector originally. It is very important to always carefully consider the expected integration times for the observed objects. For weak sources integration times can be very long (up to several hundreds of seconds), whereas for bright stars integration times of only a few seconds can already lead to saturation.

Although the CCD detectors have a large dynamical range it is not always possible to simultaneously assure a good signal to noise ratio for weak objects of interest while not saturating very bright stars. The choice for the exposure time is thus dependent on the scientific goals of the specific observation. A photometric analysis and calibration is only possible if a large number of the objects of interest is not saturated.

For the planning of the observations two artefacts in the detector signal have to be further considered. A non-negligible amount of pixels on the detector is "dead", i.e. they produce no physically significant signal at all. Additionally, a certain amount of pixels is activated by cosmic rays such as electrons, γ-rays, muons, etc., with several pixels around the impact position of the cosmic ray being saturated or showing high numbers of counts. Whereas the amount and distribution of dead pixels stays the same over time, the number of "cosmics" is dependent on the integration time and their distribution is random. To eliminate both of these artefacts from the final image used for the scientific analysis, several exposures (at least 5) are taken of the science object with the telescope being slightly moved between each exposure. This method is called "dithering" or "jittering". Due to this offset in every frame and the random nature of the cosmic rays the artefacts will be located at different positions in each exposure. When we subsequently align the frames and determine the median of each pixel we can easily remove these extreme values. This method also has the advantage that it improves the signal to noise ratio without saturating brighter sources.

2.3 Data Reduction

The intensity over the detector area is not constant due to small scale variations as well as large scale gradients in the sensitivity. These variations are due to imperfections in the production process of the detector (e.g. "bad" pixels, variations in the doping of the semiconductor, residuals on the surface of the chip) as well as contaminations in the optical path (e.g. dust on lenses or filters). Since we want to compare objects imaged by different areas of the CCD chip, we have to correct for these intensity variations. We do this by taking exposures of a flat background ("flat field"), e.g. a patch of sky without any bright objects ("sky flat") or a homogeneously illuminated area in the dome ("dome flat"). By dividing our observations with the resulting normalised flat field, it is possible to eliminate the variations in the sensitivity from the science frame. For this correction, we assume

that these sensitivity variations are constant over time, at least for the duration of an observation.

To determine the photometric luminosities of the observed objects, the measurements obtained by the instrument have to be calibrated by comparison to stars of known intensity, so called "standard stars". Of course it is advantageous if standard stars are directly located within the frames containing the science target, but this cannot always be guaranteed, for example when observing galaxies in fields without stars. Then dedicated special observations of standard stars have to be carried out to determine the calibration factor for the instrumental magnitudes.

2.3.1 Bias

The voltage produced by electrons collected in each pixel (and later digitised to "counts") is artificially superposed during the read-out process in the detector electronics by a small constant positive value. This value, the so called "bias", must be subtracted again from all CCD-Images in the first step of data reduction. To determine its value, we make images of short integration time without exposing the CCD to light, i.e. the shutter remains closed during the exposures. The measured result is a sum of the bias and the electrons produced thermally, the "dark current". However, the contribution of the dark current is very small (in order of 0.01 e- s-1 pix-1) at operating temperatures of the CCD and thus practically negligible.

2.3.2 Flat-field

It is impossible to make the individual pixels of an array absolutely identical. The differing pixel properties cause intrinsic differences in the pixel sensitivities with the consequence that two adjacent pixel areas, illuminated with the same number of photons, will not produce the same number of counts. Moreover, the large-scale illumination pattern caused by the optics ("vignetting") as well as dust particles on the glass surfaces in the camera (filter, entrance window) will introduce instrument- and time-dependent artefacts on the raw images that must be removed. We can correct for this error by exposing the detector to a surface that shows homogeneous emission over its field of view, which allows us to directly record the variations of the pixel sensitivities. This is called flat-fielding and is a standard empirical method applied in optical astronomy.

2.4 Basic principles of photometry

With CCD detectors it is possible to measure the radiation flux reaching the Earth from astronomical objects. Astronomical units are often different from those commonly used in physics. For example, the intensity of a star is given in "magnitudes". We will now try to clarify the connection between this unit and the physical properties of stars.

The flux F from a star reaching an observer in a distance d is given by

$$F = \frac{L}{4\pi d^2} \tag{1}$$

where L is the luminosity of the star. The SI unit for the luminosity is J/s or W; in astronomy the cgs system is commonly used where the corresponding unit is erg/s (1 erg/s $= 10^{-7}\,\mathrm{J\,s^{-1}}$. However, it can be much more useful to work in units that take our Sun as a standard, since this allows us an easy comparison. This is why in the majority of cases the radiation flux is given in solar luminosities ($1 L_\odot = 3.846 \cdot 10^{26}\mathrm{W}$).

The flux of a star is connected to its surface temperature by the Stefan-Boltzmann law:

$$F = \sigma T_{\mathrm{eff}}^4 \tag{2}$$

where $\sigma = 5.67 \cdot 10^{-8}\,\mathrm{W\,m^{-2}\,K^{-4}}$ is the Stefan-Boltzmann constant. The effective temperature T_{eff} is the temperature a black body with the same surface as the star would need to emit the same radiation power. With $L = A \cdot F = Area \times Flux$ it is simple to derive a relation between luminosity and temperature:

$$L = 4\pi R^2 \sigma T_{\mathrm{eff}}^4 \tag{3}$$

where R is the radius of the star. Taking the example of our Sun with $R_\odot = 6.96 \cdot 10^8\,\mathrm{m}$ and $L_\odot$ given as above this leads to $T_{\odot,\mathrm{eff}} = 5770\,\mathrm{K}$.

A further parameter of prime interest is the mass of the star. The mass considerably influences the evolution of the star, with the fate of the star being almost entirely determined by its initial mass[7] Unfortunately, it is not trivial to deduce the mass of a single star from easily accessible observables.

The intensity of the incoming flux from a star, and thus the number of photons detected by a detector, decreases with the square of the distance (see eq. 1). As the distance is in most cases unknown it is not directly possible to specify the luminosity. The problem is aggravated by the circumstance that every instrument exhibits a different sensitivity to the light of the star. Thus it is necessary to introduce different magnitude scales:

The *instrumental magnitude* is a relative unit, allowing the comparison of different sources in one observation, or possibly in different observations with the same instrument and under similar conditions. The measured counts are converted to a logarithmic scale with arbitrary zero point:

$$m_{\mathrm{inst.}} = zero\ point - 2.5 \log counts \tag{4}$$

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[7]Another important factor for stellar evolution is the star's "metallicity", i.e., its amount of elements other than Hydrogen and Helium.

To accurately determine the total counts a point spread function (PSF) is fitted to the profile of the star or simply all the counts within an aperture centered on the star are summed up (this requires the subtraction of an estimated value for the background offset). With the help of standard stars with known magnitudes the instrumental magnitude can be converted to the apparent magnitude.

The *apparent magnitude* is the brightness of the star as perceived from Earth. The first to introduce an apparent magnitude scale was Hipparch (ca. 150 b.c.). Hipparch classified the stars by eye in brightness classes ranging from 1 for the brightest stars to 6 for barely visible stars. Later this classification was quantified assuming a logarithmic perception of the brightness by the human eye. Thus, a difference of 5 magnitudes now corresponds to difference in the brightness by a factor of 100. For the flux ratio of two stars this translates to

$$\frac{F_2}{F_1} = 100^{(m_1 - m_2)/5}. \tag{5}$$

The offset between instrumental magnitudes and apparent magnitudes can be determined with the help of standard stars. It is then possible to calibrate the observed magnitudes using this offset. The magnitudes obtained with different instruments and under different circumstances can now be compared to one another as all artefacts of the observation itself should be eliminated.

Finally, the *absolute magnitude* is defined as the apparent magnitude the stars would have if they were a distance of $10\,\text{pc}$ away [8]. This distance has been chosen arbitrarily to be able to physically compare stars at different distances, i.e. to have an absolute scale for the brightness of a star. The ratio of the flux reaching us (the apparent magnitude) to the hypothetical flux that would reach us if the star stood at a distance of $10\,\text{pc}$ can be calculated. Essentially only the apparent magnitude m and the absolute magnitude M feature in the resulting formula:

$$\frac{F\,(10\text{pc})}{F} = \left(\frac{d}{10\,\text{pc}}\right)^2 = 100^{(m-M)/5} \tag{7}$$

If we know the distance of the star it is possible to determine its absolute brightness from its measured apparent magnitude m. However, normally the distance to the star is not known. If in such a case the absolute brightness of the star can be estimated by other means (e.g., the spectral class), then the distance of the star simply is determined by

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[8]The distance $1\,\text{pc}$ (parsec) is the distance at which one astronomical unit (approximately the mean of the aphelion and the perihelion of the Earth's orbit around the Sun, $1\,\text{AU} = 1.496 \cdot 10^{11}\,\text{m}$), appears under an angle of one arcsecond (1"). For small parallactic angles p and $1\,\text{rad} = 2.063 \cdot 10^5$" the distance is given by

$$d = \frac{1\,\text{AU}}{\tan p} \simeq \frac{1}{p}\,\text{AU} = \frac{2.063 \cdot 10^5}{p''}\,\text{AU} = \frac{\text{AU}}{p''} = 3.262\text{ly} = 3.086 \cdot 10^{16}\text{m} \tag{6}$$

$$d = 10^{(m-M+5)/5} \; [\text{pc}]. \tag{8}$$

The expression $m - M = 5\log_{10} d \,[\text{pc}] - 5$ is called the distance modulus. For two stars at the same distance the following relation holds:

$$\frac{F_2}{F_1} = \frac{L_2}{L_1} = 100^{(M_1-M_2)/5}. \tag{9}$$

Thus for known luminosity and absolute magnitude of a reference star, it is possible to determine the luminosity of any other star at the same distance from its absolute magnitude.

For both the apparent and absolute magnitude scales, wavelength-dependent and bolometric magnitudes are distinguished. The *bolometric magnitude*, M_{bol}, corresponds to the flux integrated over all wavelengths, from the radio waves to x-rays. In practice it is difficult to determine the bolometric magnitude, as the photon detecting devices only cover certain wavelength ranges. Additionally the atmosphere of the Earth is opaque to large amounts of the spectrum. Hence direct measurements of the bolometric magnitude must be performed by space observatories to avoid absorption of a large part of the incoming photons. Nevertheless, we can use theoretical considerations to estimate the bolometric magnitude from earthbound observations. Alternatively, we can use a wavelength-dependent magnitude, M_λ, corresponding only to a part of the luminosity in a certain wavelength range, for example a luminosity in the "blue".

For a detailed measurement of the spectral intensity of the emitted light we need spectroscopic observations. However, we can obtain a rough spectral information by using well-defined filters, which are only translucent at a certain part of the spectrum. The most commonly used system of broad band filters is the so called Johnson system and modifications of it. The following table gives the four standard filters of the Johnson system in the visible spectral range with central wavelength λ_0 and width of the filter $\delta\lambda$.

Table 1: Spectral setup of the Johnson filter system (taken from Binney & Merrifield (1998))

Filter	λ_0 [nm]	$\delta\lambda$ [nm]	spectral range
U	365	66	ultraviolet
B	445	94	blue
V	551	88	visual (green)
R	658	138	red
I	806	149	infrared

For photometry normally several filters are measured so that it is possible to calculate a distance-independent colour index[9]

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[9]In the following equation m_V often is replaced by the filter name V

$$m_{\mathrm{B}} - m_V = M_{\mathrm{B}} - M_V = B - V. \tag{10}$$

The colour index gives principle information on the spectral properties of the object, e.g. that a star is bluer and thus hotter than another star. Not all filters exactly correspond to the Johnson system and there are various other filter systems in use. A correction for a different transmission behaviour compared to a perfect Johnson filter can be applied during the calibration with standard stars. In a first order approach the calibration of the instrumental magnitude assuming a slightly different transmission curve for the B filter is calculated according to

$$B = b + b_0 + c_{\mathrm{B}}(b - v) + a_{\mathrm{B}} \cdot airmass \tag{11}$$

where B is the apparent magnitude in the Johnson system and b the measured instrumental magnitude. The zero-point b_0, the colour correction c_{B} and the correction for the airmass a_B have to be determined by comparison with the standard stars.

2.5 The Hertzsprung-Russel-Diagramm

At the beginning of the 20th century, enough stars where photometrically measured to perform statistical analysis on them. It was noticed then, that basic parameters of the stars changed with spectral classification. O stars at the one end of the classification scheme were identified as young, massive and luminous objects, while M stars on the other end were low mass and faint objects. The relation between the spectral type, which is physically connected to the effective temperature T_{eff}, and the luminosity is referred to as the Hertzsprung-Russell diagram (HRD). Hertzsprung and Russell independently discovered this correlation at the beginning of the 20th century. The majority of the stars in the diagram is located along the so called main sequence, which is characterised by the hydrogen burning phase during the main phase of a star's life. Other areas in the diagram are populated by stars at different stages during their life. For a stellar population of the same age, as is the case in stellar clusters, the HRD shows the evolutionary history of the population.

In practice, however, it is not simple to compose a HRD, as absolute magnitudes or luminosities and the spectral types of the stars are not well known. This is especially the case for the spectral types: spectra for every object would have to be obtained for their classification. In observational astronomy the colour magnitude diagram (CMD) is used instead. In this diagram the apparent magnitude V is plotted over the colour index (e.g. $B - V$), which is connected to T_{eff}. Both quantities are directly deducible from observations with two filters. The colour magnitude diagram is equally comprehensive as the Hertzsprung-Russell diagram and important conclusions about stellar evolution can be derived from it. In the following sections important methods for the calibration and analysis of such a diagram will be discussed.

2.5.1 Shift of the main sequence

Assuming that all stars in a star cluster are approximately at the same distance and have the same chemical composition as well as similar ages, the distance to the cluster can be determined. For this purpose the measured, apparent magnitude is plotted versus the colour, e.g. V versus $B - V$. The position of the main sequence is then compared to the position of a cluster with known distance, assuming that the absolute magnitude of the main sequence is the same for star clusters. Small effects due to different metallicities (see section 2.6) are ignored here. The assumption is justified for old stellar populations as only low mass stars with uniform development are still located on the main sequence. The main sequence of the observed cluster is shifted in direction of the magnitude axis and aligned with the main sequence of the cluster used for comparison. The amount by which the diagram has to be shifted corresponds directly to the distance modulus of the cluster, neglecting attenuating effects by interstellar dust. The distance of the cluster can now be determined by taking into account the distance of the cluster used for comparison. Alternatively to the comparison with a cluster of known distance, the measured main sequence can also be compared to predictions of its position by theoretical models of stellar evolution. Today, detailed models of the evolution of stellar populations exist. Evolutionary tracks can be calculated, showing the displacement of stars of defined mass in the plane of the colour magnitude diagram during their lifetime. Theoretical curves representing the position of the stars of a stellar population of the same age can be calculated from these curves. The latter curves are called "isochrones", for example the isochrones denominated Y^2(Yonsei-Yale) Isochrones (Yi et al., 2003).

Assuming that all stars within the cluster were formed at about the same time, one can also determine the age of the cluster from the colour magnitude diagram: The most massive and luminous stars have a short lifetime and thus only stay on the main sequence for a short period of time (several million years). The reason for this is the mass-dependent luminosity of stars and therefore the mass-dependent time a star spends on the main sequence:

$$L \propto M^a \tag{12}$$

$$\tau_{nuc} \propto \frac{M}{L} = M^{(}a - 1). \tag{13}$$

The exponent a is mass dependent as well (e.g. 3.5 for $2M_\odot < M < 20M_\odot$). This means, that first the massive stars at the top left end of the colour magnitude diagram drift into the red giant branch, after all their hydrogen has been burned in their cores. With increasing age of the stellar population lower mass stars also start moving into the red giant branch. This shifts the so called *turn-off point* in the diagram to lower masses and luminosities. From the position of the turn-off point it is hence possible to determine the age of the cluster. For old clusters this method allows a good determination of the age, for young clusters caution is advised, as often no unambiguous turn-off point can be defined. The position of

the turn-off point is also dependent on the metallicity of the cluster, which has to be considered when doing the comparisons.

2.6 Stellar clusters

Stellar clusters can be roughly divided into two groups: Open clusters and globular clusters (GC). The main difference between the groups are the age and the mass of the clusters. Globular clusters are much older and more massive than open clusters. GCs are among the oldest objects found in the universe. Occasionally, they even seem to be older than the universe itself! However, the ages of GCs have large errors and comparisons between different measurements can show large discrepancies. The stars in globular clusters of our Galaxy are very metal poor, they consist almost only of hydrogen and helium (which is approximately the composition of the matter found in the universe after the big bang) with only minor contaminations by higher elements, i.e. metals. The deficiency in metals is an indication that these objects must have formed in the very early phases of the universe when the interstellar matter was not yet enriched with heavier elements. Stars with such a metal deficiency are classified as population II stars.

Globular clusters are spherically symmetric objects, dispersed all over the Milky Way halo with a slight concentration towards the centre. They constitute a major portion of the halo of our Galaxy up to radii of 55 kpc from the galactic centre. In total about 150 globular clusters are known to orbit our Milky Way. It is assumed that the majority of GCs formed at the beginning of the evolution of our Galaxy. Thus their distribution, velocity and composition give information about the evolution of the Milky Way.

Open clusters on the other hand have a wider range of ages between a few Myr up to a few Gyr. The metallicity of open clusters varies as well depending mostly on the age and the location within the host galaxy of the open cluster. Therefore open clusters are good tracers for the evolution of galaxies, as many properties (age, metallicity, etc.) of their stars are more easily determined than for isolated field stars. Open clusters are also very good tracers of star formation, since most stars are formed in clusters.

Most of the open clusters of the Milky Way are located within the disc of our Galaxy. There are about 1000 open clusters known within the Milky Way. However, their total number is likely much higher, since their lower surface brightness means that we are not able to detect them to such a great distance as the more compact GCs and many open clusters are hidden behind dust clouds and the Galactic center.

As all stars of a stellar cluster are located at about the same distance and have the same metallicity and age[10] they are especially well-suited for the analysis with a Hertzsprung-Russell diagram. The colour magnitude diagram of the star cluster Messier 92 (M92) has a characteristic form as can be seen in figure 2. Stars

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[10]There are however globular clusters for which two red giant branches are found indicating a second generation of stars ((Milone et al., 2008)).
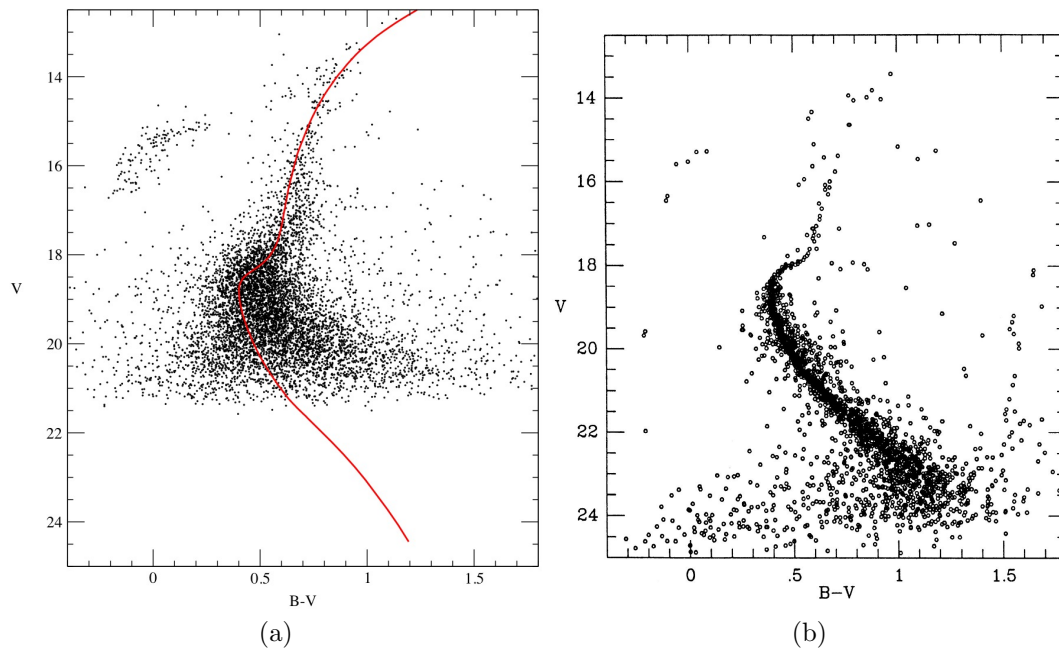
Figure 2: Colour magnitude diagram of the bright globular cluster M92: (a) Colour magnitude diagram from data observed with the 70 cm telescope of the MPIA on 07/06/2004 by K. Tristram. A Y^2 isochrone for a metallicity of $[Fe/H] = -1.9$, an age of 16 Gyr and a distance modulus $\Delta = 14.8$ mag is plotted into the diagram. (b) Diagram published in 1988 by (Stetson & Harris, 1988). The observations of Stetson and Harris go four magnitudes deeper than the observations with the 70 cm telescope.

in different evolutionary phases such as main sequence stars (hydrogen burning phase) and giants (helium burning phase) can be located easily.

3 Experimental prerequisites

The setup of the experiment is given by the sketch in figure 3. Light gathered by the primary mirror is imaged on the CCD-detector, which is cooled with liquid nitrogen within an evacuated cryostat. To read out the detector, a pulse generator, amplifier, and analogue-digital-converter is needed, which are put into a separate box, mounted on the telescope itself. A graphical user interface allows easy use of the camera and access to the data and can be started from the available workstation. As with every astronomical instrument in use, continuous changes and improvements are performed to increase the ease of use and the availability of the data.
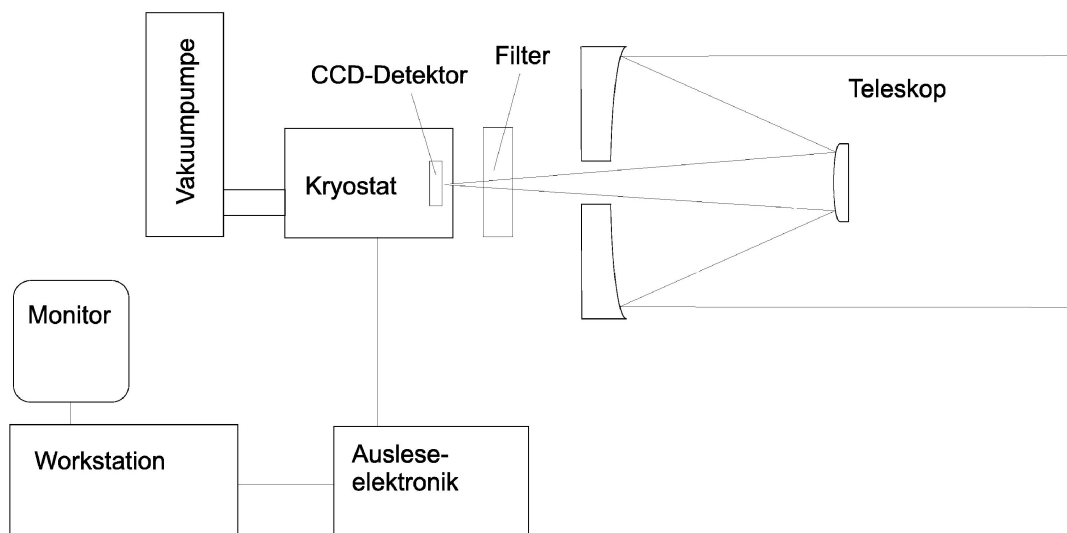
Figure 3: Diagram of the instrumental setup as of January 2006.

3.1 Hardware

The experiment is performed with the 70 cm telescope at the Max-Planck-Institut für Astronomie (MPIA) on the Königstuhl mountain. The telescope is depicted in figure 4, the general setup can be seen in figure 3.

The telescope structure was built at the institute in order to have a facility to test cameras and other astronomical devices built at the MPIA. Therefore it has a parallactic mounting. This enables to switch between two different foci: Cassegrain mode - which is used in our experiment - and Coudé focus, which decouples the beam through the polar axis and enables to attach large instruments on the ground floor, where the beam ends. In the past, another reason for this kind of mounting was the easy compensation of Earth's rotation by just rotating around the polar

Figure 4: The 0.7 m KING (**K**önigstuhl **I**nstrument for **N**ight-sky **G**azing) tele-scope at the MPIA. The major components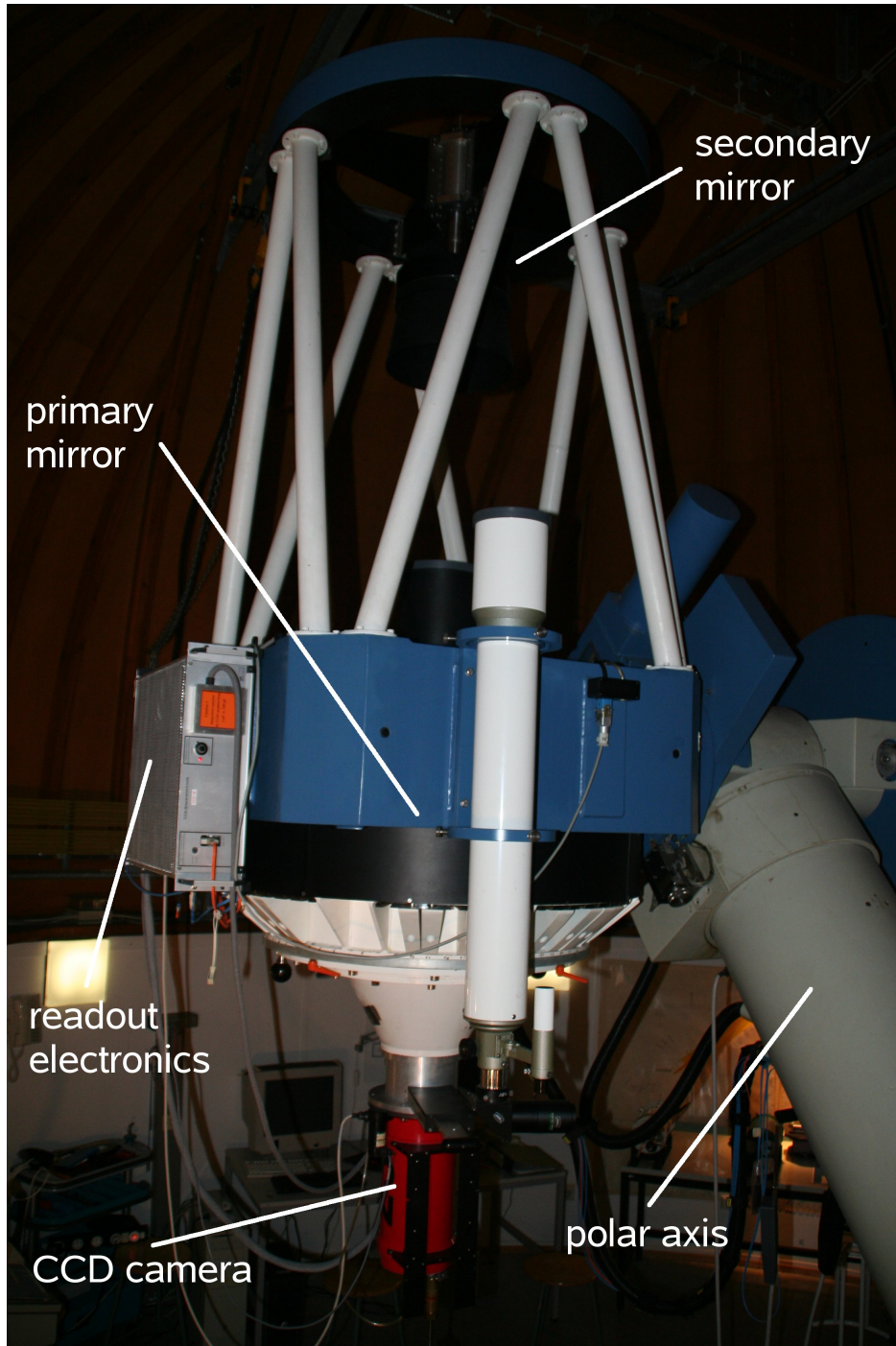 of the telescope are indicated in the picture. The CCD-camera is attached below the main mirror in the Cassegrain focus. It is situated in the red cylindrical dewar, which is cooled with liquid nitrogen. Also visible is the parallactic mounting of the telescope.

axis. After reflection by the main mirror with a diameter of $70\,\text{cm}$ the light is reflected another time by the secondary mirror and reaches the camera (for the Cassegrain case) through a hole in the primary mirror. The position of the secondary mirror can be changed to adjust the focus of the telescope (by pressing the respective button on the control panel). The camera used for this experiment is located in the Cassegrain focus below the main mirror in a cryostat cooled by liquid nitrogen (see figure 4). Before cooling the camera with liquid nitrogen the dewar has to be evacuated to a pressure of about $p = 10^{-6}\,\text{mbar}$ to suppress heat conduction and avoid water vapour to freeze out on the camera and detector. This is achieved by a dual stage vacuum system with a rotary vane pump for the preliminary vacuum and a turbo molecular pump for the high-vacuum. Just above the camera a slider containing up to five filters is located. Usually they are ordered by wavelength, but before using them it has to be checked which filters are installed in the slider. With the help of the camera optics the light gathered by the telescope is focussed on the CCD detector.

The detector is a *Loral/Lesser n2k2eb BI* thinned, back illuminated CCD with 2048×2048 picture elements and a pixel size of 15 microns. CCD-cameras in general are characterised by a very high quantum efficiency of up to 90%, which makes them ideal for astronomical observations. This explains their widespread use in all observatories over the world. The pixel scale of our camera is 0.55 "/pixel corresponding to an optimal adaption for the seeing conditions on the Königstuhl mountain. Atmospheric turbulences smear out the ideally point-like stellar sources to a profile similar to a Gaussian with a full width half maximum of 2" to 3". The pixels saturate at a level of 60000 electrons at a gain value of 5.

The box on one side of the telescope contains the readout electronics. It contains a pulse generator for the read-out procedure, an amplifier, analogue digital converter as well as the controller for the heater of the CCD-camera (in order to keep it at the ideal temperature of $115\,\text{K}$. The read-out electronics are controlled by a UNIX workstation called "*KING*".

So far it is not possible to control the telescope position from the workstation, although this feature is planned to be implemented in the near future. The controls are located in a special console in the telescope dome. During the experiment you will only need the controls for declination, right ascension, guiding, mirror protection, and the focus. The observed data is stored via network-attached storage (NAS) on a Linux computer. This computer can also be used for data visualisation and analysis during the experiment and the final data reduction and analysis. The Dome can be handled via a remote control by pressing at the same time the red button and another functional button for rotating left/right or opening/closing the slit.

3.2 Software

This section gives an overview of the software used throughout the experiment and data analysis. Details on the needed commands are given in the respective sections, where they are used. The CCD camera is controlled via command line

commands on the king computer (see section 8 for details). The data reduction, analysis and part of the visualisation is done with the help of Python and serveral packages for the use of astronomical data reduction and analysis.

Images can be displayed with the viewer SAOIMAGE DS9 or Python's matplotlib package. For the final data analysis, any plotting/fitting tool (e.g. Python/Matplotlib, Origin, GNU-plot, XMGRACE, etc., or even Microsoft Excel) can be used.

3.2.1 Python

Because of the huge variety of freely available packages and its easy extensibility, Python is today the most used programming language in astronomy. In this experiment, Python is used in several ways. There are console scripts available for the control of the CCD camera at the telescope, which can be executed from the terminal. For the analysis of the globular cluster, `Jupyter-notebooks` (JPN) are used. They provide an easy to follow workflow and individually executable script blocks. We recommend to also use JPN for the characterization of the CCD, however, you can also use IPython, Spyder or anything else you are familiar with.

Jupyter-notebook

You can start a Jupyter-notebook from the terminal with:

cl> jupyter−notebook

Which will open the home page of JPN in a browser. From there you can open existing notebooks by browsing the file system or create new ones by clicking on "NEW" in the upper right corner.

In a JPN you have cells where you can insert your code and run the cell by clicking the button "RUN" on top. If you want to delete the current output, you can either use the tab "CELL->CURRENT OUTPUTS / ALL OUTPUT -> CLEAR" or you can restart the kernel and clear all output under "KERNEL -> RESTART & CLEAR OUTPUT". More cells can be added from the tab "INSERT -> INSERT CELL ABOVE / INSERT CELL BELOW".

Usefull Python packages and commands

Basic Python

```python
print("Text to print", variable)  # print statement for printing text
                                  or variables, multiple arguments
                                  must be separated by a comma.
x = []                            # creating an empty list
x.append[1]                # appends 1 to the list x
x = [1, 2, 3, 4]              # creating a 1-dimensional list
x = [[1, 2], [3, 4]]       # creating a 2-dimensional list
```

Looping over a list is very simple in Python. For example:

```python
for value in x:
    print(value)
```

```
# Note: there is also "enumerate", which yields the index of each
                                  element and the element.
for idx, value in enumerate(x):
    print(idx, x)
```

Numpy

`Numpy` (Oliphant, 2006) is a fundamental package for data analysis. It provides the numpy-array structure which can be used for handling multidimensional data in a comfortable way. Many convenience functions can be used for mathematics and basic statistics on the data.

Usefull commands:

```
import numpy as np
x = np.array([1, 2, 3, 4])    # creates a numpy array
np.savetxt("filename.dat", x) # saves the array x under filename "
                                  filename.dat"
y = np.loadtxt("filename.dat") # reads in the data stored in "
                                  filename.dat"
```

You can perform mathematical operations on arrays, like adding (+), subtracting (-), multiplying (*) or dividing (/) arrays. Exponents are obtained by attaching "**". Additional useful operations are:

```
z = np.sqrt(x) # computes the element wise square root of the array x
                  .
z = np.exp(x) # computes the element wise exponential function of the
                  array x.
z = np.mean(x, axis=None) # yields the mean of the flattened array x.
                            If you want to specify to compute
                            it only on a specific axis, this
                            is possible with the "axis"-
                            keyword.
# Other useful functions with the same scheme are:
np.median(x, axis=None)
np.min(x, axis=None)
np.max(x,axis=None)
np.std(x, axis=None)  #computes the standard deviation of x
```

Basic slicing and indexing of numpy arrays

If you want to select a slice out of a numpy array you can use indices in brackets. See the examples below for different use cases.

```
x = np.array([1, 2, 3, 4, 5, 6, 7, 8, 9])
y = x[np.array([0, 3, 5])] # returns the elements with index 0, 3 and
                              5 from x
z = x[3:5] # returns the elements with index interval [3,5). Note
                              that the upper index is NOT in the
                              interval.
z = x[3:] # returns all values with index >= 3
z = x[:3] # return all values with index < 3
```

This is also possible for multidimensional data. Thereby the dimensions are separated by comma and indexing can be used as shown above. Additionally ":" can be used to get all values from one dimension.

```python
a = np.array([[1, 2, 3], [4, 5, 6]])
b = a[:,2] # returns [3, 6]
c = a[0, :] # returns [1, 2, 3]
d = a[0, :2] # returns [1, 2]
```

Astropy

`Astropy` (Astropy Collaboration et al., 2013; Price-Whelan et al., 2018) is the most important and versatile package for astronomy in Python. The complete documentation can be found at `https://docs.astropy.org/en/stable/index.html`. For the purpose of this experiment we are mainly interested in its capability of handling FITS files (`https://docs.astropy.org/en/stable/io/fits/index.html`).

Handling FITS files:

```python
from astropy.io import fits
hdul = fits.open("fits_image_filename") # opens a FITS file
print(hdul[0].header["HEADERKEYWORD"]) # prints the value stored
                                       # under HEADERKEYWORD in the header
                                       # of the FITS file (can also be
                                       # assigned to a variable instead)
image = fits_file[0].data # gets the image data from the file,
                          # returns a numpy-array.
# Alternatively you can directly load the data without opening the
                          # file before with:
image = fits.getdata("fits_image_filename")
fits.writeto("filename.fits", image) # saves the image variable as a
                                      # fits file.
hdul.close() # closes the file
```

Matplotlib

If you want to visualize data in Python, `matplotlib` (Hunter, 2007) is usually your first choice. It can be used for plotting data or showing images. The full documentation, with lots of examples can be found at `https://matplotlib.org/index.html`. For the analysis of the data taken in this experiment, the example codes below should be sufficient.

Basic plotting with matplotlib:

```python
import matplotlib.pyplot as plt
plt.figure() # opens a new figure
plt.title("Titel of the figure") # optional title for the figure.
plt.xlabel("Xlabel") # You can use LaTeX inline equations ("$ $") in
                     # the string for Greek letters and
                     # superscript / subscript letters.
plt.ylabel("YLabel")
plt.ylim(0, 5000) # optional, can be used for setting the plot limits
                  # . Also valid with "xlim".
```

```
plt.yscale("log") # optional, sets the scale on the y-axis to be
                                  logarithmic. Same for xscale.
plt.scatter(x, y, color="black") # scatter plot of x- and y-data.
                                  Color argument can be used to
                                  adjust the colour of the points.
plt.plot(x, y) # line plot of x- and y-data. Colour argument can be
                                  used to adjust the color of the
                                  line.
plt.errorbar(x, y, yerr=yerr) # line plot with error bars on the y-
                                  axis data. Argument "fmt="o"" can
                                  be added to get a scatter plot
                                  instead. Also takes the colour
                                  argument.
plt.show() # shows the plot, after all plot commands are finished.
```

Basic image visualization with `matplotlib` (presuming an "image" variable that stores the 2 dimensional image as a matrix):

```
import matplotlib.pyplot as plt
from astropy.visualization import simple_norm
norm = simple_norm(image, "log", percent=100) # The percent keyword
                                  sets the percentage of the image
                                  values used to determine the pixel
                                  values of the minimum and maximum
                                  cut levels. Can be tuned to get
                                  more details
fig, ax = plt.subplots()
ax.set_title("Title of the Figure")
im = ax.imshow(image, norm=norm, origin="lower", cmap="viridis") # "
                                  cmap" argument sets the   colormap
                                  which is used for the plot
fig.colorbar(im, ax=ax)
fig.show()
```

Plotting histograms with `matplotlib`:

```
plt.figure()
plt.hist(data.flatten(), bins = 100) # "data" should be the variable
                                  name for which you want to plot
                                  the histogram (must be a numpy
                                  array). The "bins" parameter takes
                                  the number of bins for the
                                  histogram.
plt.show()
```

Fitting with Python

There are various ways to fit a function in Python. A simple-to-use one is "curve_fit" from the `scipy` (Virtanen et al., 2019) package. It takes a fit function and your x- and y-data and returns the optimised parameters from the fit function and the corresponding covariance matrix.

```
def fit_func(x, param1, param2):
    return param1*x+param2
```

```
popt, pcov = curve_fit(fit_func, xdata, ydata)
print(popt, np.sqrt(np.diag(pcov))) # prints the results and the
                                    corresponding errors
```

You can also use the results for plotting the fitted curve by adding:

```
plt.plot(xdata, fit_func(xdata, *popt))
```

to your plot.

Creating file lists

For processing many files at the same time, you need to create a list of the file-paths of the files you want to use. In Python you can use the `glob` function for this. It takes a string and returns all the file paths that match the string. Thereby you can use wildcards (*) to retrieve multiple paths at the same time. For example:

```
from glob import glob
all_dark_fits = glob("./dark*.fits")
```

returns a list of file paths for all fits files that start with "dark" in the folder from which the python script is run. **Caution! The file list from `glob` is not necessarily a sorted list. If you want a sorted list add:**

```
all_dark_fits = sort(all_dark_fits)
```

You can loop over such a list with:

```
for path in all_dark_fits:
    print(path)
```

3.2.2 SAO-Image DS9

The best way to get an idea of how SAO-Image DS9 works is to open an image and to play around with it. Most of the menu structure and buttons should be self-explanatory. We would like to give just a few hints, which might be useful for our image analysis:

1. For most of our data-sets, a reasonable representation can be obtained by pressing the buttons *Scale → zscale*

2. Push the right mouse button while moving over the image. This will change the color-scale as well.

3. The color-scale can be reset via the menu *Color → Reset Colormap*

4. In order to take a look at the differences between two images, load them in two different frames. Comparison by eye can easily be done by pushing *Frame → prev/next*.

As for all scientific equipment also the setup of this instrument is constantly being improved. Changes in the hardware and software are planned for the near future. The layout as described above should hence not be considered final.

You should also keep in mind that you are dealing with expensive equipment, which is also in use by other people of our institute to obtain and handle scientific data, so please be cautious. However, you are encouraged to bring in and follow your own ideas (in coordination with your supervisor).

Figure 5: Liquid Nitrogen vessel kept at its designated position in the telescope room. Read and follow the safety instructions displayed in the black panes located above the vessel.

4 Part 1 - Characterization of the CCD

Disclaimer regarding liquid nitrogen

Before handling liquid nitrogen (LN2), students must go through the instructions provided in the telescope room (see Figure 5) and adhere to them strictly. The supervisor will reiterate these instructions to ensure that the experiment is carried out without any injuries. In case of any mishaps, the emergency contact numbers listed on the contact sheet (see Figure 6) should be called immediately. This sheet is displayed right next to desk in the telescope room and also near every emergency exit.

70 CM TELESKOP

Useful contacts

TELESCOPE MANAGER	CONCHI CARDENAS	conchi@mpia.de	457

TOPIC	PERSON	TEL.
EMERGENCY CALL IN-HOUSE		**500**
GENERAL EMERGENCIES	**POLICE**	**0-110**
MEDICAL EMERGENCIES AND FIRE		**0-112**
RECEPTION (only working hours)		06221-528-0

TOPIC	PERSON	E-MAIL …mpia.de	TEL.
CRYO/VACUUM	WERNER LAUN	laun@…	279
CCD	VIANAK NARANJO	naranjo@…	290
CCD ELECTRONICS	RALF KLEIN	klein@…	336
ELECTRICITY/SAFETY	ALEXANDER FRANK	frank@…	250
MECHANICS	ARMIN HUBER	huber@…	331
OPTICS	CONCHI CARDENAS	conchi@…	457
SOFTWARE & IT	FLORIAN BRIEGEL	briegel@…	256
KING MAILING LIST	mailing list	king @…	--
OUTREACH FELLOWS MAILING LIST	mailing list	outreach-fellows@…	--

Figure 6: Emergency contact numbers.

4.1 Preparations

Before any scientific observations can be carried out with a CCD camera it has to be cooled down with liquid nitrogen.
What is the reason for the cooling? (Keep in mind that a CCD is a silicon-based semi-conductor.)
To allow the cooling of the dewar (the "can" containing the CCD chip), it has to be evacuated with a vacuum pump. This is normally already done by the supervisor. After arrival at the telescope you will have to power down the pumps, close all valves and detach the pumping station from the telescope to allow the telescope to be moved.
In the next step, the instrumentation has to be switched on from the console for the telescope controls. Now the electronics which can be controlled via a Linux PC are able to "talk" to the CCD and images can in principle be taken. Before that, however, the readout electronics have to be reset and newly initialized. This is also done from a terminal on the Linux PC *king* via the following command:

```
> ccdread -R -I -C 1
```

The parameters mean the following: -R tells the electronics to reset the CCD, -I to re-initialize it and -C 1 short-circuits the pixels so that they are "empty".
You should now change to your private data directory using the terminal command `cd` (create directory with `mkdir`), where it is possible to make a first exposure with the camera:

```
> cd data/name_of_group
> ccdread -C 0 -f test.fits -E 5.0 -O -C 1
```

The last command tells the CCD to make an image. -C 0 tells the CCD that an image will be taken, -f is followed by the filename (in this case test.fits), -E is followed by the exposure time in seconds (here 5.0 seconds) and -O reads out the CCD. It is important to note that at this stage the detector still has ambient temperature and the coverage of the primary mirror is normally still closed. That means that most of the signal we would see on the image are charged particles that were produced mainly thermally[11]. To visualise the image, the program ds9 is used.

```
> ds9 &
```

The ampersand (&) allows running a program in the background, so that it does not block the terminal window. In ds9, colour mapping can be changed using the right mouse button.
The image shows a strong gradient in read-out direction as those parts that are read out later show a stronger dark current signal. On one side of the image the so called **overscan region** can be seen. This region is not a physical part of the CCD chip but is added electronically to the image. These pixels contain no image information but only the positive offset (**bias**) applied to all detector pixels. This offset is required as the analog-digital-converter can only handle positive

---

[11]cosmic rays might leave also some traces

pixel values and due to fluctuations or bad behaving pixels negative values might otherwise occur. In any case, the bias signal has to be subtracted from all data obtained with the camera in the so-called **bias correction**!

The image can be analysed with the iraf command imexam:

```
> cd ~
> iraf
vocl> cd path/to/your/directory
vocl> imexam
```

On the image display pressing **m** gives statistics on pixels around the cursor position, **s** plots a surface plot, **c** plots a column, **l** a line and so on. Play a little bit around with the first test image and get an idea how the signal varies over the detector and check also the values in the overscan region.

4.2 Required measurements

4.2.1 Properties of the dark current

It was already mentioned above that charges that were thermally excited due to the "heat" of the surrounding environment can cause a signal on the CCD. In this section you will learn how to cool a CCD and how the band gap of the semiconducting CCD material can be derived.

4.2.1.1 Theoretical background   The theoretical dependency of the dark current I from the temperature T can be derived from Fermi statistics and has the following form:

$$I = const. \; T^{\frac{3}{2}} e^{-\frac{E_\mathrm{g}}{2k_\mathrm{B}T}} \tag{14}$$

where $k_\mathrm{B} = 8.617 \cdot 10^{-5} \frac{\mathrm{eV}}{\mathrm{K}}$ is the Boltzmann constant and E_g the band gap of the semiconductor. Measuring the dark current as a dependency of the temperature enables the determination of the band gap of the semiconductor material of the detector and thus the determination of this important characteristic of the instrumental setup. Apart from the decrease of the dynamical range of the detector by the additionally freed charge, the dark current also deteriorates the sensitivity of the system by its stochastic nature (thermal noise).

To reduce the dark current to negligible levels, CCD detectors are cooled. It is possible to suppress the thermal excitation of electrons to less than 0.01 electrons per pixel and second. For astronomical detectors in the optical wavelength range liquid nitrogen, at a temperature of approximately $-196°$C, is used for cooling. Chips used for observations in the infrared have to be cooled by liquid helium because of the smaller band gaps in these devices. To avoid a deterioration of the semiconductor properties at very low temperatures the chip used in this instrument is heated to an optimal operating temperature of roughly $-100°$C. For short exposure times and high luminosities, for example in digital cameras, no cooling at all has to be applied.

4.2.1.2 Execution of the dark current measurements The cryostat should be evacuated before it is cooled. The dewar is filled from the nitrogen container by inserting the rod into the dewar from below. If necessary, nitrogen has to be obtained at the large tank outside the main building and transported into the telescope building in the container. **When handling the nitrogen, glasses and gloves have to be used to avoid serious burning!** Don't worry: Your supervisor will guide you through the cooling process.

Once the dewar is filled with nitrogen you can start measuring the dark current. *What is actually a good measure for the dark current of a CCD with 2048x2048 pixel?* (Keep in mind that all pixels are probably slightly different and some might not be working at all.)

As the CCD is getting cooler and cooler you should read out the chip multiple times to ensure a good sampling of the curve (for this repetitive job it is advantageous if you have a basic knowledge in scripting). The command you need is:

```
> ccdread -C 0 -b 2 -g 5 -T -F dark -E 30 -o dark_current
         -u "name1 name2" -O -C 1
```

The parameters here are: -b 2 leads to a binning factor of 2, meaning that 4 pixels (2 in each direction) will be regarded as one, -g 5 prescribes a gain factor of 5 (equivalent to a certain amplification of the signal), -T saves the current temperature of the CCD, -F is followed by the basis of the filename (a running number will be automatically added), -o gives the possibility to provide a name for the object that is observed and -u lets you specify the observer's names. While the image data is stored in the image part of the .fits-file all the other information is saved into the so-called .fits-header (you can think of it as a "label" for the image). Ask your supervisor, as this step can be easily automated using common programming languages.

4.2.2 Flat-field correction

Both small scale (different sensitivity of single pixels) as well as large scale variations (dust, optical effects) have an influence on the detector signal and thus on the resulting image. To correct for these effects every (scientific) raw frame has to be divided by a so-called master flat-field, which is an image of a structureless, well illuminated (white) surface. By combining several individual images we neglect artefacts on single frames and also reduce the noise. **This correction has to be done for every filter that will be used for the scientific observations** as it accounts for small imperfections (e.g., dust) in the optics.

4.2.2.1 Execution of the measurements To obtain flatfields, point the telescope to the white surface hanging on the inner side of the dome by pressing the RA and DEC buttons on the control console. Make sure that the cover for the primary and secondary mirror are open and take a first test flat-field image with the command:

```
> ccdread -C 0 -b 2 -g 5 -T -F flat -e 30 -o "flat_filter_"
```

```
            -u "name1 name2" -O -C 1
```

While before we had an upper case -E followed by the integration time in seconds we now have to type a lower case -e followed by the integration time. The difference is that with -E the shutter within the camera is not opened (good for taking dark frames) while the -e opens the shutter and photons can be detected. To obtain a sensible signal at longer integration times the illumination inside the dome has to be dimmed. You should test the optimal setting for each filter before starting the measurement. To obtain a master flatfield frame, take at least five exposures at relatively long integration times to achieve a good signal to noise ratio for each filter. However, the maximum pixel value across the array should still lie well within the region of linearity.

4.2.3 Linearity and dynamical range of the CCD

To quantify the sensitivity to incoming photons and the limitations of the chip the linearity and dynamical range have to be analysed. For this, we measure the detector signal for flat-field images for different illuminations (in this case for different exposure times) in one filter. The filters can be changed with the movable filter-frame located above the CCD dewar.

4.2.3.1 Execution of the measurements The setup is similar to the one for the master flat field (section 4.2.2.1), but now we take exposures of the uniformly illuminated area with different integration times. The collected charge and hence the signal should be proportional to the integration time, assuming the illumination is proportional to the integration time. In total, you should take about 15 images with increasing exposure time, of which the last 2-3 should already be in the saturated area.

4.2.4 Sensitivity of the detector and noise properties

According to Poisson statistics, the variance of the pixel-to-pixel fluctuations in a uniformly illuminated area scales with the number of incoming electrons per pixel, N_e, such that

$$\sigma_e^2 = N_e \tag{15}$$

For astronomical data reduction, the gain, κ, that is the number of excited electrons per unit of the analogue digital converter (ADU), is an important characteristic of the system. It is a required parameter in signal processing. The final signal $N_{e,d}$ in ADU is connected to the original number of incoming photons N_e by

$$N_e = \frac{\kappa}{\eta} N_{e,d}, \tag{16}$$

where η is the quantum efficiency of the detector. Similarly,

$$\sigma_e^2 = \frac{\kappa^2}{\eta^2} \sigma_{e,d}^2. \tag{17}$$

Using the above relations, we get

$$\frac{\kappa}{\eta} N_{e,d} = \frac{\kappa^2}{\eta^2} \sigma_{e,d}^2, \tag{18}$$

so that,

$$\sigma_{e,d}^2 = \frac{\eta}{\kappa} N_{e,d}. \tag{19}$$

Two other sources of noise exist in the data aside from the photon noise. The read-out noise of the gate amplifier, σ_R, and the so-called Pixel Response Non-Uniformity (PRNU) noise, σ_{PRNU}, which occurs due to the fact that the quantum efficiency of the pixels are not exactly identical. The PRNU noise increases linearly with signal level such that

$$\sigma_{PRNU,e} = N_e f_{PRNU}, \tag{20}$$

where f_{PRNU} is the detector-dependent, characteristic PRNU factor, which is typically on the order of 0.01. Therefore, for this typical value, the photon noise and PRNU noise are exactly the same at a signal level of 10 000 electrons, ($\sigma_{PRNU,e} = 0.01 \cdot 10\,000 = \sigma_e$), and PRNU noise dominates at high signal levels. The total noise of the measured signal in ADU is given by

$$\sigma_{tot,d}^2 = \sigma_{e,d}^2 + \sigma_{R,d}^2 + \sigma_{PRNU,d}^2 \tag{21}$$

Since the underlying cause of PRNU noise is a physical one and directly proportional to the signal level, two frames with the same signal level should have the same PRNU noise. Therefore, by subtracting the two flat frames with the same signal level, one eliminates the effect of PRNU all together. The only source of noise left in this difference image, σ_{diff}, is then a combination of read-out noise and photon noise, such that after propagating the errors,

$$\sigma_{diff,d} = \sqrt{2} \cdot \sqrt{\sigma_{R,d}^2 + \sigma_{e,d}^2}. \tag{22}$$

Combining eqs. (19) and (22), and assuming $\eta = 1$, one can derive

$$\sigma_{diff,d}^2 = 2 \left(\sigma_{R,d}^2 + \frac{N_{e,d}}{\kappa} \right). \tag{23}$$

4.2.4.1 Execution of the measurements The setup is similar to the one for the linearity measurements (4.2.3.1), but now we take two measurements of the uniformly illuminated area with the same integration time for a range of exposure times in order to remove the effect of PRNU. In total, you should take about 20 pairs of images with increasing exposure time in the I-filter, of which the last 2-3 should already be in the saturated area.

4.3 Analyses

4.3.1 Analysis of dark current measurements

1. Use `Astropy` and `glob` to read in and extract the temperature information from the headers of the FITS files of all dark measurements. Be efficient and build a loop over all files found by `glob` (that applies to almost all the next tasks).

2. Use `numpy` to determine the median bias values and the standard deviations in the overscan region for all dark measurements (check one of the FITS files with ds9 to get the range in pixel values of the overscan region which you can use for slicing the arrays).

3. Correct each of the dark measurements with their corresponding median bias value and save them with `Astropy` (you should give them meaningful names to indicate that these are the bias-corrected files e.g., by adding _b to the FITS file names.). Don't forget to always subtract the bias from all of your measurements also later on!

4. Extract the median dark current values of the bias-subtracted FITS files of the dark measurements with `numpy`. (use ds9 again to get the proper range in pixel values of the image region of the CCD).

4.3.1.1 Tasks and discussion points for the protocol

- Verify the dependency according to eq. 14 by plotting the data and the theoretical curve in a sensible plot (see Figure 7).

- Determine the band gap of the detector chip by fitting the dependency.

- Determine the bias and its scatter for two images, one at the beginning of the experiment and one at the end.

- Why do we have to determine and subtract the bias value for every image individually?

- Discuss reasons for possible deviations from the expected band gap value.

4.3.2 Analysis of Flat-field correction

1. Use `Astropy` and `numpy` to subtract the bias from the individual flat-field images similar to the dark current correction in the previous section.

2. Combine the individual flat-field images to a single image (the master flat-field) (Tip: Store all images in a single 3-dimensional array and take the median along the primary axis.)

3. Normalize the master flat-field to one by dividing it by its median.

4. Use `matplotlib` to obtain a histogram for the normalized master flatfield.
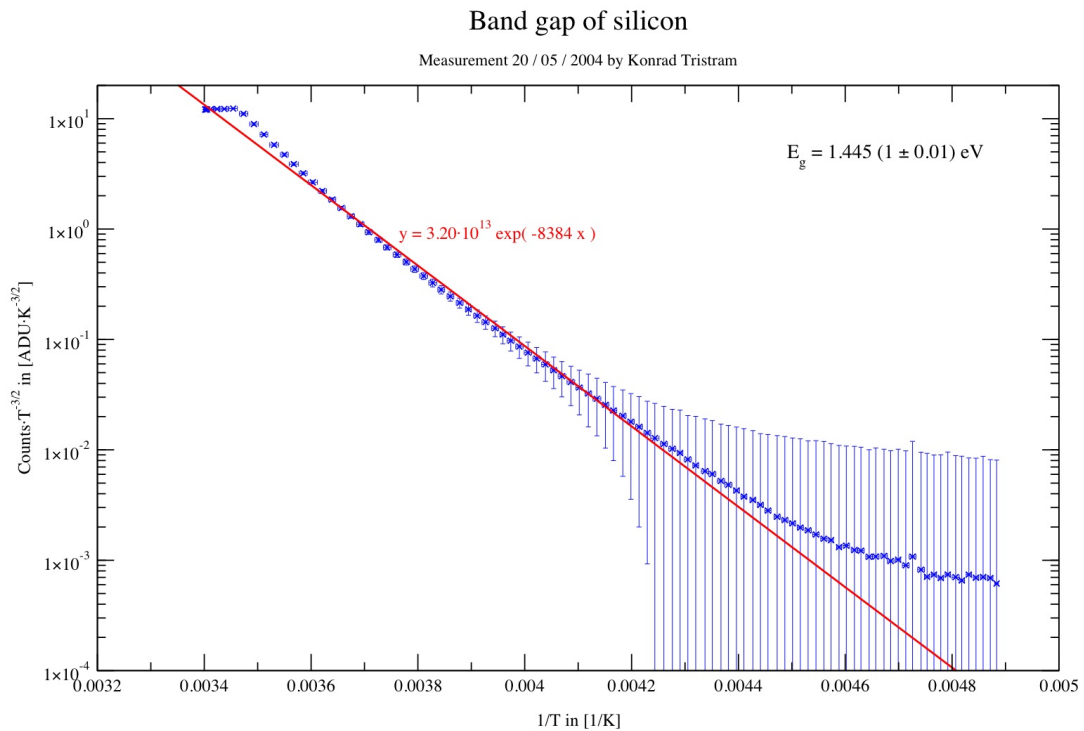
Figure 7: Determination of the band gap E_g for the detector according to formula 14. To better see the deviation of the measured signal from the theoretical behaviour the dependency has been transformed into a linear dependency. The band gap of this measurement determined by Konrad Tristram on 20/05/2004, $E_g = 1.445\,(1 \pm 0.01)\,\mathrm{eV}$, is considerably higher than the band gap of silicon, $E_{g,\mathrm{SI}} = 1.15\,\mathrm{eV}$.

4.3.2.1 Tasks and discussion points for the protocol

- Perform a flat-field correction for one of the single flat-field images (i.e., divide the image by the normalized master flat-field). Plot and compare the image before and after the flat-field correction. (Tip: the FITS images can be displayed and saved with `matplotlib` in an interactive plot window, or using ds9.)

- Plot and discuss the histogram for the normalized master flat-field.

- Why do we have to normalize the master flat-field?

- How many pixels of the master flatfield have a value close to 1 as expected and where do the deviations come from?

- Why do we need to divide all of our observations by a master flat-field?

- Why do we use the median of the pixels for the image combination and for the normalization and not, for example, the mean value?

- Why was it necessary to set the illumination differently for each filter?

- Why do we have to measure the CCD response to a "flat" surface? Why are the nails holding the blanket and the warped surface of the blanket itself not detrimental to the dome flat?

4.3.3 Analysis of linearity and dynamical range

1. Use `Astropy` and `numpy` to subtract the bias from the images similar to the dark current correction in the previous section.

2. Perform a flat-field correction for all images by dividing them by the normalized master-flat from the previous section. (Tip: Store all images in a single 3-dimensional array and build a loop over them.)

3. Use `Astropy` to extract the exposure time information from the headers of the FITS files.

4. Use `numpy` to get the values for the median and the standard deviation for all the bias- and flat-field-corrected images.

4.3.3.1 Tasks and discussion points for the protocol

- Verify the linearity and the dynamical range of the chip and determine the linear regime (for both filters) by plotting the signal of each flat-field image against its integration time.

- Determine the deviation from a perfect linear relationship and specify the intensity where saturation sets in (see Figure 8). Use only those frames that are not saturated for the fit. Estimate the error of the measurement from the standard deviation of the pixel values.
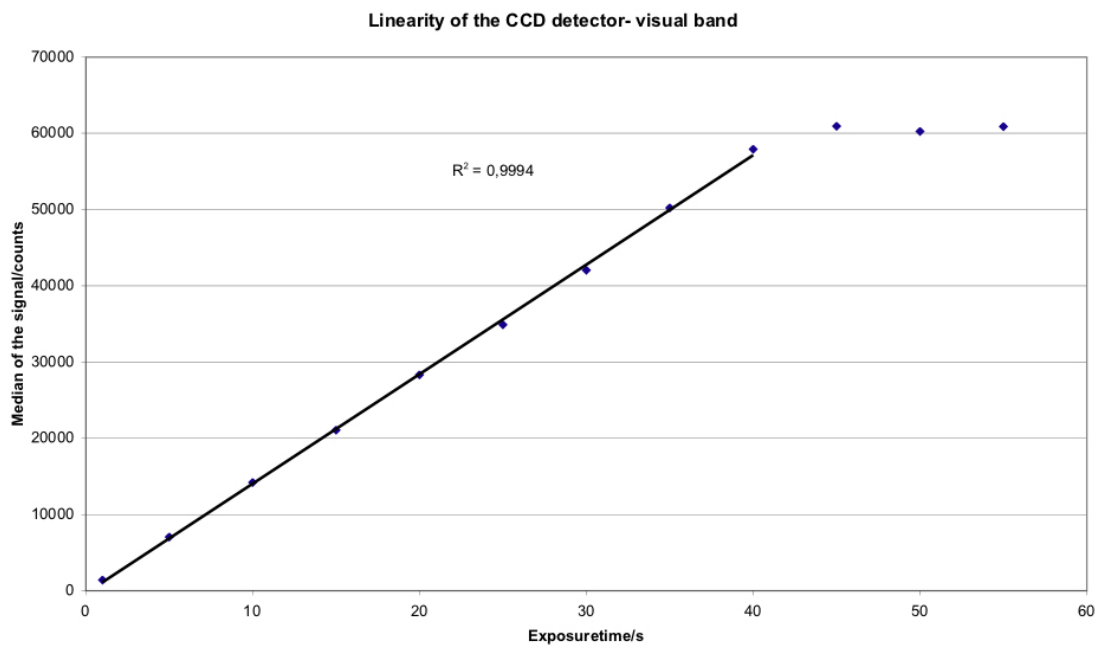
Figure 8: Linearity and dynamical range of the detector. The illumination is assumed to be proportional to the integration time. The plot shows that the detector signal is proportional to the integration time (and hence the illumination) with less than 0.5 % deviation. The detector saturates at approximately 60000 counts for a gain factor of 5.

4.3.4 Analysis of sensitivity and noise properties

1. For one pair of images where on average the signal does not exceed ~ 30000 counts. Determine the read-out noise, $\sigma_{R,d}$, from the overscan region by using `numpy`.

2. Use ds9 to choose at least a 200×200 pixel region that is uniformly illuminated near the center of the image. *It is important to choose a uniformly illuminated area because in the following steps we are working with the raw images and will not be performing any bias-subtraction or flat-field corrections!*

3. Use `np.std` to determine the total noise, $\sigma_{\text{tot},d}$, and `np.median` for the median signal, $N_{e,d}$, in each of these two images within the chosen regions.

4. Use `numpy` to subtract the signal of one image from the other. **Caution! The fits files from the CCD camera are unsigned 16 bit files and also interpreted as such by `astropy.io` Convert the read in image with:**

```
image = image.astype('float64')
```

before the subtraction.

5. Determine the noise in the difference image within the chosen region, $\sigma_{\text{diff},d}$, by using `numpy`'s `std`.

6. For all pairs of images, use `numpy` to extract the median signal, $N_{e,d}$, within the chosen regions for each pair of images. You can choose to do this for one of the images in each pair, or for both images and take the average of the two values.

7. Subtract the signal of one image from the other **in each pair** of images. (Tip: use `glob` to create two lists of filenames, one for each image in a given pair. Then use `numpy` to subtract one from the other. Be careful to do the subtraction for images with the same exposure time.) For **all pairs** of images, use `numpy` to extract the noise in the difference image within the chosen region, $\sigma_{\text{diff},d}$, by using the function `std`.

4.3.4.1 Tasks and discussion points for the protocol

- Find and discuss the photon noise, the read-out noise, and the PRNU noise for one flat image. Which noise dominates? Explain.

- Plot the variance of the signal versus the signal itself for all measurements with increasing exposure time in the I-filter and extract the electron sensitivity κ using eq. (23). Explain what this plot means.

- Calculate the electron sensitivity κ using eq. 19 and explain discrepancies with the previous method. Do the calculation for just one image with average counts of $\sim$30 000 ADU.

- Discuss the quality of the analogue digital converter.

5 Part 2 - Photometry of globular clusters

In this section you will use either your own KING data or archival images of the Hubble Space Telescope (HST) of the globular cluster BS90[12] to produce a color magnitude diagram that will allow you to estimate the distance, age, and metallicity of the cluster.

Overview of the tasks

1. Determine the zero point for each image (see section 5.1)

2. Perform PSF fitting for each image (see section 5.2)

3. Cross-match the PSF fitting results and plot the CMD. Fit isochrones to it in order to estimate the distance, age, and metallicity of the globular cluster (see section 5.3)

5.1 Zeropoint calibration

As for all physical measurements, a careful calibration is essential for the detailed analysis of an astronomical observation. It is necessary to calibrate every measurement to a standard scale, since each observational facility and detector will have different characteristics. For astronomical observations this is the apparent magnitude scale. In order to get the zero-point magnitude we have to compare the counts of standard stars[13] in the Hubble images with reference values of the apparent magnitude from an astronomical data base called SIMBAD. For this purpose, open the V and I image of BS90 with ds9 and overplot one of the images with the catalog "Cool evolved stars in SAGE-SMC and SAGE-LMC (Boyer+, 2011)".
The catalogue can be loaded into ds9 as follows:

1. *Analysis → Catalogs → Search for Catalogs*

2. Set *Name or Designation* = J/AJ/142/103

3. Click *Retrieve*

4. One catalogue will show up in the list in the window. Select it and then click on *Catalog*

---

[12]The images were taken with the HST filters F555W (V) and F814W (I).

[13]The reference stars we will use are not really standard stars, in the sense that they were calibrated themselves from true standard stars. This is due to our limited field of view, so the probability of a true standard star being located there is very low.

5. Now the objects will be overplotted on the image. On the pop-up window, you can sort the list after I-magnitudes, such that you can look at only those: $Sort = $ Imag

Now you have to select ~ 10 stars for which you write down the V- and I- magnitudes and measure the counts in **both** Hubble images. To measure the counts you have to define a region around the reference star (with *Edit $\rightarrow$ Region*; click into the image to create a region that can then be dragged to a different location). To get the right counts you have to make the region the right size, so that you get all the flux of the star without including any flux of another object nearby. If two stars are too close together choose another reference star instead. To get to the region statistics you have to double-click on the region and then go in the region menu to *Analysis $\rightarrow$ Statistics*; *sum* gives you the counts and *error* the standard deviation. You can copy and paste the region to the other Hubble image and repeat the count measurement (you may have to slightly change the radius of the region).

After you have determined the counts and apparent magnitudes from the catalogue entry, you can compare your instrument magnitude with the CATALOG objects to calculate the individual zero points for all ten objects for both filters with the following equation:

$$\text{zeropoint} = m_{CATALOG} + 2.5 \log_{10}(\text{counts}) \tag{24}$$

Take the median to get your final zero point values for both filters and if necessary neglect extreme outliers.

5.2 PSF photometry with Python

In crowded regions, for which the PSFs of different stars overlap, aperture photometry does not work any more. Therefore, we use `IterativelySubtractedPSFPhotometry` from the `photutils` package (Bradley et al., 2019) for Python. The task is especially designed to perform PSF photometry on crowded fields. First, we need to extract a statistical PSF from the image and then try to fit all stars in the image with that PSF. To extract a PSF from the image the position of $\sim$ 10 - 30 isolated, unsaturated stars is needed. The task `EPSFBuilder` from `photutils` is used for building an effective PSF (ePSF; Anderson and King, 2000) by combining the postage stamps of the selected PSF stars into an empirical oversampled PSF model that describes what fraction of a star's light will land in a particular pixel (`https://photutils.readthedocs.io/en/stable/epsf.html#build-epsf`). To detect all the stars in the image, `photutils`' `IRAFStarFinder` is used. The task makes use of the DAOPHOT MMM algorithm to determine the background of the image. Simplified, the algorithm estimates the true sky value by $3 * \text{median} - 2 * \text{mean}$ of the image. In an iterative process, `IterativelySubtractedPSFPhotometry` tries to fit all potential stars with the

PSF and subtracts the results of this fitting from the science image. After all stars are analysed and subtracted, the background is recalculated and subtracted from the original science frame, after which the next iteration begins (i.e., another fitting of the PSF to all potential stars of the new background-corrected science image) That way we can disentangle and correctly account for the flux of overlapping sources. However, this yields only the instrumental magnitude of these sources. To get the correct apparent magnitude, you have to use your calculated zero-point values from the previous section.

5.2.1 Using the `jupyter-notebook`

To simplify the work flow of this part of the experiment, a `jupyter-notebook` is provided. To open it, type:

```
cl> cd
cl> jupyter−notebook
```

in the terminal and browse from the homepage to the folder "CMD" in the directory that was created by your tutor for you. There you will find a file named "PSFPhotometry.ipynb". Once the JPN is opened, you can simply follow the steps given there and which will be described in a bit more detail below.

Definition of fit parameter

Before you can start, you need to specify some parameters.

1. First you need to specify the filter for which you want to perform the photometry (must be either "V" or "I").

2. Then you give the zero-point magnitude that you determined in the previous section for this filter. It is used to convert the flux measurements to the apparent magnitude.

3. Lastly you need to define the size of the PSF in the images. Use logarithmic scaling in ds9 to find the radius in pixels that contains all of the features of the PSF of a rather bright star.

Loading the data with astropy

After you set all parameters, you can simply run this cell to read in the image data to the "image" variable.

Select the stars to determine the PSF

1. Choose stars from the image: Running this cell will open a figure in which you can pan and zoom the image to select the star which will be used for determining the PSF. Double-click the left mouse button to select a star. You should mark a set of 20-30 rather isolated and unsaturated stars in

order to get a good PSF. Double-click the right mouse button to close the figure and proceed with the next cell.

2. Evaluate choosen stars: Start with the first cell. It will open a figure, in which you can evaluate your previous selected stars. Select only stars with a rather symmetrical brightness distribution of the PSF and which are not overlapping with other stars. If you find less than 15 stars to be suitable, please repeat step one and preselect more/other stars.
   Please run the next cell after the figure closed. It will mask out stars, which are too close to the edges of the image and save the star positions in a format which is suitable for the next step.

- *NOTE: your selected stars will be saved in a file, so you can also reload stars from a previous run and skip the first to parts.*

Build PSF model

1. Run the first cell to create a PSF model out of your selected stars.

2. Run the second cell to visualize the obtained PSF model. You can tune the "min_max_percentage" parameter to display as much details as possible. If you find something strange, you can try to select other stars and repeat the previous steps. Save an image of the PSF for your records.

Performing iterative PSF photometry

Finally you perform the photometry on the image using your determined PSF model. Depending on the threshold you set, the execution of this task takes up to 30 min.

5.3 Cross-match the PSF fitting results and plotting the CMD

For the CMD we plot the V vs the V-I magnitudes, which means that only stars that were found in both images by your photometry can be used. The cross-match and the plotting are executed in the `jupyter-notebook` "Crossmatch_and_plotting_CMD.ipynb". Determine which isochrone fits the distribution of stars best by playing around with the age and metallicity values. You also have to determine the right shift in the V axis, which will then give you the distance to the cluster.

1. Load the results from the previous section by running the first cell.

2. Next the cross-match of the sources is performed. Depending on the number of fitted sources, this may take a while. Once you did the cross-match you can reload the cross-matched sources later by simply running the second cell instead.

3. When the cross-match is finished you can proceed to plot the CMD. But before you can create the figure, you must set the parameters for some theoretical isochrones that are overplotted on your measured data. Additionally you need to specify the magnitude offset between the theoretical isochrones (which are by default in absolute magnitudes) and your measurements.

4. Plot the CMD and save it for your records.

6 Optional part - Scientific observations

In the steps described in the last section you have analysed the technical aspects of the detector and should have verified that the properties of the instrument are suitable for astronomical research, so that you can now proceed to take your own observations. However, due to the generally unfavourable weather conditions on the Königstuhl this may not be possible. If that should be the case, your supervisor will instead provide you with archival data to work with.[14]

6.1 Tasks

- Find a suitable globular cluster for the observations.

- Observe the globular cluster and relevant calibration data in two filters.

6.2 Execution of the observations

An astronomical observation has to be planned carefully in advance. To decide on which cluster to observe by you have to consider its position on the sky during the possible observation time. It is advisable to make a detailed plan of all the steps needed for the observation to allow a smooth work flow during the night and avoid wasting precious telescope time. You should think about the observing method to optimize the signal to noise and eliminate artefacts in the raw images.

Before the actual observations can be performed the telescope has to be focused. There are two buttons in the console for the telescope controls changing the position of the secondary mirror and thus the focus. Point the telescope to a clear part of the sky, make a test exposure, display the image on the computer and check whether you find an appropriate number of stars in the image. If that is the case set the focus to an arbitrary initial position and increase/decrease the focus for subsequent exposures. Write down the focus positions in your protocol. Take care to change the focus position only in one direction as the positioning mechanism exposes the behaviour of a hysteresis. With the IRAF task `starfocus` you can then analyze the images and determine the optimal position of the focus.

For the calibration of the magnitudes you also need observations of standard stars for all the filters with which you observed the globular cluster. Don't forget these

---

[14]It is very common in astronomy to use archival data and many new discoveries have been made from the re-analysis of "old" data.

as otherwise all your measurements are useless! Your supervisor will be able to help you find a suitable field for standard stars.

6.3 Observation

6.3.1 Introduction

The majority of the visible stars are in the hydrogen burning phase and thus populate the main sequence in the Hertzsprung-Russell-Diagram (HRD). In a stellar population with a defined age, like a globular cluster (GC), it is possible do deduce its evolution history, its age and its distance from the form of its HRD. Producing an HRD in practice is rather difficult since both absolute brightness and spectral type of a star are often not exactly known. In observational astronomy it is therefore common to use a color-magnitude-Diagram (CMD) in which the apparent magnitude, e.g. V, is plotted against the color, e.g. B-V. These values are possible to determine directly from observations. This is the goal of this experiment. According to the time of year, ergo, the visibility on the sky and some properties (size, magnitude) one GC is chosen by the assistant. Its images are then taken with the 70cm KING telescope in two wavelength ranges, B and V (corresponding to the Johnson-System). After "reducing" the images to their scientific content and their evaluation, the measured instrumental magnitudes of the stars in the GC can be translated to the standard apparent magnitudes which produce its CMD. By fitting the theoretical model-isochrones ("lines of equal age") to the plotted CMD, the age and the distance of the GC can be determined.

6.3.2 Data acquisition

Before taking images of the GC some preparatory work is necessary. In general, "raw" images coming from the telescope have instrumental defects which we have to correct. They are removed later from the images, enabling an exact calibration of the photon fluxes across the array. For that, images of two kinds have to be acquired - "Bias" and "flat-field". In the following it is described how to do that.

6.3.3 Bias

To measure the bias perform 5 integrations of 5 seconds each with the shutter closed. Pay attention to the position of the shutter during the integration. Is it necessary to measure the bias with each filter?

6.3.4 Flatfield

For the sake of redundancy and in order to remove single pixels saturated by cosmic rays, take at least 5 dome-flat exposures with each filter (see section 4.2.2). A "master flat-field" can then be constructed by combining these single flat-fields and calculating the median for each pixel. Each scientific exposure has to be divided by the normalized master-flat to correct for its small- and large-scale pixel-to-pixel variations.

If the weather and time should allow it, another method to perform flatfield calibration is to make images of the twilight sky when the Sun is already below the horizon but the sky is still not completely dark. The available period for doing this is relatively short - if it is still too bright the CCD saturates immediately; if it is already too dark, the low brightness of the sky will necessitate longer integration times, causing stars to appear on our flatfield image. Nonetheless, the sky-flatfield is preferable to the dome-flatfield since the CCD is illuminated in the same way as in the scientific exposures. It also avoids possible effects introduced by the artificial dome lights, such as weak patterns of scattered light and non-uniform illumination.

Sky-flatfields are performed analogous to dome-flatfields. Even though the twilight sky may seem homogeneous to our eyes, it is probable that the CCD will detect some bright stars. To avoid detecting them always with the same pixels, move the telescope after each exposure approx. 10 arcsec in the same direction along one coordinate. When deciding with which filter to start (considering the limited time), take into account that the colour of the sky is blue.

6.3.5 Finding the focus

After we have taken the necessary calibration data, we have to set the focus, i.e., align the position of the focal plane of the telescope with the plane of the CCD-Chip. For that we slide the secondary mirror along the optical axis by pressing the + or - "focus" button on the control console. But first, we should make a short "acquisition image" of the night to make sure that a couple of sufficiently bright stars are visible within the field of view. If that is the case, we can perform a series of 7 measurements, starting with one position of the secondary mirror and scanning a value range of approx. 20 (as displayed on the control console). Write down which value displayed on the control console corresponds to which image name. That is necessary to find out which position of the secondary mirror corresponds to the best image. The optimal focus position can be calculated with the IRAF task "starfocus" which takes the images and measures the point-spread function (PSF) width of stars (selected interactively) in each image and makes a parabolical fit to the values. This has to be done for all filters that will be used in the measurements.

After the focus is set, the actual imaging of the GC can be performed. For that, point the telescope to the sky coordinates of the GC. Because the pointing accuracy of the KING telescope is not perfectly calibrated there is a small offset between the coordinates displayed on the control console to the real sky coordinates. Therefore, we first have to make a test image to find the necessary shift we have to apply to the coordinates so that our object will be centered in the image. This test image is also useful for deriving the integration time for a single exposure with the given filter. Given the linearity of the CCD, an integration time should be chosen at which the central (= brightest) region of the GC will be at the upper end of the linear range of the detector. This will optimise the efficiency of the observations since the time needed to read out the CCD is relatively long (approx. 1 min for

resolution 2048x2048) and independent of the exposure time.

Finally, we can start to image the GC. In order to avoid the negative effects of observing the same sky area with the same pixels (possibly "bad pixels" or pixels locally saturated by cosmic rays), we use an observing method in which the object is moved ("dithered") across the array by a few arcseconds after each exposure, so that in each image the pixels observe a different part of the sky or object. Hence, for every new exposure, the telescope has to be repositioned by approx. 10 arcsec according to the chosen observing pattern. This pattern can be a triangle, a rectangular, a spiral, a cross, etc.

The derived apparent magnitudes can now be plotted in a colour-magnitude diagram (e.g., V vs. V-R) and evaluated with the help of model isochrones. Shifting the isochrones along the V-axis allows to choose the one that fits the CMD best, which gives the estimated age of the GC. Since the isochrone units are absolute magnitudes, the comparison of the positions of the fitted isochrone on the vertical axis allows to calculate the distance modulus m-M. Given the weather conditions on Koenigstuhl, usually the quality of the resulting CMD is not perfect. To achieve better fitting constraints, it is recommended to make separate histograms of V-R for a certain slice in V to identify the average value of V-R and its dispersion.

7 Glossary of CCD commands

The CCD can be controlled from the command line prompt on the *king* computer.

7.1 Basic commands

Start CCD:

```
> ccdstart
```

Stop CCD:

```
> ccdstop
```

Check temperature:

```
> ccdtemp
```

Make exposures (see below):

```
> ccdread …
```

7.2 Resetting the electronics

```
> ccdread -R -I -C 1
```

-R reset the electronics

-I initialize the electronics

-C 1 short circuit the pixels

7.3 Taking a dark exposure

```
> ccdread -C 0 -b 2 -g 5 -T -f dark.fits -E 60
         -o "dark_current" -u "name1 name2" -O -C 1
```

-C 0 …

-b 2 set the binning factor to 2,i.e. 2 pixel in x-direction and 2 pixel in y-direction will be binned into one single pixel.

-g 5 set the detector's gain factor to 5

-T write the temperature into the header of the .fits image file

-f specify the full file name hereafter

-E specify the integration/exposure time in seconds hereafter (shutter will remain closed!)

-o specify the object's name hereafter

-u specify the observer's name hereafter

-O read out the CCD

-C 1 short circuit all pixels

7.4 Taking a flatfield or science exposure

```
> ccdread -C 0 -b 2 -g 5 -T -F filename -e 60
         -o "flatfield" -u "name1 name2" -O -C 1
```

-C 0 ...

-b 2 set the binning factor to 2, i.e. two pixels in x-direction and two pixels in y-direction will be binned into one single pixel.

-g 5 set the detector's gain factor to 5

-T write the temperature into the header of the .fits image file

-F specify the basis of the file name; a running number will be attached automatically, e.g., filename0001.fits

-e specify the integration/exposure time in seconds herafter (shutter will be opened!)

-o specify the object's name hereafter

-u specify the observer's name hereafter

-O read out the CCD

-C 1 short circuit all pixels

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