

Task 11: Spectroscopy - Data Reduction and Interpretation

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1 Introduction

While many different techniques have been established in observational astronomy, spectroscopy counts unquestionably as one of greatest importance, and it has been a major key for understanding the properties and physics of many different cosmic objects, such as stars, nebulae, galaxies, and others. The basics of the method go way back to Isaac Newton's optical experiments between 1666 and 1672 and his findings that, using a diffractive element, light can be decomposed in contributions of different wavelength. Still, only since the discovery of stellar lines in the spectrum of the Sun by Joseph v. Fraunhofer in 1813, and their explanation based on the *spectral analysis*, developed by Robert Bunsen and Gustav Kirchhoff, spectroscopy arrived at its full potential. By the means of spectroscopy today's astronomers are able to probe objects far away, and the night sky has become a laboratory. The observed spectra are used to investigate different properties of the emitting objects, such as their temperatures and/or chemical composition. They also allow to determine the objects' radial velocities¹, which can be used to map the dynamics within the Milky Way, to discover planets around other stars, or to measure the cosmological redshift, one of the key evidences for the expansion of our universe.

In this course you will learn, how this kind of information can be derived from the raw spectral images, as recorded during an astronomical observation. Equipped with the proper tools (software) you will find out, which steps need to be undertaken during the data reduction in order to correct for instrumental and environmental effects and to arrive at a fully extracted one dimensional spectrum. You will be provided with data of different objects, with distinct spectral appearance in order to learn, how to discriminate them.

2 Spectroscopy

2.1 Spectroscopy in Astronomy

Typically, as depicted in Fig. 1, an astronomical spectrum consists of contributions from three different spectral features:

¹The velocity component parallel to the line of sight.

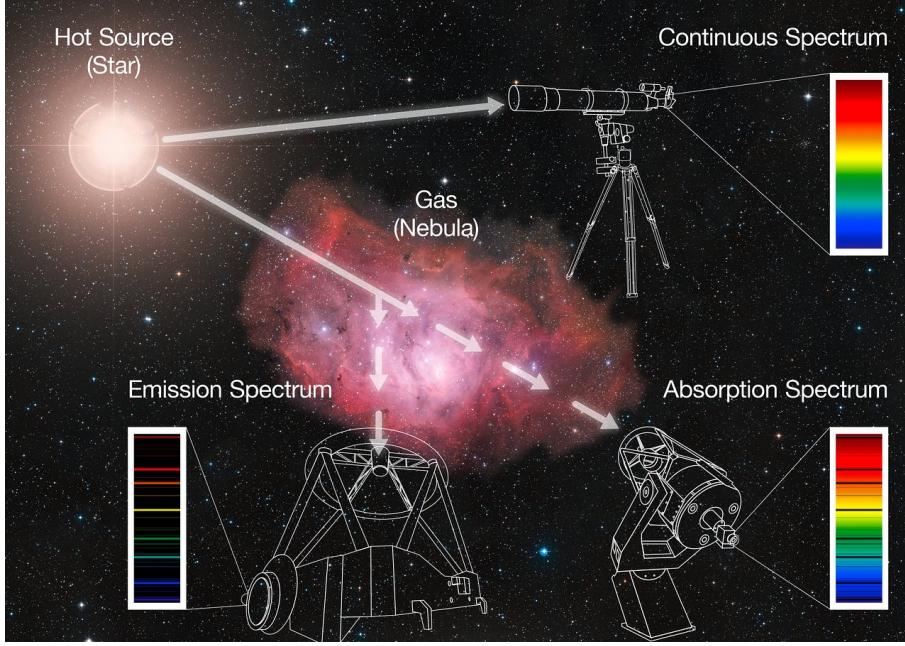


Figure 1: An observed spectrum usually consists of a continuum, superimposed by absorption and emission features. Which of those components dominate, depends on the properties of the light source and the matter along the line of sight.

- The standard black-body radiation, emitted by sources with a specific temperature, is observed as a **continuum**. Following Planck's law it shows a characteristic flux distribution. As those spectra peak at a specific position in frequency space, which depends alone on the bodies' temperature, they can in turn be used to determine that property.
- **Absorption** features are observed in spectra, when matter, in general colder than the initial light source, is situated along the line of sight between the source and the observer. Photons of specific wavelengths excite electrons to higher energy levels, if the difference in energy between the levels ΔE equals the photon energy hc/λ . These distinct absorption features serve as *fingerprints* of the absorbing material.
- When gaseous matter is observed without a hot light source in the background along the line of sight, it shows a specific **emission** spectrum. This again is created when excited electrons move from higher to lower energy levels, and the energy is emitted in form of electromagnetic radiation.

Therefore, the analysis of spectral features allows conclusions on the observed target's temperature, its chemical composition, and also its ionization states. Energy transitions in molecules due to rotation and/or vibration states are also observed, although typically at lower energies, at wavelengths in the near-infrared (NIR) or the infrared (IR) of the electromagnetic spectrum. Those lines are often tightly packed, leading to the observation of the so called *molecular bands*.

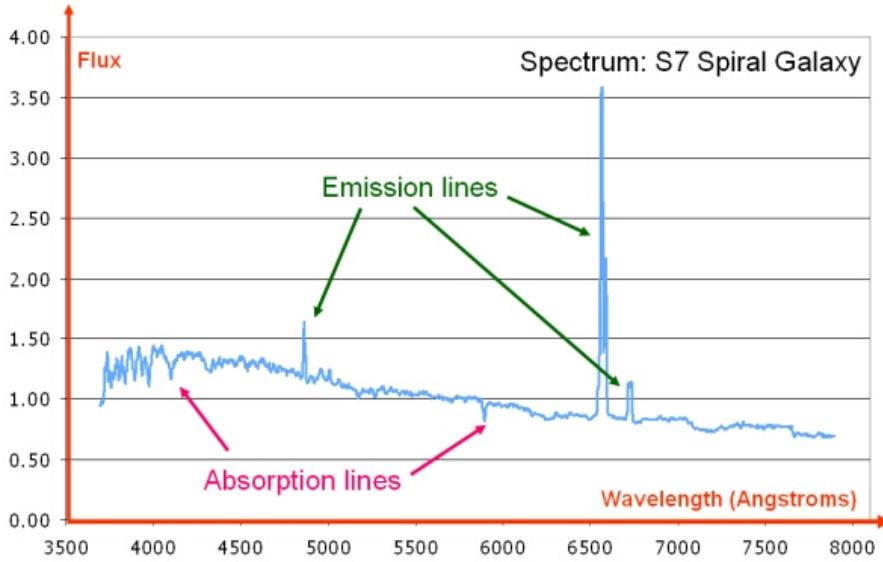


Figure 2: This spectrum of an S7 spiral galaxy shows emission and absorption line features on top of the continuum. Credit: Swinburne University of Technology, based on Santos et al. (2002).

2.2 Basics of a Spectrograph

In order to record an astronomical spectrum, the telescope needs to be equipped with a spectrograph. Various kind of spectrographs are used today. They differ in the ways, how the light is injected into the instrument and therefore in the optical elements used, the choice of the dispersive element, which redistributes the incoming white light along the wavelength axis in the image frame, and also in the camera that is used to record the image. As in photometry, in these days Charged Coupled Devices (CCDs) serve as detectors to store the final raw image (see Sec. 2.3). Still, all spectrographs have some main components in common that are presented in the following, using a standard slit spectrograph as an example (see Fig. 3). This is the kind of instrument that has been used to collect the data for this course:

- The **slit** serves as the entrance pupil of the instrument. It is located in the focal plane of the telescope and defines, which light enters the device. During target acquisition the object to be observed is placed on top of the slit, which is depicted in Fig. 4. Only light along the slit is fed into the spectrograph. While in the following the spectra are formed along the dispersion direction in the final image, the slit is mapped along the spatial direction.
- A so called **collimator** is used to parallelize the incoming light rays (or equivalently to flatten the wave forms). This is needed in order to homogenize the angles of incidence at the dispersive element to prevent blurring of the spectrum.
- The **dispersive element** is the main part of the spectrograph. Behind this element the initially white light is ordered with respect to a diffraction pattern, specific to the used element. While different kind of elements can be used, they all have in common that the diffraction angle of the outgoing beams are wavelength dependent. Historically,

simple prisms were deployed for this purpose. In today's applications, diffraction gratings, or the combination of both, the so called *grisms* are usually utilized. In those cases the diffraction pattern (Fig. 5) follows the grating equation

$$d(\sin \theta_i \pm \sin \theta_d) = m\lambda, \quad (1)$$

where d is the grating's groove spacing, θ_i the angle of incidence, θ_d the diffraction angle, m the spectral order, and λ the wavelength. The (+) on the left hand side of the equation applies for reflective gratings, and the (-) for transmissive gratings respectively. The angular dispersion, the angular spread of a spectrum of order m between the wavelength λ and $\Delta\lambda$, is then obtained by differentiating the grating equation and assuming the incidence angle θ_i to be constant:

$$\frac{d\theta_d}{d\lambda} = \frac{m}{d \cos \theta_d}. \quad (2)$$

The dispersion increases at higher spectral orders and with decreasing groove spacing. The choice of the right angular dispersion depends on many factors, such as the dimensions of the detector, the integrated camera optics, the desired wavelength coverage, as well as the resolving power one strives for. The spectral resolution (or resolving power) R is a measure of the instrument's ability to separate adjacent spectral lines. Together with the *limit of resolution* $\Delta\lambda$, which is the wavelength difference of two lines that still can be distinguished from one another (see Fig. 6), the resolving power is defined by:

$$R = \frac{\lambda}{\Delta\lambda}. \quad (3)$$

Spectrographs with a resolving power below 1,000 are typically called low-resolution instruments, whereas instruments with $R > 10,000$ are often referred to as high-resolution spectrographs.

- The final spectrum is imaged by the optics of the **camera** onto the **detector**. While in former times photographic plates have been in use, these days CCDs are the current standard. Together with the camera's optics the dimension of the CCD limits the wavelength coverage of the observed spectrum, while the pixel resolution defines the sampling.

2.3 Charged Coupled Devices - CCDs

Astrophysical measurements in the optical wavelength regime are nowadays usually carried out with CCDs (Charged Coupled Devices). Compared to photographic plates they have a higher quantum efficiency, a larger dynamical range, repeatable and nearly linear characteristics and a broad spectral sensibility. The incident stream of photons $S(\omega)$ is detected in each pixel in the form of signal electrons $N_{e^-}(x, y)$ and digitized during the the readout (ADU(x, y)).

Fundamentally, a charge coupled device (CCD) is an integrated circuit etched onto a silicon surface forming light sensitive elements called pixels. Photons incident on this surface generate charge via the photoelectric effect that can be read by electronics and turned into a

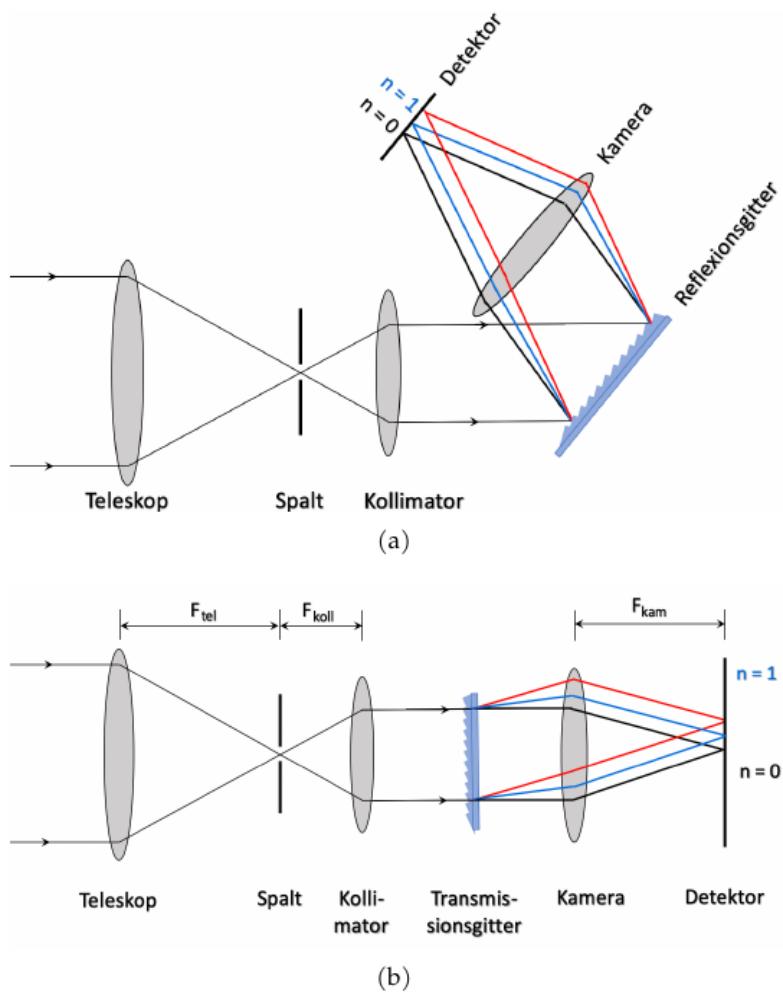


Figure 3: Schematic overview of a slit spectrograph. After entering the instrument through a slit, which here serves as the system's entrance pupil, the light is dispersed at the dispersive element (*top*: reflective grating; *bottom*: transmissive grating). The spectrum is then focused by the camera onto the detector, where the final image is recorded. The choice of the dispersive element does not change the basic set-up. Credit: V. Pfeiffer (2020).

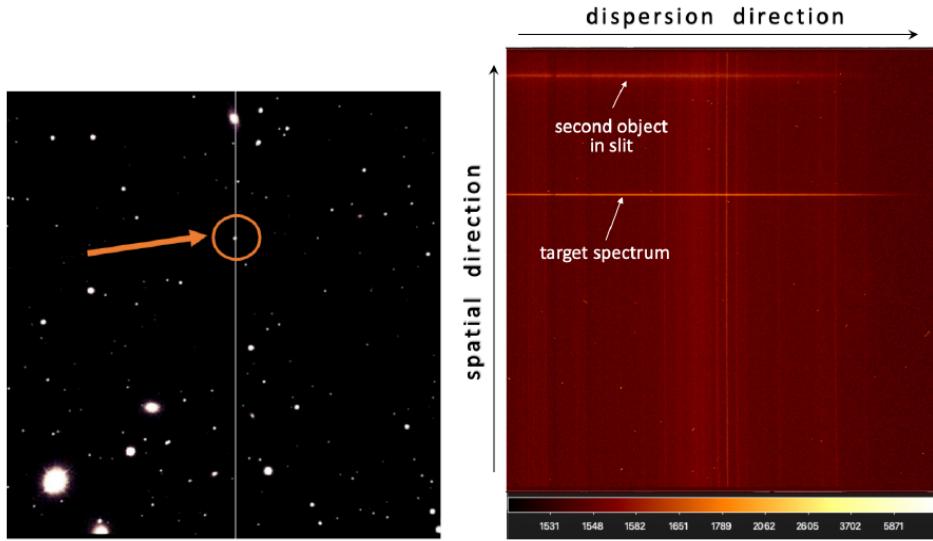


Figure 4: *Left:* Acquisition image showing the position of the slit on sky and objects falling onto it. *Right:* The final image, in which the coordinates along the slit are mapped along the spatial direction, while the spectrum is dispersed along the orthogonal direction. Two distinct targets fall on top of the slit here, so that the spectra of both can be observed. Credit: V. Pfeiffer (2020).

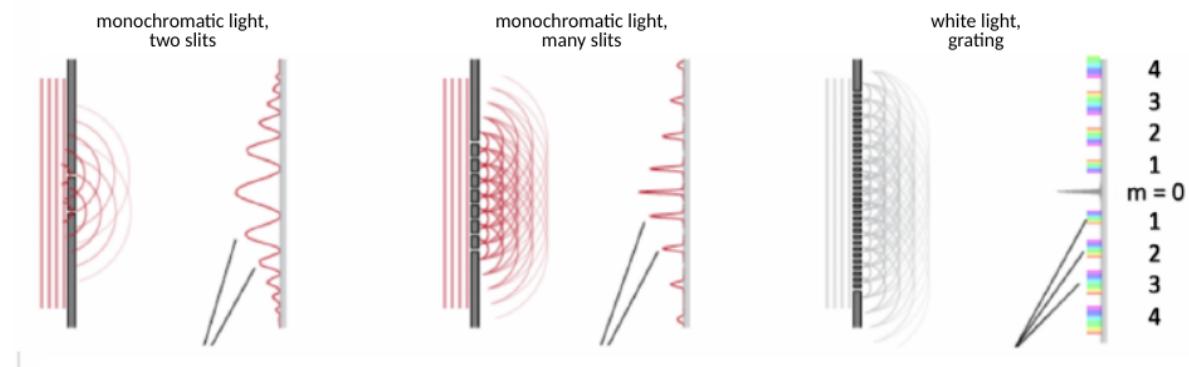


Figure 5: Different diffraction patterns. *Left:* Monochromatic light through two slits. *Middle:* Monochromatic light through many slits. *Right:* Polychromatic light through a grating. With increasing number of illuminated slits the interference peaks become narrower, therefore the resolving power increases. While there is no dispersion at the 0th order, the angular dispersion increases with spectral order. Credit: V. Pfeiffer (2020).

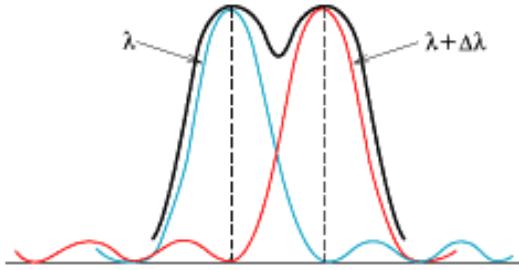


Figure 6: The Rayleigh criterion for the resolving power R of a spectrograph. It is defined by the wavelength difference $\Delta\lambda$, at which two adjacent lines can still be distinguished. Credit: Shimadzu Corporation (2021).

digital copy of the light patterns falling on the device. CCDs come in a wide variety of sizes and types and are used in many applications from cell phone cameras to high-end scientific applications.

The function of a CCD can be visualized as an array of buckets (pixels) collecting rain-water (photons). Each bucket in the array is exposed for the same amount of time to the rain. The buckets fill up with a varying amount of water, and the CCD is then read one bucket at a time. This process is initiated by pouring water into the adjacent empty column. The buckets in this column transfer their “water” down to a final pixel where the electronics of the camera read-out this pixel (the computer measuring the bucket) and turn it into a number that can be understood and stored by a computer.

Of course, this is an oversimplification - in fact, this “model” described above is actually wrong in some ways. All the pixels in a CCD are actually shifted simultaneously, not one column at a time. Fig. 7 shows how that works in principle. In practice the charges generated by the photons on the pixel need to be kept for a certain period of time (the integration time) via a potential well. In order to create a potential well, a voltage is applied to each pixel (called “Vorspannung”). This means that each pixel has a priori a number of electrons in its potential well, regardless of the number of photons received from the science target or sky background and is often called *bias*.

All CCDs benefit from working at lower temperatures. Thermal energy alone is enough to excite extraneous electrons into the image pixels and these cannot be distinguished from the actual image photoelectrons. This process generates an additional component which is called *dark current* and which is dependent on the temperature of the CCD. Modern CCDs are nowadays cooled down to -120°C with liquid nitrogen (LN2), which is a compromise between minimum dark current and maximum quantum efficiency. Lower temperatures would result in lower dark current but also lower quantum efficiency.

Although not discussed in detail, it is clear from Fig. 7 that each pixel of a CCD is a complex element of it. Naturally, the pixels differ slightly in their efficiency to convert the incident photons into electrons, also called *quantum efficiency*. This effect is also wavelength dependent. Furthermore, there are several on-chip sources of noise. The most prominent one is associated with the readout of the detector and the conversion from the analog (electrons) to digital (analog-to-digital Unit = ADU) signal and is called readout noise (RON). This comes from the amplifier(s) involved and adds some noise to the signal. This is a character-

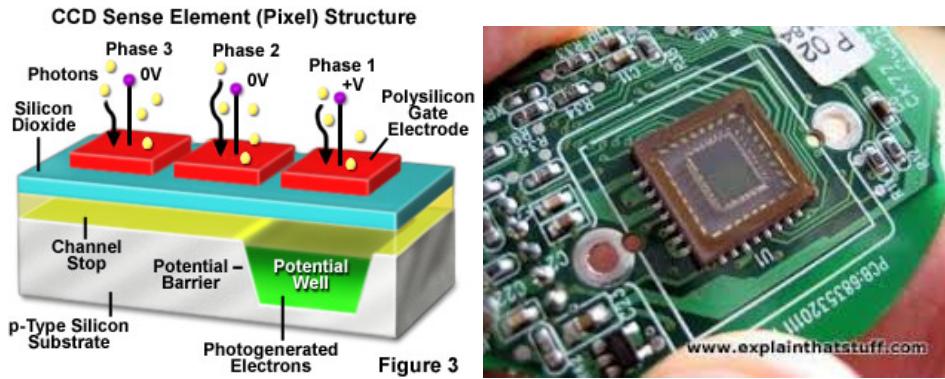
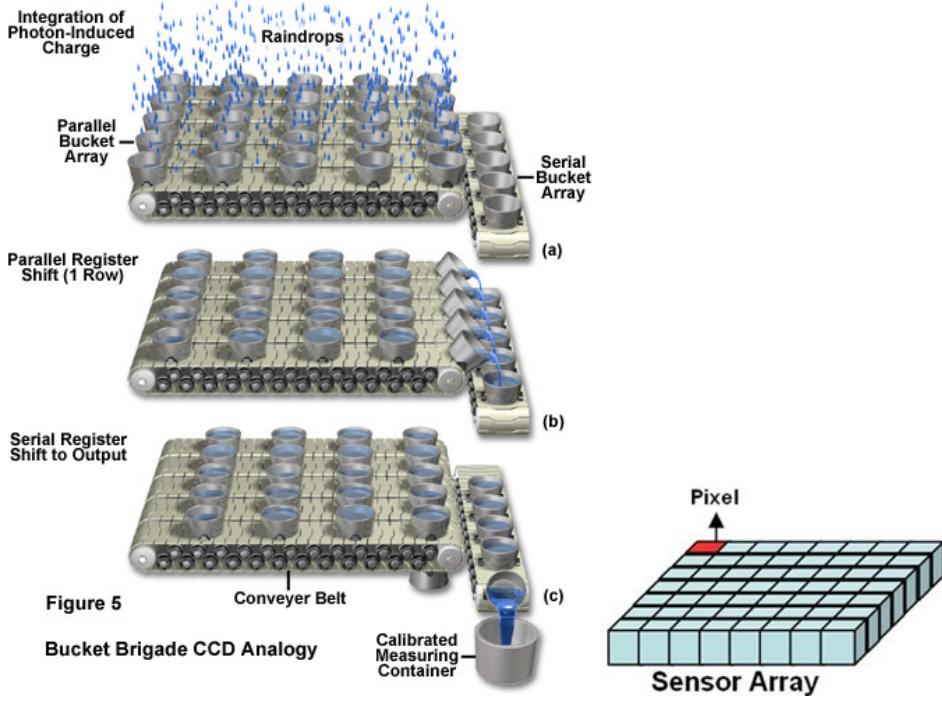


Figure 7: Top left: Working principle of a CCD. Top right: Sketch of an array. Bottom left: Sketch of a pixel. Bottom right: A CCD in reality.

istic of the detector and of the read-out-mode used. The RON is normally given in electrons RMS. If the conversion factor from electrons to ADU called *gain* is known, the readout noise can be converted accordingly. Knowledge of the amplitude of the readout noise is fundamental for some astronomical observations. The signal from the night sky and the target causes shot noise due to the random arrival of photons. Since each photon is an independent event, the arrival of any given photon cannot be precisely predicted; instead the probability of its arrival in a given time period is governed by a Poisson distribution. The statistical properties of a Poisson distribution are well understood, while they are not for the distribution underlying the on-chip sources of noise. Thus the observer must make sure that his observations are shot noise or “photon-noise” dominated.

Those CCD specific sources of error need to be taken into account during post processing of the data. They are typically the first steps in data reduction, and will be covered in Sec. 5.1.

2.4 Outline of the Task

The aim of this course is to familiarize the students with spectroscopy in astronomy. In addition to conveying a basic understanding of the physics behind the observational technique, the process of data reduction will be introduced. Therefore, the students will be provided with all the raw data they need. It comprises of observations of 15 different targets, as well as a full set of calibration frames, in order to correct for instrumental and environmental systematics. The data sets will be further described in Sec. 3. For the data reduction itself the students will use the software package IRAF, developed by the National Optical Astronomy Observatory (NOAO). IRAF provides all the tools needed in order to extract the final clean spectra from the raw data, and to inspect them. While some basics on the software are presented in Sec. 4 to ease the start of the tasks, a rudimentary walkthrough through all the steps of the data reduction follows in Sec. 5. There, some explanations on the background of each reduction step, as well as suggestions regarding parameters that need to be defined, are provided. Still, the students are encouraged to use online help during the course to get a deeper insight on the procedures. Once the target spectra are extracted and calibrated, they should be analyzed from a scientific standpoint. Your assistant will support you and, depending on your time left, suggest proper tasks for you. An overview of what could be done is found in Sec. 6.

For the tasks of this course you can use the CIP-pool computers provided by the Kirchhoff Institut für Physik (KIP)². The systems can be accessed remotely, for which you can use your personal URZ account credentials. Alternatively, you will be provided with a temporary URZ account. On the URZ web page you can chose between any of the five available servers. After the log in, you will be asked to chose a desktop environment. If you choose "LXQT", the desktop will appear. To work with the data, you will need to download the material for this task from the Astrolab webpage and unpack it. Make sure you know, at which location it is stored, so that you can access it. The software you need, basically the FITS file viewer DS9 and the data processing software IRAF (see Sec. 3 and Sec. 4), can be found and started via the symbol in the bottom left corner. Click on it and navigate to the uppermost entry "Bildung". There you should find the two needed programs. If you use one of the other UNIX desktops, then the programs can be found in the upper left corner under "Anwendungen" → "Bildung".

3 Data Sets

The data for this course was taken between September 12th and 18th 2018 during the OPTICON NEON summer school at the *Osservatorio Astrofisico di Asiago*, Northern Italy. For the spectroscopic observations two of the site's cassegrain reflectors have been used, namely the 1.82 m Copernico telescope, and the 1.22 m Galileo telescope (see Fig. 8). Both telescopes are equipped with a low resolution spectrograph, and the observations were performed in a broad wavelength range in the optical between around 350 nm and 800 nm. Altogether, 15 distinct targets, among them stellar sources (stars of different spectral type, a white dwarf), as well as deep-sky objects (different galaxies) were observed. For the purpose of calibra-

²<https://wiki.kip.uni-heidelberg.de/KIPwiki/index.php/EDV:CIP-Pool/RemoteZugang>



(a) 1.82 m Copernico telescope



(b) 1.22 m Galileo telescope

Figure 8: The two reflectors at the Asiago observatory, which were used for the observations.

tion, standard bias and flat frames were taken, and for the wavelength mapping, emission spectra of specific arc lamps were recorded. In addition to that, so called standard calibrator stars have been observed to perform the final flux calibration.

The data files you will work with come in the FITS (Flexible Image Transport System) format. This open format, developed by NASA in the early 1980s, serves as data container. It usually consists of some table data, in this case the raw pixel data of the CCD image, as well as a lot of information about the observation. This information is encapsulated in the so called FITS header and usually contains the names of the used telescope and instrument, the mode and time of the observation, some environmental data, and more. The FITS files can be inspected by any proper FITS viewers. We recommend to use DS9, which is one of them. After you have started the program, a blank window with some panels will show up. You can then open an image by clicking on Datei → Öffnen (top left corner), and navigating through the open-file-dialog to the file in question. After opening the file, you will see the image itself (Fig. 9). You can zoom by using the mouse wheel, and also change the image scale ("skalierung" in the upper panel). A logarithmic scale is usually helpful, as it highlights differences in signal over the whole bandwidth. Furthermore, you can vary the contrast by hovering over the image (up-down, left-right), while holding the right mouse button pressed. Besides that, the file header can be accessed via Datei → Display Header..., which will list all the parameters in a table. Each of those header lines consists of the header *keyword* (name of the parameter), the keyword's *value*, and some *commentary*. Ask your assistant, in case you need any help at this point.

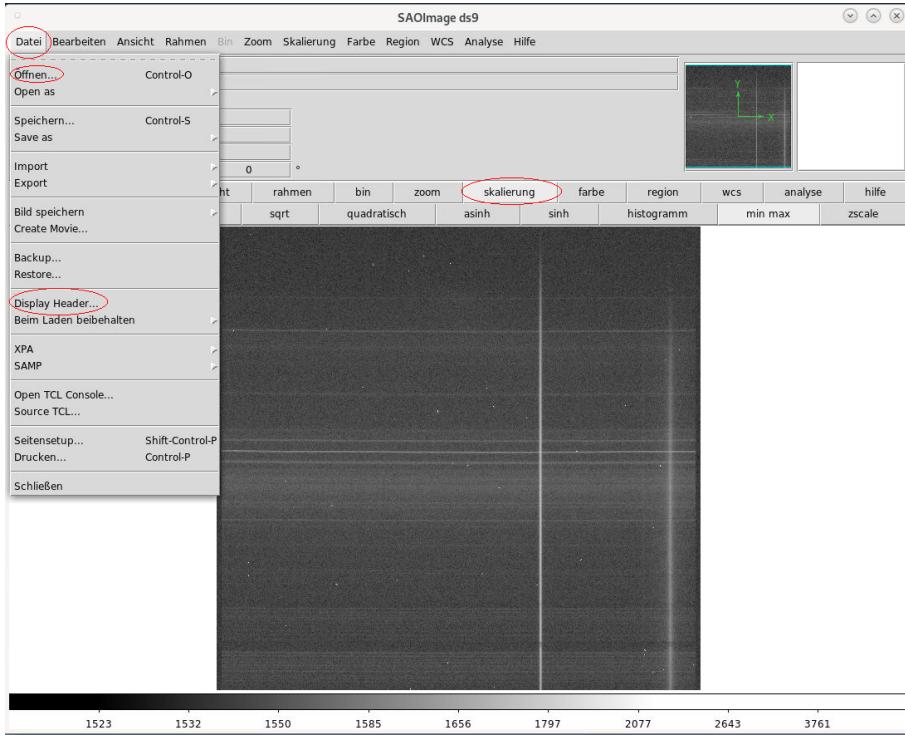


Figure 9: The window panel of the FITS viewer DS9. The important elements are outlined.

Copernico Data

The Copernico telescope with a mirror diameter of 1.82 m can be used for photometry, spectroscopy, as well as for polarimetry. Different grisms can be applied to the Asiago Faint Object Spectrograph and Camera (AFOSC) for the spectroscopic observations. The specifics of the instrumental set-up are listed in Table 1, while the different exposures are summarized in Table 2. For the wavelength calibration two sets of emission spectra are available: three frames of a HgCd arc lamp, exposed at an exposure time of 60 s, and three exposures of a Neon lamp, taken with 30 s. These sets were recorded back-to-back with only a short time in between, so that they can be safely combined for the calibration (see Sec. 5.3 for details).

Galileo Data

In Cassegrain configuration the 1.22 m Galileo telescope is equipped with a low resolution Boller & Chivens spectrograph. Table 3 summarizes some elemental instrumental specifics, and the exposures needed for this course are listed in Table 4. Here, for the wavelength solution the emission spectra of an HeFeAr arc lamp are provided. As during the observation run the lamp exposures were taken subsequent to each science target observations, it is generally recommended for the wavelength calibration to use the specific lamp spectrum, corresponding to the target in question. However, it should be noted that the quality of the lamp spectra varies, so that some might be easier to handle than others. Also, you might be short on time, and as the wavelength calibration is a time consuming step, you might consider to fall back onto one and the same wavelength solution (one single lamp exposure sequence) for all targets.

Table 1: Some specifics on the mode of observations with the Copernico telescope.

AFOSC spectrograph with grism #4	
slit width (on sky)	1.69 as
resolving power R	311
grating constant	300 mm $^{-1}$
central wavelength	580 nm
wavelength coverage	336 nm - 774 nm
dispersion	0.28 nm/pixel
CCD scale	0.25 as/pixel

during observations	
gain	1.91 e $^-$ /ADU
read-out-noise (RON)	7.1 e $^-$
CCD temperature	-95 °C

Table 2: List of observations with the Copernico telescope and raw files provided for the course.

type	exposure time [s]	# of exposures
bias	0	10
flats	8	10
HgCd arc lamp	60	3
Ne arc lamp	30	3
BD+28 4211 (calibrator star)	300	1
object 1	300	2
	600	1
object 2	600	3
object 3	600	3
object 4	300	1
object 5	600	3
object 6	600	4
object 7	600	3
object 8	600	3

Table 3: Some specifics on the mode of observations with the Galileo telescope.

Boller & Chivens spectrograph	
grating constant	300 mm ⁻¹
wavelength coverage	300 nm - 800 nm
CCD scale	1.25 as/pixel
during observations	
gain	0.97,e ⁻ /ADU
read-out-noise (RON)	3.4 e ⁻
CCD temperature	-80 °C

Table 4: List of observations with the Galileo telescope and raw files provided for the course.

type	exposure time [s]	slit width [microns]	# of exposures
bias	0	-	10
twilight flats	varying	176	8
flats 5''	2.5	294	11
flats 3''	5	176	11
flats 1''	20	118	30
HeFeAr arc lamp	15	176	6
	15	294	4
HR 1544 (calibrator star)	30.0	294	1
	5.0	294	1
object 9	600	176	3
	600	118	1
	100	118	1
object 10	600	176	3
object 11	600	176	2
object 12	300	176	3
object 13	600	176	4
object 14	600	176	3
object 15	600	294	1

4 Introductory Remarks on IRAF

Data reduction is in general a non trivial task in astronomy. As observational data typically is processed by CCDs, the raw images need to be corrected for any effects that are specific to those detectors. While the same applies to photometric observations, in spectroscopy another key step is necessary. Initially, all the spectral information is distributed across a two-dimensional array, and therefore it needs to be extracted properly, in order to arrive at a one-dimensional flux distribution across the wavelength axis. IRAF provides all the needed tools to get there, and is flexible enough, so that it can handle data from different instruments. This section serves as a manual to guide you through all the different steps. If you need any further help, you can talk to your course's assistant, but you can also try to find the solution to your problem online. You will find a lot of helpful information on the web, still we recommend the following sources:

- <https://iraf.net/>³
- <http://research.iac.es/sieinvens//SINFIN/CursoIraf/Intro.php>
- https://www.stsci.edu/instruments/wfpc2/Wfpc2_dhb/appendixA3.html
- A Beginner's Guide to Using IRAF (J. Barnes)
- A User's Guide to Reducing Slit Spectra with IRAF (P. Massey, F. Valdes, J. Barnes)
- Reducing Spectra with IRAF (H. Lietzen, S.-M. Niemi)

Once you have found and started IRAF on your computer, a console will open, which will be your interface to the program (see Fig. 10). Before you start the data reduction, some initial steps need to be performed. Some of them are essential, while others will simplify your further work:

- execute: `set imtype = "FITS"`
- use `cd <path>` to move to working directory (where the data is located)
- define the dispersion direction by typing: `set dispaxis = 1` for the Galileo data set, and `set dispaxis = 2` for the Copernico set respectively

All procedures in IRAF, also called *tasks* in the software's terminology, need to be loaded first before they can be called. The tasks are distributed over many different packages. IRAF's main packages are outlined in Fig. 10. Calling one of them loads the package and lists the next level of subpackages, where again you can chose the next, and so on. Once the last level is reached, all the tasks in the last loaded subpackage are loaded and will be ready to be used. In the following, when a task is introduced in this document, it will be indicated, in which subpackage it is located. For example, the task `imcombine` can be found in the subpackage `immatch`, to which you can navigate by: `images -> immatch`. The

³Especially useful as a reference to the individual tasks and parameters, which need to be defined.

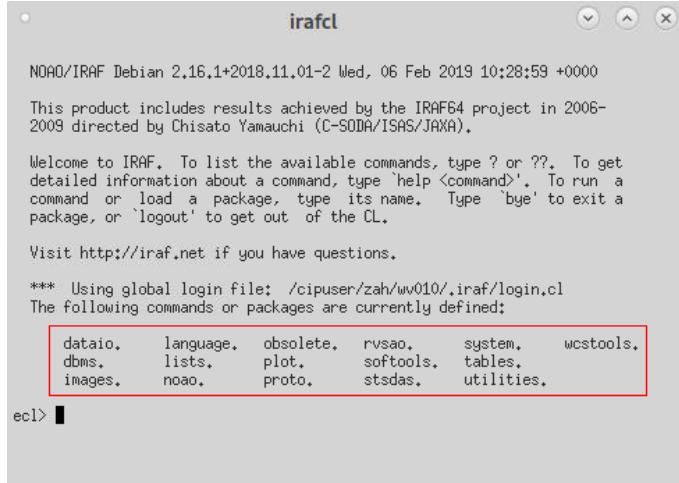


Figure 10: The standard IRAF console. The available main packages are outlined.

command `bye` is used to move one level back up in the packages hierarchy. The most fail-proof way to execute an individual task is to call the command `epar`, followed by the task name (e.g.: `epar imcombine`). You will then see the full list of parameters, which you can modify, simply by moving to the line in question and typing in the desired parameter's value. Once finished, you can now execute the task by typing `:go`. You can also escape the parameter list without execution by `Ctrl-d`. Alternatively, tasks can also be executed directly by typing the task's name, followed the mandatory parameters, without entering the parameter list.

5 Data reduction

In the following all tasks that need to be performed in IRAF to arrive at fully reduced one dimensional spectra will be described. The most important parameters that you need to set/modify to some recommended values will be listed and shortly explained. However, you are encouraged to use any online references to find other options and to try them out. Default values can be used for parameters not listed explicitly. For the sake of clarity, we have subdivided the tasks into Calibration, Spectrum Extraction, Wavelength Solution, and Scientific Corrections. In Fig. 11 the whole process and data flow are summarized within a floatchart.

5.1 Calibration

In Sec. 2.3 some properties of CCDs were introduced. Applying the tasks laid out in the following, the raw images will be corrected for bias, trimmed to useful areas on the detector, and normalized to the flat field lamp exposures taken in context of the observations.

Bias Correction

In the course of the observations bias images were taken. Those are typically 0 s exposures with the shutters closed, which can be used to determine the bias level. As the signal in each

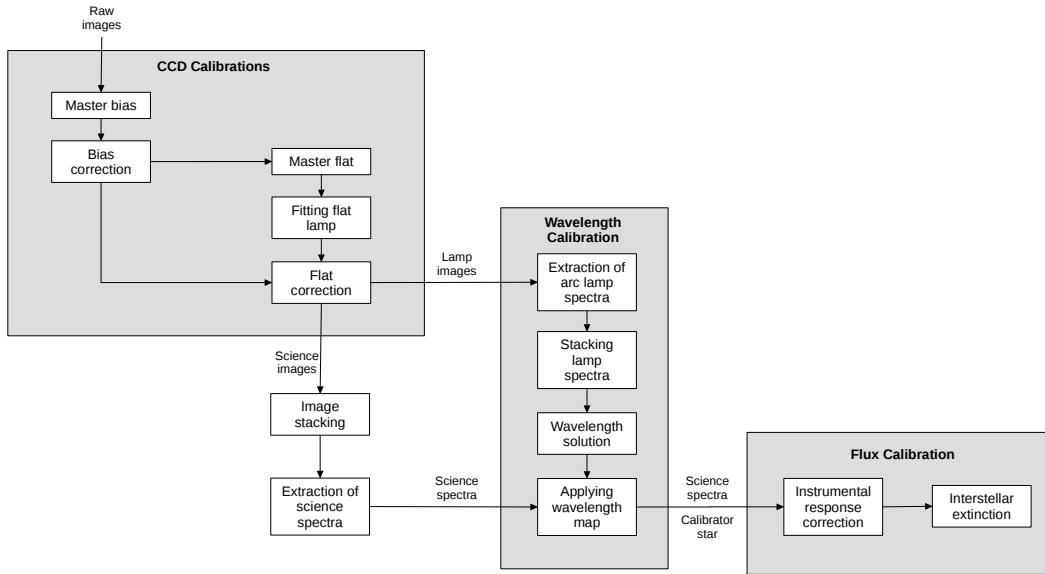


Figure 11: The data reduction process as a whole. All raw images need to be corrected for CCD related effects. A wavelength solution needs to be found to map pixel coordinates onto a wavelength grid, and eventually a flux calibration is performed and applied onto the science spectra.

pixel is subject to stochastic variations, robustness is achieved by combining the different bias images first by some kind of averaging.

i. Combination of the bias images:

The raw bias exposures will be combined to one image, the so called *master bias*.

TASK	:	zerocombine	
PACKAGE	:	noao -> imred -> ccdred	
input	=	@bias.txt	Text file with list of filenames of the input files.
output	=	M_bias	Name of the created master bias file.
ccdtype	=		CCD image type to combine. Leave this blank.
rdnoise	=	rdnoise	CCD read-out noise; value from FITS file header will be used.
gain	=	gain	CCD gain; value from FITS file header will be used.

Whenever more than one file is required as input, you can chose between two ways of defining them. In the parameter edit view you can enter all the files, separated by a comma (note that if those are FITS file, you do not need to provide the file ending ".fits"). The second option, as applied in the list above, is to provide IRAF with a text file, in which all the filenames are listed line by line. In the latter case the name of the text file needs to be prefixed by "@".

ii. Applying the bias correction & Trimming:

The master bias is used to correct all images. Also, the original images can be trimmed here. This is useful in order to get rid of the under- and/or overscan areas on the detector. Those are virtual pixel rows/columns at the boarders of the image that alternatively can be used to estimate the bias level. If they are not trimmed, they can become problematic in the later process, in particular with tasks that expect some smooth signal variations across the image. Check the science images before trimming to find the regions to be discarded.

TASK	:	ccdproc	
PACKAGE	:	noao -> imred -> ccdred	
input	=	@all.txt	Text file with list of all input files to be corrected.
output	=	@all_bias.txt	Text file with output filenames. Needs to be of the same length as the input list.
ccdtype	=		CCD image type to correct. Leave this blank.
fixpix	=	no	Fixing bad CCD lines or columns.
oversca	=	no	Applying overscan strip correction.
trim	=	yes	Image trimming.
zerocor	=	yes	Applying zero level (bias) correction.
darkcor	=	no	Applying dark current correction.
flatcor	=	no	Applying flat field correction.
trimsec	=	[26:2065,7:2044]	Trim section for Copernico data set images.
	=	[7:2047,37:512]	Trim section for Galileo data set images.
zero	=	M_bias	The master bias calibration frame.

Note that you can leave the value of the parameter "output" empty. In this case the corrections will be applied to the input images, and the original data will be overwritten. In the list above, the input, as well as the output, are provided within text files (prefixed by "@"). In this case the output file needs to be of the same format as the input, since the input file names will be mapped line by line to the product file names. The easiest way to achieve this, is to copy and to rename the input text file, and also to modify the individual entries by "search & replace" within a standard text editor.

Flat Normalization

While in photometry flat-fielding is performed to correct for pixel-to-pixel variations and potential vignetting (not uniform illumination of the chip) on the two dimensional image, in spectroscopy for the flat calibration a continuous reference lamp spectrum along the dispersion direction is utilized. For the observations of this course dome flat exposures were taken for this reason. As with the bias frames, they will be combined first. The combined master flat will then be used to fit a smooth response function along the dispersion axis. This is done, in order to get rid of the lamps' own specific spectral distribution. This way only the higher frequency variations are then corrected for during the last step.

i. Combination of the flat images:

The bias corrected (and trimmed) flat field exposures will be combined to one image, the so called *master flat*.

TASK	:	flatcombine	
PACKAGE	:	noao -> imred -> ccdred	
input	=	@flat_biascor.txt	Text file with list of bias corrected flats to combine.
output	=	M_flat	Name of the created master flat file.
reject	=	crreject	Type of rejection during averaging.
ccdtype	=		CCD image type to combine. Leave this blank.
rdnoise	=	rdnoise	CCD read-out noise; value from FITS file header will be used.
gain	=	gain	CCD gain; value from FITS file header will be used.

ii. Fitting the flat lamp spectrum:

In this step the created master flat will be used to determine and to isolate the pixel-to-pixel variations to be corrected for. For this the flat is collapsed across the spatial axis, so that a one dimensional regression can be performed, in order to fit the flat lamp's spectral distribution. This way only the high frequency variations are left over. This is done in an interactive mode, during which the order of the fitting function may be adjusted, until the result is satisfactory.

TASK	:	response	
PACKAGE	:	noao -> imred -> kpnoslit	
calibrat	=	M_flat	Calibrator image for determination of the response (the created master flat).
normaliz	=	M_flat	Image for determination of the normalization (the created master flat).
response	=	flat_response	Name of the output response file.
interac	=	yes	Interactive mode.
functio	=	spline3	The fitting function (cubic spline by default).
order	=	3	Order of the fitting function. In interactive mode this can be adjusted.

When you have chosen to do the fit interactively, then you will be asked to verify that. Type "yes" here. In the following, you will see a plot, similar to Fig. 12, with the collapsed flux on the y-axis, and a pixel dimension (position) on the x-axis. A dashed overplot depicts your current fit, which you want to nicely resemble the flat lamp spectrum. In order to modify the fitting order, keep the plotting window active, type :o 5 (to change it for example to 5), and confirm with ENTER. You will see the

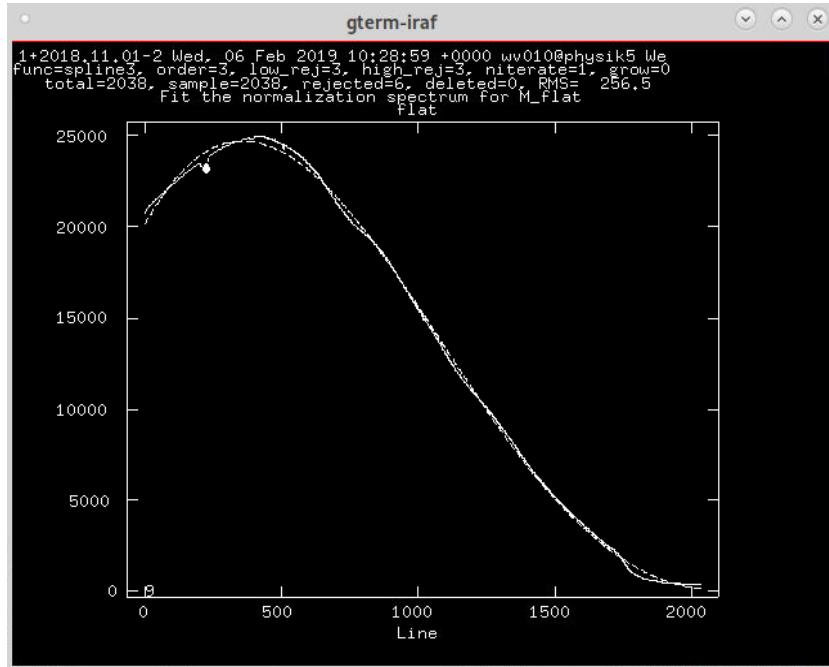


Figure 12: The collapsed flat lamp spectrum together with the current fit (dashed line).

typed commands in the IRAF console. After changing the fitting order, you can redo the fit by pressing "f". Outliers, single data points with significant deviations, can be eliminated by hovering the cursor on top and pressing "d" for delete. Once you are satisfied with the result, press "q" to quit the task. The fitting is sufficient, when the lower frequency variations are well resembled by the fitting curve.

iii. Applying the flat correction:

The flat normalized variation map is now used to correct the science data for the unwanted higher order variations. All features apparent in the flat field exposures, on top of the fitted lamp spectrum, will be corrected for.

```

TASK      : ccdproc
PACKAGE   : noao -> imred -> ccdred

input     =      @all_bias.txt  Text file with list of all bias corrected input files.
output    =      @all_flatcor.txt Text file with output filenames. Needs to be of
                                the same length as the input list.
ccdtype   =
fixpix    =
oversca   =
trim      =
zerocor   =
darkcor   =
flatcor   =
flat      =      flat_response The response file to be used (created in the previous task).

```

5.2 Spectrum Extraction

In the following step the one dimensional spectra will be extracted from the two dimensional images. As most of the sources were observed more than once, the individual images will be combined (stacked) in the first step.

i. Image stacking:

The flat normalized target exposures will be combined to one image. Weighting will be applied according to the exposure times inside of the file headers.

TASK	:	imcombine	
PACKAGE	:	immatch	
input	=	@objx.flatcor.txt	Text file with list of flat normalized target exposures.
output	=	objx	Name of the output image.
combine	=	average	Type of combine operation.
reject	=	crreject	Type of rejection during averaging.
weight	=	exposure	Weighting scheme.
rdnoise	=	0	CCD read-out noise.
gain	=	1	CCD gain.

ii. Spectrum extraction:

In the raw images the spectra follow a distinct trace along the dispersion direction. In this task this trace is searched for in the target image. Once its position in the spatial direction is found in each wavelength bin, the spectrum of the target can be extracted within a given width, centered on the trace position. This is done in two steps. During the first step the pixel intensities along the dispersion axis are summed up to find the rough position of the spectrum (Fig. 13a). A so called *aperture* is then placed on top of the spectrum, along which the spectrum is traced in the second step. This trace is again fitted by a smooth function (Fig. 13b) that is then used for the final extraction. During the extraction the background is also modeled and subtracted from the flux. Therefore, the background level is sampled in a specific region around the aperture, defined by the parameter "b_sampl". **Modify those values, if the background is overestimated, which could happen, when the regions are chosen too close to the actual target.**

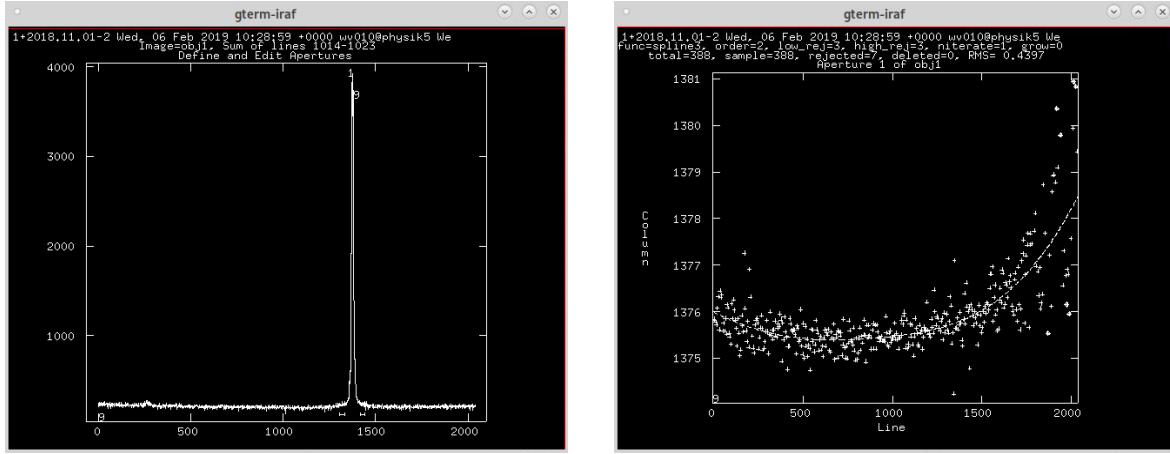
TASK	:	apall	
PACKAGE	:	noao -> imred -> kpnslit	
input	=	objx	The flat corrected and stacked target image.
output	=	objx_spec	Name of the output file (the extracted spectrum).
recentre	=	yes	Recentring apertures.
resize	=	no	Resizing apertures.
edit	=	yes	Editing apertures.
trace	=	yes	Tracing apertures.
b_funct	=	chebyshev	Background function.

b_order	=	2	Background function order.
b_sampl	=	-70:-40,40:70	Background sample regions.
b_naver	=	-15	Background average mode and pixel number.
b_niter	=	2	Number of background rejection iterations.
width	=	5	Profile centering width.
radius	=	5	Profile centering radius.
nfind	=	1	Number of apertures to be found automatically.
bkg	=	yes	Background subtraction.
t_nsum	=	5	Number of dispersion lines to be summed at each step along the dispersion.
t_step	=	5	Step along the dispersion axis between determination of the spectrum positions.
t_funct	=	spline3	Trace fitting function.
t_niter	=	1	Trace rejection iterations.
backgro	=	fit	Background to subtract.
weights	=	variance	Extraction weights.
clean	=	yes	Detect and replace bad pixels

When the routine is started, you will be prompted to confirm the search of apertures, the number of apertures, as well as the opportunity to edit the apertures. Confirm all by pressing `ENTER`. In the following plot (Fig. 13a), you will see the selected and numbered apertures. Proceed by pressing "`q`" and confirm the tracing of the aperture until the trace plot pops up (Fig. 13b). As before with interactive fitting, you can modify the fitting order iteratively. Note that the scatter might be increased at one side of the trace. This is an expected feature due to a low flux level. Do not care too much about the fit in those regions. If desired, you can as usual delete outliers with "`d`". When satisfied with the fit, that is, when the regression curve follows the trace reasonably well, proceed with "`q`", and confirm the following questions. Eventually, the extracted spectrum will be plotted. Finish the task by pressing "`q`" again.

From now on you can plot and inspect the extracted spectra with the tasks `splot` or `specplot`, both to be found in `noao -> imred -> kpnoslit`. Just type the name of the task, followed by the name of the spectrum to be plotted. In `splot` you can zoom in by "`z`" and move left/right using ";" and/or ".". Redraw the graph in full scale again with "`c`". Check the online reference for `splot` to see all the commands.

You are provided with spectra of different sources. For point like sources, such as stars, the choice of aperture position is trivial. However, this is not necessarily the case with extended sources. They cover some extended regions on the sky, and the placement of the slit acts like a cross-section. In those cases it happens that you see several spectra of different strength along the spatial axis (x-axis in Fig. 13a), as different regions of the object along the slit are probed. Here, depending on, what you are interested in, it might be useful to extract more than one spectrum from the image. Apart from the spectral continuum, in some of those objects you might also notice an emission



(a) Flux summed up along the dispersion direction to identify the position of the spectrum (aperture). The identified aperture, as well as the background regions are marked.

(b) Fitting the trace of the spectrum: The x-axis resembles the dispersion direction, and the y-axis the position of the spectrum in the spatial direction. The dashed curve illustrates the current fit.

Figure 13: Plots during the extraction with the task `apall`

spectrum (distinct lines) in the images, which arises from emission regions within the galaxies in question.

5.3 Wavelength Solution

While the target spectra are already extracted, if you plot them, you will see that the flux is plotted against pixels along the dispersion direction. This is the case, as there is no wavelength information available in the data. In the process of the following wavelength calibration, a mapping from pixel to wavelength space is created, which is often also called *wavelength solution*. To do this, one generally needs some spectra with lines at well known positions. For this reason, emission line spectra of arc lamps have been taken during the observations that in the following will be used.

Processing emission lamp spectra

In the first step the emission line spectra need to be extracted in the same format as the target spectra are present. As several exposures per lamp type, as well as different lamp types were used, those need to be combined afterwards. During the final step, the combined spectra need to be inspected, and the emission lines identified. Providing the lines' wavelengths at rest frame, the wavelength solution can be established.

i. Extraction of lamp spectra:

The lamp spectra are extracted by the same task as the target object spectra. However, some particularities need to be accounted for during the extraction. While the target spectra typically exhibit some spectral continuum along the dispersion direction, this is not the case with the spectra of the arc lamps. Their images rather show emission at distinct positions, which are continuous in the spatial direction. Therefore, the trace of the spectrum can not be traced, as it was done with the sky targets. Due

to that, for the extraction of lamp spectra a science target serves as a reference for the aperture and trace of the spectrum (see parameter list below). Of course, the reference needs to have been extracted already, during which the trace information was stored in a database within your working directory, and is available to IRAF. That way the lamp spectrum will be extracted following the same trace on the detector as the science target. Consequently, in principle such an extraction should be done for each target separately. However, as this might be quite extensive, and we do not expect the wavelength solution to differ too much along the slit, you should be fine with a single lamp extraction (using only one particular reference) that in the following you can apply to all targets. Furthermore, it should be made sure that no background is modeled and subtracted in the course of the extraction, as this would eliminate the emission lines signals.

TASK	:	apall
PACKAGE	:	noao -> imred -> kpnoslit
input	=	lampimage_flatcor A flat corrected arc lamp image.
output	=	lamp_spec Name of the output file (the extracted spectrum).
referen	=	objx Reference for the extraction: the name of the flat corrected and stacked target image that was used before already for target extraction.
recentre	=	no Recenturing apertures.
resize	=	no Resizing apertures.
edit	=	no Editing apertures.
trace	=	no Tracing apertures.
bkg	=	no Background subtraction.
backgro	=	none Background to subtract.

ii. Combining lamp spectra:

Since there are several arc lamp spectra available and they might differ in quality, it might be useful to combine them before the line identification. This step is not mandatory, if you make sure that the lines in your extracted spectrum show enough flux for further analysis. Still, for the 1.82 m Copernico data set two different arc lamps were used. As the Ne lamp exhibits emission lamps mostly at one side of the spectrum, while the HgCd lamp on the other, in this case it is in fact necessary to combine them to one final spectrum. This is achieved with the task `scombine`.

TASK	:	scombine
PACKAGE	:	noao -> imred -> specred
input	=	@lamp_specs.txt Text file with list of extracted lamp spectra to be combined.
output	=	lamps_combined Name of the combined output spectrum.
scale	=	none Image scaling.
zero	=	none Image zero point offset.

```
weight      =          none  Image weights.
```

iii. Line identification & wavelength fitting:

After the arc lamps spectra have been extracted and combined, they are ready to be inspected. In the following, within the task `identify` the spectrum needs to be compared to a given reference, and the emission lines to be identified. You can use Fig. 14 as a line reference for the Ne and the HgCd lamps of the Copernico data set, and Fig. 15 for the HeFeAr lamp of the Galileo data set respectively.

```
TASK      : identify
PACKAGE   : noao -> imred -> kpnoslit
images    =  lamps_combined  Name of the combined lamp spectrum.
coordli   =          User line list. Leave this blank.
fwidth    =      7.5  Feature width in pixels.
cradius   =      5.  Centering radius in pixels.
thresho   =     10. Feature threshold for centering.
functio   =  spline3  Wavelength fitting function.
order     =      3   Fitting function order.
```

Once you execute the task, the lamp spectrum will be plotted. Use the given references for line identification. In order to enter a particular line's wavelength, move the crosshair to center it on the line. Keep it centered and press "**m**" to mark the feature. If the line is recognized, its pixel location will be estimated, and you will be prompted for the wavelength. Enter it in units of Ångstrom and press ENTER. If for some reasons the line is not recognized, you can retry by moving the crosshair a bit. If the plot is too small for proper identification, pull one of the corners of the window to expand it and press "**r**" to redraw the spectrum. In order to arrive at a satisfactory wavelength solution, you should identify and mark several lines, well distributed across the whole spectrum. From experience a number between five and ten should work very well, although even three or four correct ones might suffice. It is however important to make sure that lines near the boarders of the spectral range are identified, as otherwise extrapolation of your solution might quickly lead to significant deviations. After marking several lines, you can start the fitting by pressing "**f**". A residual plot to your fit will be shown. Here you can as usual modify the fitting order and identify outliers that you might want to delete. You can do this with "**d**", which will erase the data point nearest to the cursor. After that, update the fit by "**f**". Pressing "**q**" brings you back to the line identification screen. Now your current wavelength mapping should have been applied, so that the x-axis is given in wavelengths. You can further identify and mark lines, after which you can redo the fit, or you can finish and proceed again with "**q**". In that case confirm to write the result into the data base.

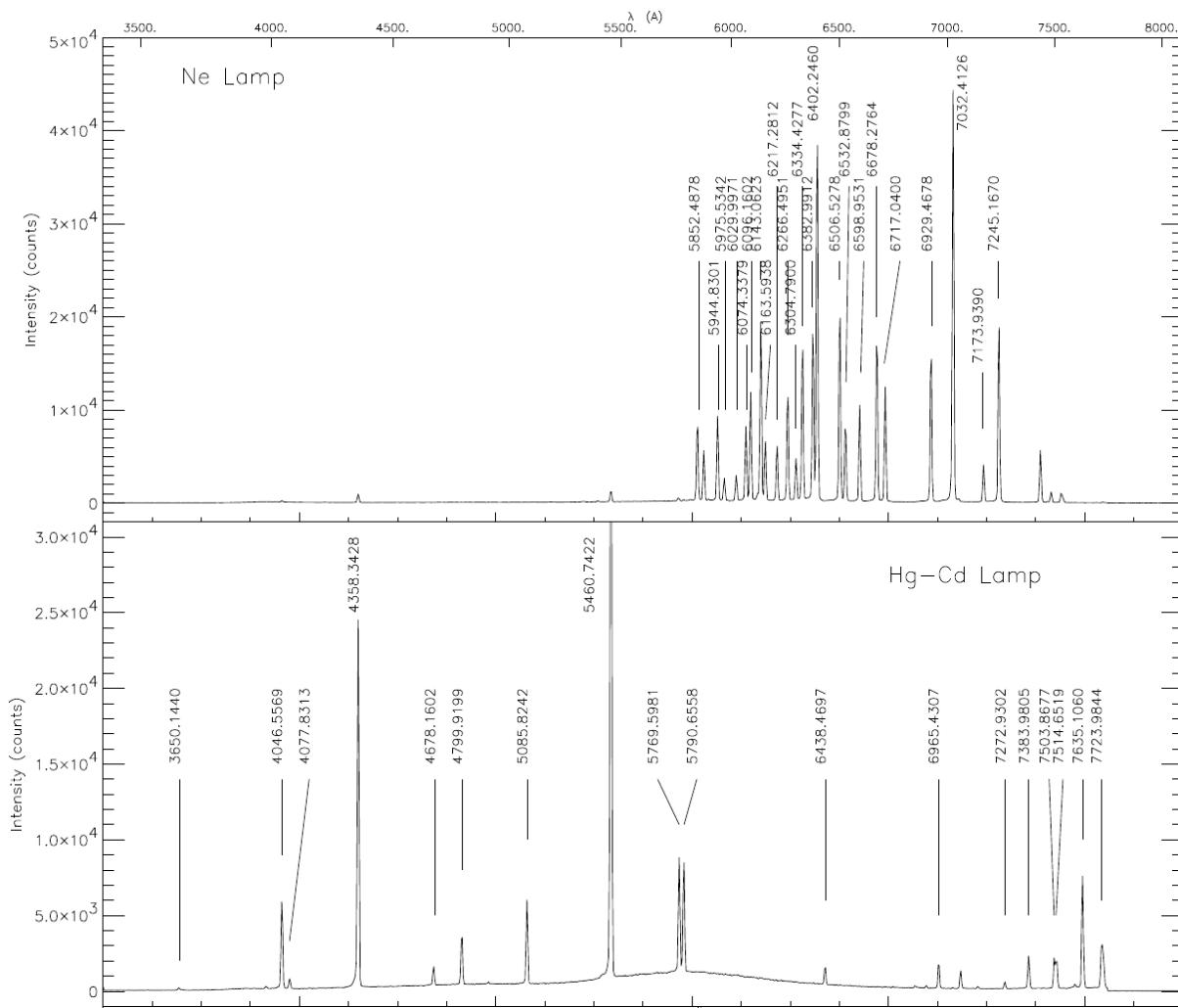


Figure 14: Wavelength reference for the Ne and HgCd emission lines to be used for the Copernico data set.

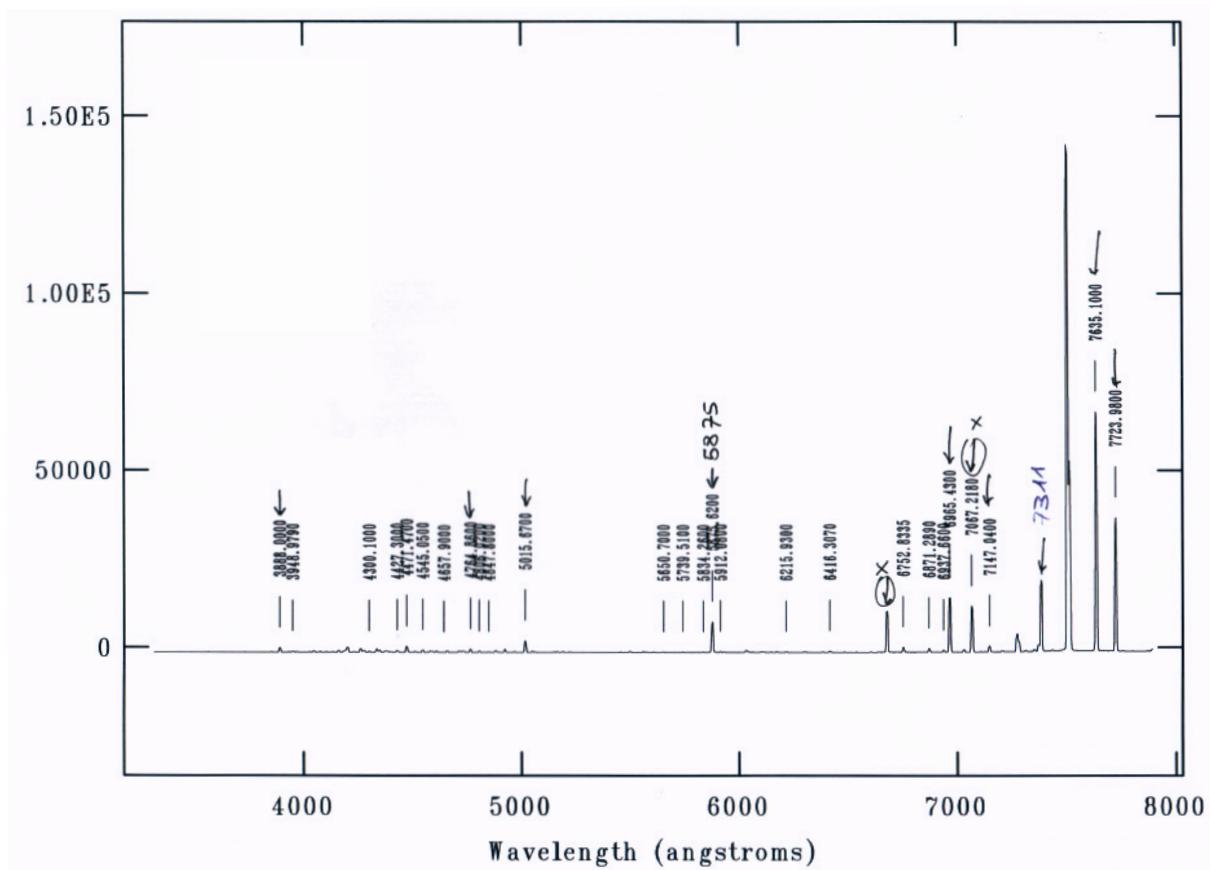


Figure 15: Wavelength reference for the HeFeAr emission lines to be used for the Galileo data set.

Wavelength mapping in science spectra

After finishing the last step, you are now equipped with a proper wavelength solution, which now needs to be applied to the extracted science target spectra. This is achieved by the following two steps.

i. Editing file header:

The file header of the target spectrum needs to be edited. In this step a new header keyword is created, in which the name of the wavelength reference is stored. This is needed for the step afterwards, in which the actual wavelength solution is applied.

TASK	:	hedit	
PACKAGE	:	imutil	
images	=	objx_spec	Extracted target spectrum.
fields	=	REFSPEC1	Header keyword to be edited.
value	=	lamps_combined	Name of the lamp spectrum used for line identification.
add	=	yes	Add keyword rather than editing.

ii. Applying the wavelength solution:

With the provided information in the file header, in this step the wavelength solution is finally applied.

TASK	:	dispcor	
PACKAGE	:	noao -> imred -> kpnoslit	
input	=	objx_spec	Extracted target spectrum.
output	=	objx_speccor	Name of the output spectrum.

You can check, if everything went well by plotting the spectra as usual with `splot` or `specplot`. The flux should now be plotted against wavelength.

5.4 Flux Calibrations

At this stage you can already work with the extracted spectra. Depending on what aspect you are interested in, the following steps are not needed. However, the flux distribution in the spectra is currently still affected by the instrument, as well as by further effects, such as the interstellar extinction. In the following, we want to correct for both.

Instrumental Response

The whole instrument, the spectrograph together with the telescope, exhibit a specific finger-print in terms of flux feed-through against wavelength. Usually, as in our case, observations of a flux calibrator star are being used to correct for this instrumental response. Such a flux

correction is possible, as the spectrum of the standard star is known from previous observations. One compares the recorded spectrum of the calibrator star to its known reference (basically one divides one by the other), which yields the instrument's sensitivity function. That one is consequently used to correct the spectra of all science targets. This is accomplished in three steps, listed in the following. Note that you will need to have processed the image/spectrum of the standard calibrator star the same way, as you have done for the science targets. The spectrum needs to be extracted, and the wavelength solution applied.

i. Setting up flux calibration:

In this step the flux calibration is prepared by providing the information on the used calibrator star. IRAF will recognize the star ID and will find it in its data base. The airmass of the observation (height above horizon) needs to be provided, as well as the exposure time. In our case this information is given in the file headers of the spectra.

TASK	:	standard
PACKAGE	:	noao -> imred -> kpnoslit
images	=	calibstar_spec Spectrum of the calibrator star.
output	=	std Output flux file used in the following (you can leave the default here).
caldir	=	onedstds\$spec50cal/ Directory with calibration data (Copernico data set).
	=	onedstds\$spec16cal/ Directory with calibration data (Galileo data set).
star_nam	=	xxx Star ID in IRAF's calibration list (bd284211 for Copernico data set; hr1544 for Galileo data set).
airmass	=	Airmass of the observation. If blank the value from the file header will be used.
exptime	=	Exposure time of the observation. If blank the value from the file header will be used.

When using the calibrator star bd284211 together with the Copernico data, you can leave "caldir" unmodified at its default value. The calibrator star still should be found.

ii. Fitting the sensitivity function:

Provided with all the needed information, in this step the sensitivity function will be fitted.

TASK	:	sensfunc
PACKAGE	:	noao -> imred -> kpnoslit
standard	=	std Input standard star data (output from previous task).
sensitiv	=	sens Output sensitivity function image name.
functio	=	spline3 Fitting function.
order	=	3 Fitting function order.

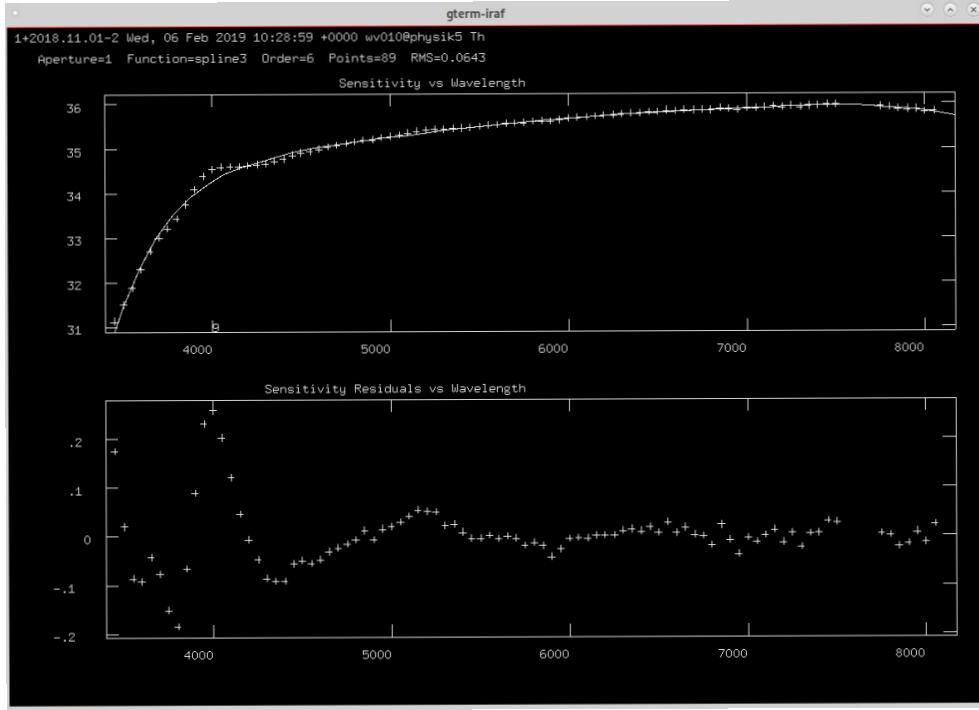


Figure 16: Fitting the sensitivity function with the task `sensfunc`. *Top*: Data and the current fit. *Bottom*: The residuals to the fit.

When the task is executed, two plots as in Fig. 16 will appear. The upper plot shows the data to which the fit of the sensitivity function is performed, as well as the fit itself. As before, you can modify the fitting order and redo the fitting with ”**f**”, delete outliers with ”**d**”, and finish with ”**q**”. The lower plot shows the residuals to the fit. The aim is to have the residuals normally distributed around zero along the y-axis.

iii. Applying the flux correction:

In this step the fitted sensitivity function will be used to correct the flux within the science targets’ spectra. To correct for atmospheric extinction, the airmass of the observation, provided by the header, is also needed.

TASK	:	calibrate
PACKAGE	:	noao -> imred -> kpnoslit
input	=	objx_speccor Input science spectrum to calibrate.
output	=	objx_fluxcor Output spectrum file name.
extinct	=	yes Applying atmospheric extinction correction.
flux	=	yes Applying flux correction.
airmass	=	Airmass of the observation. If blank the value from the file header will be used.
exptime	=	Exposure time of the observation. If blank the value from the file header will be used.

Correction for interstellar extinction

The interstellar space within the Milky Way is not empty. Consequently, the light is affected by it. In particular, scattering due to the interstellar medium needs to be taken into account when investigating spectral distributions. As the effect is wavelength dependent, light of shorter wavelengths is affected stronger, and the observed target appears redder than it truly is. The dust and matter distributions within our galaxy are heterogeneous, and therefore not isotropic with reference to the Earth's location. Consequently, the amount of interstellar, or galactic, extinction varies with the direction of the observation. In our last step we will try to correct for that. Usually, the degree of reddening is described by E_{B-V} , the object's *color excess*, which is determined by

$$E_{B-V} = (B - V) - (B - V)_0, \quad (4)$$

where $(B - V)$ is the observed colour index, and $(B - V)_0$ the true value, specific to the source. As with more matter along the light path, both, the reddening, as well as the extinction in general increase, the two effects can be linked by $R = A(V)/E_{B-V}$, the *relative visibility*, where $A(V)$ is the extinction measured in the V band. Together with the extinction values given in Table 5, the reddening is corrected for with the task `deredden`. After the correction is applied, compare the two products (before and after the dereddening).

TASK	:	deredden
PACKAGE	:	noao -> imred -> kpnoslit
input	=	objx_fluxcor Input spectrum to be corrected.
output	=	objx_dered Name of output spectrum.
value	=	xx Extinction value from Table 5.
R	=	3.1 Relative visibility $A(V)/E_{B-V}$.
type	=	A(V) Type of extinction parameter.

6 Analysis & Interpretation of the Results

Once the spectra are extracted and calibrated, they can be used for target identification and analysis. When inspecting the plots of the spectra, you should notice that the signals become weak and noise dominated below and above some specific wavelengths. Typically, the spectra are not useful below around 3,600 Å or above 7,700 Å. Also, in some circumstances you might not be satisfied with the flux correction, in particular with the result of the de-reddening. In those cases you still can do your analysis with either the corrected or the uncorrected spectrum, although the overall flux distribution may not be reliable.

Depending on the data set and science objects you reduced, you will be left with very different spectra. In the course of the data underlying workshop stellar sources, as well as extra galactic targets have been observed. If you are already experienced with the look and features of astronomical spectra, you can try to identify the types of the objects observed. Keep in mind that the spectra in general are combinations of some continuum, overimposed by absorption and emission features, related to the chemical composition of the material.

Table 5: Extinction parameters $A(V)$ in the V band (V. Pfeifer 2020; based on the NASA/IPAC Extinction Calculator⁴).

target	extinction $A(V)$
object 1	8.262
object 2	0.416
object 3	0.139
object 4	0.658
object 5	4.154
object 6	0.187
object 7	0.117
object 8	0.223
object 9	1.643
object 10	11.321
object 11	1.305
object 12	0.15
object 13	3.127
object 14	1.055
object 15	0.067

You are also encouraged to use online resources and literature to research the spectral characteristics and peculiarities of different astronomical objects, which should help you with the identification.

Stellar Sources

The shape of the continuum is mostly governed by the objects' surface temperatures. In the Harvard stellar classification the sequence O B A F G K M defines stellar spectral types with decreasing effective temperature. O, B, and A type stars are historically referred to as early types, while cool stars (G, K, and M) are known as late type stars. Main sequence stars of spectral type O show temperatures above 30,000 K, while late type stars of type M are significantly cooler between 2,400 K and 3,700 K. Although the chemical composition of main sequence stars is quite similar, the spectral features still vary. Typical appearances of stars of different spectral type are illustrated in Fig. 17a. Which features become prominent with decreasing temperature, is plotted in Fig. 17b. While the hottest O stars show a strong ultraviolet continuum and singly ionized helium lines in emission and/or absorption, neutral helium gains significance in B-typed stars. With decreasing temperature hydrogen lines become stronger, with a maximum strength around the spectral type A0. Moving further to F stars, metallic absorption lines become visible. These become more and more significant, until in K-typed stars the metals dominate the spectrum. Going further to lower temperatures, molecular bands, such as those of titanium oxide, get noticeable. Some further characteristics of stars of different stellar types are summarized in Table 6, which together with wavelengths of the most important features (Balmer series H_α , H_β , H_γ , H_δ , H_ϵ , He I

⁴https://ned.ipac.caltech.edu/extinction_calculator

Table 6: Characteristic features of stellar spectral types.

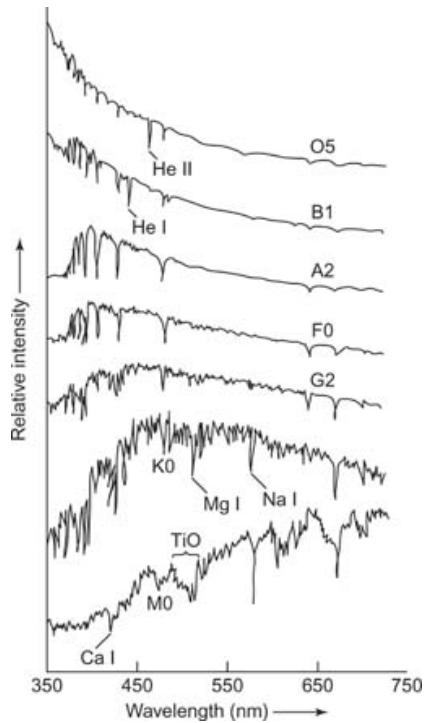
Type	Color	Eff. Temperature [K]	Main Characteristics
O	blue	25,000 - 50,000	Singly ionized helium lines either in emission or absorption; Strong UV continuum.
B	blue	11,000 - 25,000	Neutral helium lines in absorption
A	blue	7,500 - 11,000	Hydrogen lines at maximum strength for A0
F	white	6,000 - 7,500	Metallic lines become noticeable
G	yellow	5,000 - 6,000	Solar-type spectra; Absorption lines of neutral metallic atoms and ions (Ca II) grow in strength
K	orange	3,500 - 5,000	Metallic lines dominate; Weak blue continuum
M	red	2,000 - 3,500	Molecular bands of titanium oxide noticeable
L	brown	1,000 - 2,000	Stars with masses insufficient to run the regular hydrogen fusion process (brown dwarfs) contain lithium that in hotter stars is destroyed
T	infrared	1,000	Cool brown dwarfs with methane in the spectrum

and He II lines, Ca I, Ca II H&K doublet, TiO bands) that you can look up online, should help with the identification of the stars in your sample and the determination of their spectral types.

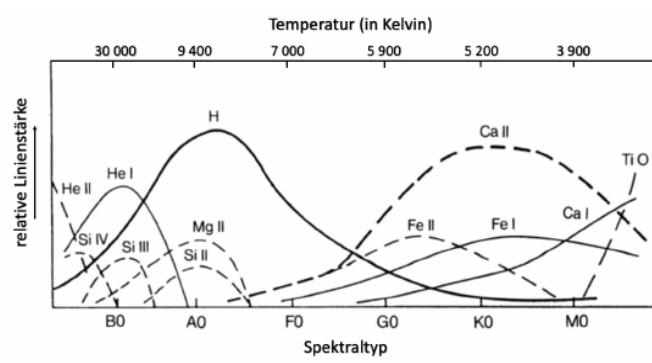
Another stellar sources hidden within the data sets are two kinds of white dwarfs. One of them shows a typical hydrogen dominated atmosphere, where mostly significant hydrogen absorption can be seen. The other is a white dwarf in a close binary system, a so called *nova*. The spectrum of such novae can differ, but Balmer emission serves as an indicator for such ones. In our case some helium emission is also noticeable.

Extragalactic Objects

As the spectra of galaxies are combinations of billions of stars, as well as some emission line regions, they are naturally multifarious and more difficult to identify. Overall, the spectra are relatively flat, which is a direct consequence of summing up many spectra of different temperatures. A break in the spectrum at around 4,000 Å due to absorption of high energy radiation from metals in stellar atmospheres and by a deficiency of hot, blue stars, is fairly common. It is very prominent in elliptical galaxies, only weak in spirals, and is not observed at all in irregular galaxies. In general, the features serve as indicators for the galaxies' stellar populations, while ionized gas reveals itself by emission. The most common absorption lines found in galaxies are those of the Ca II H&K doublet, the G-band at 4,304 Å, Magnesium at 5,175 Å, and Sodium at 5,894 Å. Prominent emission features can be found at wavelengths of the Balmer series, O II, O III, and S II. They typically are tracers for hot interstellar gas.



(a) Relative line strengths of absorption features with spectral type/eff. temperature. (Credit: Houghton Mifflin Harcourt, 2020)



(b) Relative line strengths of absorption features with spectral type/eff. temperature. (Credit: H. Karttunen, 1990)

Figure 17: Spectral stellar features with varying stellar temperature and spectral type.

A special type of galaxies within our sample are those with active galactic cores (or nuclei), which are referred to as AGNs. In those galaxies matter is accreted onto the supermassive black holes in their centers. Typically, a high amount of energy is released in this process, leading to radiation across the whole electromagnetic spectrum, from radio to X-ray frequencies. Furthermore, another non-thermal component in form of a relativistic jet along the black holes' rotational axis can be observed. Depending on the orientation of those galaxies with reference to the observer, different spectral characteristics are exposed. Quasars for example, which belong to the most luminous of the AGNs, outshining their host galaxies by far, are well-known for their significantly broadened emission lines, which indicate matter at high velocities of several thousands of km/s. Seyfert galaxies are significantly dimmer than quasars. Still, in particular the spectra of Seyfert-I galaxies are quite similar to standard quasars, as the inner region of the objects are visible. In contrast, in Seyfert-II galaxies the accretion discs are hidden behind the dust torus, and therefore the *broad-line* regions are not visible. Consequently, they typically show strong and narrow forbidden O III, N II, and S II emission lines. BL Lac objects are AGNs, observed more or less along their rotational axis. Their spectra are dominated by a relatively featureless non-thermal emission continuum over the entire electromagnetic range.

Redshift Determination

You should notice that in some of the extragalactic spectra the visible lines appear to be shifted towards longer wavelengths with respect to the rest frame. This is due to the cosmological redshift and could help you with the identification of the objects. Furthermore, once you found such a spectrum, you should be able to determine the amount of redshift z that is defined as:

$$z = \frac{\lambda_0 - \lambda_{\text{obs}}}{\lambda_0}. \quad (5)$$

Here, λ_0 stands for the wavelength of the feature in question at rest frame, and λ_{obs} for the corresponding wavelength, at which it is observed and measured. You can use the IRAF's `splot` task here to determine the redshift. Remember, within the plot of the spectrum you can zoom in by "z" and move (left/right) to the line in question by ",," and/or ".". To perform a fit to the line, move the cursor to the left-hand side of the feature and press "k" to mark the continuum. Repeat the same on the right-hand side of the line. A gaussian peak should be fitted and overplotted. The fit parameters are shown in the console. You can use the FWHM of the gaussian for error estimation. However, it is useful to determine the redshift from an average of several different features, as it can vary. This way your determination will be more robust. The redshifts of four of our targets, as found in literature, are listed in Table 7.

Table 7: Literature redshifts for comparison ([http://simbad.u-strasbg.fr/
simbad/](http://simbad.u-strasbg.fr/simbad/)).

target	redshift z
object 3	0.06995 ± 0.00007
object 6	0.002188 ± 0.000007
object 7	0.369 ± 0.003
object 12	0.01753 ± 0.00006