The E-ELT Design Reference Mission:
Direct imaging of terrestrial and giant exoplanets

S. Gladysz, R. Rebolo

Contents

1 DRM Science Case: Direct imaging of terrestrial and giant exoplanets
  1.1 The science case ................................................. 3
  1.2 Goals of the DRM simulations .................................. 3
  1.3 Metrics / Figures of merit ..................................... 3

2 DRM simulations
  2.1 Methodology ..................................................... 4
  2.2 Pipeline .......................................................... 5
     2.2.1 Adaptive optics .............................................. 5
     2.2.2 Static aberrations ........................................... 6
     2.2.3 Post-AO aberrations ........................................ 10
     2.2.4 Integral field spectrograph ................................. 11
     2.2.5 Coronagraphs .................................................. 12
     2.2.6 Post-processing ............................................... 14
  2.3 Inputs ............................................................ 14
     2.3.1 Scientific data ............................................... 14
     2.3.2 Technical data ............................................... 16
  2.4 Outputs .......................................................... 16

3 Results of simulations
  3.1 Simulation runs .................................................. 16
  3.2 Analysis .......................................................... 16
  3.3 Compliance with figures of merit ............................... 20
  3.4 Sensitivity to input parameters ................................. 20
## 3.4.1 XAO vs. SCAO

20

## 3.4.2 Coronagraphs: “Perfect” vs. APLC vs. Lyot

21

## 3.4.3 Distance to the star

23

## 3.4.4 Star’s spectral type

24

## 3.4.5 Influence of M1 aberrations

24

## 3.4.6 Influence of post-AO aberrations

25

### 4 Concluding remarks

25
1 DRM Science Case: Direct imaging of terrestrial and giant exoplanets

1.1 The science case

We propose to perform a direct imaging search for terrestrial and giant planets around 150 FGK stars within 20pc from the Sun. Using the E-ELT, coronagraphy and image post-processing we intend to achieve a contrast of $10^{-10}$ beyond an angular distance of $8\lambda/D$ (>50mas) from each parent star. The observations will be conducted to ensure detection (S/N=5) of the reflected light from Earth-like planets in the habitable zone of the parent stars (i.e. at 1AU orbital distance of the solar-type stars in the sample). This will allow a rather extensive demographic study of terrestrial and giant exoplanets over a wide range of separations leading to a better understanding of the formation mechanism of planetary systems.

1.2 Goals of the DRM simulations

The primary goal of this work is to test whether the E-ELT will be capable of directly imaging Earth-like planets in the habitable zones of the parent stars (i.e. at approximately 1AU from solar-type stars, and at smaller distances for K and M-type stars). This task is extremely challenging because it requires suppressing stellar light down to $10^{-10}$ of the peak PSF value at very small separations, e.g. 0.1″ for an Earth-like planet located at a distance of 10pc.

A slightly less challenging secondary goal of the simulations is to verify whether contrasts of the order of $10^{-9}$ could be reached on the E-ELT. Such capability would allow for detection of giant, Jupiter-type planets at 5AU. In both cases we assume that detections could be considered reliable if they are obtained at the S/N = 5 level.

Direct imaging and spectroscopy of exoplanets requires, for the most challenging and interesting objects, extreme adaptive optics (XAO). The XAO instrument for detection and characterization of exoplanets on the E-ELT is EPICS (Kasper et al. 2008). The study of exoplanets motivates the design of HARMONI (Tecza et al. 2009) and METIS (Brandl et al. 2008) too, although these two instruments will rely on modest adaptive optics (AO) correction provided only by the telescope (METIS will additionally rely on long wavelengths to deliver high Strehl ratio). In this report we do not focus on any specific instrument design. The goal of this work is to sample a large parameter space, and not to have a full end-to-end model of one planned E-ELT instrument. We want to know how parameters such as: level of static aberrations, efficiency of coronagraphic diffraction suppression, and distance to the target impact the final contrast curves.

1.3 Metrics / Figures of merit

We make use of the most common metric in the high-contrast-imaging field, i.e. the contrast curve. This curve shows level of suppression of stellar light as a function of angular distance from the star. It gives 5σ detection limits as a function of
angular separation from a star, and tells us which exoplanets are detectable at a given separation from the star.

2 DRM simulations

2.1 Methodology

We follow the work of Cavarroc et al. (Cavarroc et al. 2006) where the authors analysed limitations for the detection of exoplanets with a general model of an ELT with a 30m mirror. Instead, we focus on the specific design of the E-ELT.

All simulations are polychromatic, and they assume the same spectral range: 0.9 - 1.7µm. Planetary spectra are not simulated as part of this project but we make use of 100 channels in the afore-mentioned spectral range to calibrate and subtract residual stellar contribution after AO and coronagraphy, thus aiding the detection of the faintest planets (Section 2.2.6). The spectral range was pre-determined by the additional objective of this project: a cross-check with the EPICS simulation results. Originally, the 0.9 - 1.7µm range was chosen by the EPICS team because it contains the most interesting molecular bands: O₂, CO₂, H₂O, and CH₄. It should also be mentioned that the planet finder instrument for the Very Large Telescope (VLT), i.e. SPHERE, will have an integral field spectrograph (IFS) working in the similar spectral range (Fusco et al. 2006).

We achieve high dynamic range in the final, single image through three simulated techniques:

- Image sharpening through turbulence compensation with XAO (0.2m inter-actuator spacing). The impact of using telescope’s AO (0.5m inter-actuator spacing) is also explored (Section 3.4.1). In Section 2.2.1 we explain how the action of AO on the turbulent atmospheric phase can be modelled in an analytic fashion.

- Diffraction suppression through coronagraphy. It should be mentioned here that efficient coronagraphic starlight suppression is helped by first reducing the image to a narrow spot with AO, i.e. the second step in our simulations is intrinsically linked to the first. If AO works as expected the diffraction-limited spot hits the focal-plane mask and most of the light is blocked. We consider a “perfect” coronagraph, which is a theoretical construct removing the coherent portion of the stellar light (Fusco et al. 2006), the classical Lyot coronagraph (Lyot 1939), and the apodized pupil Lyot coronagraph (APLC) (Soummer 2005). In Section 2.2.5 the differences between the coronagraphs are explained and their impact on the achievable contrast is quantified in Section 3.4.2.

- Finally we implement a post-processing technique to remove the residual stellar light from the spectral data cube (Section 2.2.6). Here we use the method called spectral deconvolution (Sparks & Ford 2002) which has been successfully tested on spectral data cubes recorded with the SINFONI IFS on the VLT (Thatte et al. 2007). In all our simulations we assume an IFS is present but we do not simulate spectra and we do not have a detailed model of light
propagation through a spectrograph. Such a model depends on the design of an IFS (BIGRE or TIGER configuration, see (Antichi et al. 2009)) and is certainly outside the scope of this work. For the same reason we do not model Fresnel propagation, which depends on the number and relative separations of optical surfaces within an instrument (Marois et al. 2008).

2.2 Pipeline

2.2.1 Adaptive optics

For the wavefront sensor (WFS) we implemented the Shack-Hartmann device. This is the most widely-used WFS concept. The main component of this WFS is a lenslet array in the conjugate pupil plane. Each lenslet produces an image on the detector. The displacements of the images from the reference positions can be translated to average slopes of the incoming wavefront over individual sub-apertures (lenslets).

The two main shortcomings of the Shack-Hartmann WFS relevant for high-contrast imaging are aliasing and noise propagation. Aliasing is the result of the finite sampling provided by the Shack-Hartmann WFS: the wavefront is sampled with a grid of lenslets of linear dimension $d$, and in theory perfect wavefront reconstruction should be possible for aberrations with spatial frequencies up to $1/2d$ if the aberrated phase has no spatial-frequency content above $1/2d$. If aberrations are not band-limited, and this is always the case with atmospheric phase perturbations, aliasing occurs, i.e. high-frequency, “unseen” content is folded-back into the low-frequency range by the WFS. The control system will mistakenly try to correct these phase components which will lead to more power in the low-frequency range of the corrected wavefront. Since in the Fraunhofer approximation image formed by a lens is given as the Fourier transform of the electric field in the pupil plane, these aliased low-order aberrations will accumulate near the centre of the PSF. This is certainly undesired because this region of the image is particularly interesting in high-contrast imaging: this is where Earth-like planets are expected to be found. The solution to the aliasing problem is relatively simple: use a low-pass filter in the focal plane before the Shack-Hartmann WFS, thus removing the high-frequency content and making the phase band-limited. This concept is called the spatially filtered Shack-Hartmann WFS (Poyneer & Macintosh 2004) and it is used in our simulations.

All WFS work with light coming from a guide star and therefore wavefront reconstruction will be imperfect for fainter stars when photon noise starts playing a role. The second disadvantage to using the Shack-Hartmann WFS is that its capability to correctly sense low-order aberrations is weak because of photon noise (Verinaud et al. 2005). The pyramid WFS (Ragazzoni 1996) has a better response to photon noise for large-scale aberrations, and in theory could provide a gain in contrast of up to ten magnitudes for the smallest separations (Verinaud et al. 2005). However, end-to-end simulations have shown that the performance of the pyramid WFS is comparable, and sometimes poorer than that of the Shack-Hartmann device (Fusco et al. 2006). Because of that, and because of the existing simple and well-tested Shack-Hartmann implementation in the PAOLA simulation package used throughout this project it was decided that aliasing-free Shack-Hartmann model could approximate the behaviour of more favourable WFS to be developed during the construction of the
E-ELT.

The choice of the WFS wavelength calls for a trade-off study (Fusco et al. 2006). Briefly: using visible wavelengths for sensing and near-IR for imaging causes chromatic aberrations to appear in the corrected wavefronts. On the other hand no science photons are lost when distinct spectral ranges are used for sensing and science. Also, fast, low-noise detectors for the visible band are already available on the market, while similar detectors for the near-IR are currently in the development phase. The higher background noise for near-IR wavelengths also impacts the choice, as does the fact that stars of spectral types from G5 to M5 - which are of interest in exoplanet surveys - produce less photons in the near-IR portion of the spectrum compared to the visible (Fusco et al. 2006). For these reasons the Shack-Hartmann WFS simulated in this study uses visible wavelengths in the range 0.6 - 0.9µm.

The action of AO is modelled in an analytic fashion. For example the AO fitting error can be described as a result of the high-pass transfer function acting on the Kolmogorov atmospheric power spectrum of phase. Stars of various magnitudes (m_v = 2.5 to 10.5, covering spectral types F, G, K and M at the distance of 10pc, or G-type stars at distances 10 to 50pc) are assumed to be observed on-axis therefore we consider the transfer functions corresponding to the fitting error, the WFS noise error, dispersion error, and the AO loop delay error, but neglect the anisoplanatic transfer function. The WFS aliasing effect is also neglected because we assume a “perfect” spatial filter in front of the WFS. The PAOLA simulation package (Jolissaint et al. 2006) was used to model the action of AO on the atmospheric power spectrum. The modelled atmosphere consists of ten layers. The heights, fractional strengths, and wind speeds in the layers were taken from the DRM technical database. Seeing was set to 0.7″ at 500nm. Zenith angle was set to zero. Scintillation was neglected because this effect will be much smaller than the other effects considered in this report (Fusco et al. 2006). Generally, no amplitude errors - such as the effect of dust on the optical surfaces - have been considered here. The parameters of the simulations are summarized in Table 1.

The AO-filtered phase spectrum was used to generate independent realizations of AO residual phase using a classic, Fourier-domain filtering algorithm (McGlamery 1976). To this phase two static components have been added: static aberrations from the main mirror’s segments, and the non-common-path aberrations not sensed by AO and modelled with an f^{-2} spectrum (with f being the spatial frequency) (Cavarroc et al. 2006). The next section describes the nature of optical aberrations to be expected on the E-ELT and simplifications used in the simulations.

2.2.2 Static aberrations

The E-ELT has a five-mirror design and naturally all five mirrors will introduce some aberrations. Nevertheless in our simulations we consider only M1 aberrations because these will be by far the most significant. The same simplification has been used in the thirty Meter Telescope (TMT) XAO simulations (Troy et al. 2006; Crossfield & Troy 2007). Firstly we describe the strategies for phasing and control of M1 which are currently being investigated at ESO. This is by no means a complete description, and the reader is encouraged to consult the relevant documents referenced
in this section. The main procedures are defined here as follows:

- M1 phasing through optical sensors, segment position actuators and warping harnesses,
- M1 control through edge sensors and position actuators.

**M1 phasing** Phasing of the main mirror serves two purposes:

- to remove piston and tip/tilt errors between segments, and shape errors on the individual segments, thus increasing the coherency of the wavefront,
- to calibrate set points of the edge sensors (ES).

The baseline concept for the phasing sensor is a Shack-Hartmann WFS with 19 inner sub-apertures covering the segment, and 12 phasing sub-apertures straddling the inter-segment edges (Bonnet 2009a). The phasing sub-apertures will sense the distortion of the diffraction pattern due to the inter-segment step errors. To do so the sensor must be operated in the diffraction-limited regime which will be the case since the sub-aperture size is of the order of the Fried parameter $r_0$ at the operating wavelength of 800nm (Bonnet 2009a). The edge measurements have two main advantages over the traditional Shack-Hartmann slope measurements (Chanan 2009). Firstly, they are only sensitive to wavefront discontinuities which can only come from the segmented primary mirror, and therefore they are not affected by the aberrations on the other mirrors. This removes the complication of having to disentangle which aberrations came from which mirror. Secondly, for the same reason they are insensitive to the atmospheric aberrations. Since the distortion of the PSF is controlled by the relative edge height, the phasing algorithm compares the measured PSF shape to the one predicted by diffraction theory. Actuator commands deduced from the piston values are then sent to the segment position actuators (PACTs). Three PACTs control piston and tip/tilt of one segment.

The inner sub-apertures will be used to sense the shape errors of the segments, for example: defocus and astigmatism aberrations which could be produced by the spatial inhomogeneities of the coefficient of thermal expansion within one segment (Swat, priv. comm.). The measurements will be transformed to signals for the warping harnesses (WH) which will control the shapes of the segments with 24 actuators (Müller 2008). There is some ambiguity regarding the WH reduction factors (rms output aberration / rms input aberration) and the spatial structure of the resulting segment error (Troy et al. 2006; Müller 2008). Here we will assume the WH reduction factors predicted by simulations carried out at ESO (Müller 2008). They translate to nm-level aberrations after correction of realistic input loads due to gravity or thermal effects. Regarding the spatial frequency content of the output error we will follow ESO predictions and assume the output segment shape to be independent of the input shape. For example, correction of the simple tilt due to gravity results in a complicated, high-spatial-frequency shape which bears no resemblance to the input aberration (Müller 2008). Since the simulated output segment shapes are characterized by relatively high spatial frequencies, between approximately 5 and 10 cycles per segment, the resulting speckles in the PSF will
cover the outer regions of the image which are less interesting from the high-contrast point of view (see beginning of Section 2.2.1). In our simulations we therefore ignore correction limitations of the WH.

As regards the second function of the phasing procedure, segment alignment process described above calibrates the edge sensors on the segments. An edge sensor is a combination of one emitter and one receiver belonging to two adjacent segments. In the current E-ELT design there are two ES pairs per segment edge. ES will be used to maintain the figure of M1 which will be perturbed by wind buffeting, thermal effects and gravity. ES are only sensitive to inter-segment piston variations. Because they deliver a relative measurement they have to be calibrated. At the E-ELT this will be done using the optical phasing sensor described above. The role of ES in M1 figure control is explained below. Here we will only mention that their accuracy can reach sub-nm level (Dimmler 2009; Bonnet 2009b). As explained below, accuracy of optical phasing with a WFS, rather than the ES measurement precision, will dominate the error budget of the aligned, off-line telescope.

The level of phasing precision to be expected at the E-ELT is open for debate. On one hand results from phasing of the Keck telescope point to the possibility of achieving an accuracy of tens of nm (Crossfield & Troy 2007). This seems to be supported by the results of the Active Phasing Experiment at ESO (Gonte 2009). On the other hand simulations of ELTs alignment yield post-correction rms values on the order of hundreds of nm (Chanan 2009), or microns (Piatrou & Chanan 2010) of wavefront error on M1. Importantly, the aberrations - or rather mirror modes - left over after alignment are all of the low-order type: global tip/tilt, defocus, astigmatism, etc. (Piatrou & Chanan 2010). They are left uncorrected on purpose because their “observability” is low: attempts to measure them would result in noisy reconstruction and oscillations of the global shape of M1, which would have a very strong effect on image quality. Instead they are allowed to propagate through the main optics, up to the AO system which in principle should have excellent response to low-order aberrations. As discussed in the next section the simulations presented in this report implement low-order, high-amplitude M1 aberrations.

In the currently adopted strategy phasing of the E-ELT primary mirror will be performed about once a month. Provisions are made for a back-up strategy where phasing is done also during the science observations. At the TMT phasing will be carried out at monthly intervals outside the science queue. The procedure is expected to last no longer than two hours (Piatrou & Chanan 2010).

**M1 control** During the science operations dynamical perturbation of the M1 figure will be measured by ES and corrected with PACTs with a bandwidth of 1-3Hz. Naturally, the resulting accuracy should be better for slow aberrations. These are due to temperature variations (current ESO assumptions: seasonal amplitude of ambient temperature variations of ±15K; ambient temperature change during tracking of ±1K/hr), and gravity-load changes during tracking (E-ELT’s operational pointing range is 0 to 70deg zenith angle). In simulations, after execution of the ES-PACT loop for these quasi-static perturbations the main mirror assumes shapes characterized by low-spatial-frequency aberrations (Bonnet 2009c). Same is true for fast perturbations due to wind buffeting (ESO assumptions: wind speed 10m/s outside,
and 1.6m/s inside the dome), although the degree of correction is not spectacular. Similarly to the phasing loop, the M1 controller does not attempt to remove low-order aberrations - these are to be corrected by AO. The role of M1 controller is to compensate mainly for high-spatial-frequency errors which have small amplitudes and therefore their contribution to rms wavefront is also small.

It is expected that the wind-induced dynamical aberrations will dominate over aberrations generated by gravity and thermal effects. Hence for our simulations we adopt a model of segments’ piston, tip and tilt from a scenario of closed-loop M1 control system responding to wind disturbances alone (Dimmler 2009). ES noise is included in this scenario and perfect telescope phasing is assumed. The full model contains piston and tip/tilt values for every segment during the interval of 30 seconds, sampled every 5ms. From this set we take only one “instantaneous” M1 shape for our simulations. What is important here is the spatial structure of the closed-loop aberrations shown in Figure 1. These low-order distortions are a good representation for a class of slowly-varying perturbations to be expected to leak-through M1 control. It is important to have them in an XAO simulation since their residuals after AO will focus light close to the PSF core, which is the most interesting region in the image.

As will be explained later, high-contrast imagery is dominated by quasi-static speckles which are the result of quasi-static aberrations. While the wind-induced perturbations have timescales less than one minute, which is the timescale adopted in our simulations, and temporal noise averaging in the final image would be in practice quite efficient, we argue that the complete set of controllable aberrations (gravity, temperature variations and wind) could be assumed constant within the time-span of one minute. In the future we plan to integrate into the simulations models of these aberrations under the action of M1 controller, when such models become available. Currently we assume that M1 aberrations are completely static on the time-scale of one minute.
2.2.3 Post-AO aberrations

The spatially-filtered WFS employed in the simulations has a “perfect” transfer function, i.e. it sets to zero all aberrations up to its cut-off frequency, which by the Shannon sampling theorem is $1/2d$ for the inter-actuator spacing $d$. In the XAO case this means no PSF deformations (speckles) should be present up to $0.45''$ from the PSF centre. Still, in real systems speckles do appear within the AO control region: they are due to non-common-path aberrations, which are not sensed and therefore not compensated for by AO (Cavarroc et al. 2006; Soummer et al. 2007). These instrumental errors are usually slowly-varying and they generate quasi-static speckles, which are the biggest obstacle in detecting planets. They are too slow to efficiently average-out in long exposures, and too fast to be completely subtracted-out through PSF calibration (Marois et al. 2006; Racine et al. 1999).

We simulate these aberrations with an $f^{-2}$ spectrum which has been empirically obtained on optical surfaces. The spectrum is flat in the range $0 < f < 1/2d$. Beyond the cut-off frequency it decreases towards zero. Similarly to the case of atmospheric aberrations this spectrum is used to filter white noise to give one realization of non-common-path aberrations, i.e. a phase-screen (McGlamery 1976). The total rms power in this phase-screen is $30\text{nm}$. This adopted value is consistent with commonly assumed levels of non-common-path aberrations (Cavarroc et al. 2006; Marois et al. 2008; Verinaud 2010). This phase remains static during one simulation run.

Here we come to the problem of choosing the time-scale for the quasi-static aberrations. In this work we obtain images corresponding to one minute of integration time, compute contrast curve, and then scale this curve to 10 hours. The total variance of speckle noise at some separation from the PSF centre in an image corresponding to one minute is:

$$\sigma_f^2 = \sigma_{qs}^2 + \frac{\tau_{qs}}{\tau_a} \sigma_a^2,$$

where $\sigma_{qs}$ is the standard deviation of quasi-static speckles at the location under investigation, $\sigma_a$ is one realization of AO-filtered atmospheric speckles, and $\tau_{qs}$ and $\tau_a$ are the corresponding time-scales (one minute and 0.5sec, respectively - see Section 2.2.4). In 10 hours noise variance will be:

$$\sigma_f^2 = \frac{10\text{hrs}}{\tau_{qs}} \left( \sigma_{qs}^2 + \frac{\tau_{qs}}{\tau_a} \sigma_a^2 \right),$$

What happens when $\tau_{qs}$ is in reality twice as long as the chosen value?

$$\sigma_f^2 = \frac{10\text{hrs}}{2\tau_{qs}} \left( 4\sigma_{qs}^2 + \frac{2\tau_{qs}}{\tau_a} \sigma_a^2 \right) = \frac{10\text{hrs}}{\tau_{qs}} \left( 2\sigma_{qs}^2 + \frac{\tau_{qs}}{\tau_a} \sigma_a^2 \right),$$

(The same argument could be applied to $\tau_a$). This means that the results of simulations based on too-short time-scales will underestimate the noise in very long exposures. Unfortunately there is very little information in the literature about the temporal properties of long-lives speckles. The reported values range from few seconds to few hundred seconds (Hinkley et al. 2007). As mentioned earlier we adopt a value of 60sec, which is probably optimistic. To illustrate the effect of two different time-scales of quasi-static aberrations we also ran one simulation for an integration
time of 30sec. Figure 2 shows the impact of $\tau_{qs}$ choice on the final contrast curve, after coronagraphy (Section 2.2.5) and spectral deconvolution (Section 2.2.6). As expected the gain in contrast caused by assuming a shorter $\tau_{qs}$ of 30sec is slightly smaller than the factor of two (compare Equations (2) and (3)). Since we expect our order-of-magnitude guess for $\tau_{qs}$ to be optimistic but reasonable, the results reported here will also have an order-of-magnitude correspondence to real observations. In the future we plan to investigate the auto-correlation of quasi-static speckles induced by a full model of M1 aberrations (dynamical wind perturbations, quasi-static PACT response to gravity while tracking, etc.) similarly to work that had been carried out for filtered atmospheric aberrations after XAO (Macintosh et al. 2005; Poynier & Macintosh 2006).

2.2.4 Integral field spectrograph

To simulate an IFS we needed to have the wavelength-scaled wavefronts. After one AO-corrected wavefront is generated for the first wavelength of 0.9\(\mu\)m, the programme loops over the 100 channels in the spectral range, scaling the stored phase by each channel’s wavelength:

$$\phi(\lambda_i) = \phi(\lambda_0)\lambda_0/\lambda_i,$$

(4)
where $\phi$ stands for phase of the wavefront, and $\lambda$ is wavelength ($\lambda_0 = 0.9\mu m$). In the simulations this loop over wavelengths is contained inside the loop over time: 120 iterations over instantaneous wavefronts. The resulting instantaneous PSFs are scaled by flux expected for an integration time of 0.5sec. We take this as a guess for an XAO speckle lifetime. It has been shown \cite{Macintosh05, Poyneer06} that second-order speckles, which dominate coronagraphic images, are significantly slower than pure atmospheric speckles. These XAO coronagraphic speckles have timescales set by the atmospheric clearing time of the telescope, $\tau \approx D/\nu$, where $\nu$ is the wind speed. The value of 0.5sec, while consistent with estimates reported for simulated XAO systems on 8m-class telescopes \cite{Macintosh05}, is probably optimistic for a 42m aperture. Nevertheless, the characteristics of the contrast curves in the region of interest are dominated by the quasi-static aberrations and not by the smooth halo generated by AO-filtered atmospheric perturbations.

Equation (4) excludes differential chromatic aberrations between the channels. M1 was sampled with $1024 \times 1024$ pixels inside a $2048 \times 2048$ zero-filled array. To keep the constant plate scale ($\lambda/2D$ at 0.9$\mu$m) the wavefronts for consecutive wavelengths were embedded in ever-increasing arrays of zeros before Fourier-transforming. For $m_v = 4.7$ star the wavefronts containing AO-filtered atmospheric and M1 aberrations, together with non-common-path-errors resulted in Strehl ratios between 85 and 96% in the spectral range 0.9 - 1.7$\mu$m. In the SCAO simulations these Strehl ratios were 55 and 85%.

2.2.5 Coronagraphs

Firstly we consider the “perfect” coronagraph in order to explore the theoretical limitations to high-contrast imaging at the E-ELT. One can model such a coronagraph as a subtraction in the pupil plane of the perfect, flat wavefront \cite{Fusco06}. Since coronagraphs can only remove the coherent part of the incoming electric field, as quantified by the Strehl ratio, this idealized subtraction is less efficient for less-coherent wavefronts:

$$\text{PSF} = \left| FT \left( Pe^{i\phi} - \sqrt{SR} Pe^{i\psi} \right) \right|^2,$$

where $\phi$ stands for phase, $P$ is the binary pupil function, and SR stands for the Strehl ratio. This “perfect” coronagraph is completely achromatic and therefore the resulting multi-wavelength images can be very efficiently “cleaned” by spectral deconvolution (Section 2.2.6).

The action of a more realistic Lyot-style coronagraph is simulated in the following way. For each wavelength the electric field is Fourier-transformed to give the amplitude PSF in the focal plane. The result is multiplied by the representation of the focal-plane mask, taken to be 7.5 $\lambda/D$ in diameter. This value has been optimized for the inner working angle of 3.9 $\lambda/D$ and the current design of the E-ELT main mirror, i.e. 30% linear obscuration ratio \cite{Martinez08}. The electric field after the focal-plane mask is inverse Fourier-transformed and multiplied by the model of the Lyot stop. For the Lyot coronagraph this stop has an oversized central obscuration and undersized outer edge. Diffraction patterns induced by the spiders are also covered with masks of twice the size, i.e. 0.8m. The throughput of the Lyot
coronagraph is 63%. The corresponding attenuation of the peak of the planetary PSF is included in the computation of the contrast curve.

The more suitable APLC coronagraph has an apodizing mask in the first pupil-conjugate plane after XAO. This mask suppresses the Airy rings in the subsequent focal plane, and scatters non-coherent light outside the next pupil-conjugate plane where the stop is located. That is why for the APLC coronagraph the stop is not undersized. Only the diffraction from spiders is blocked with masks of twice the size. Transmission of the APLC coronagraph is determined by the apodizer and reaches 55%. The design was optimized for the current E-ELT configuration and the inner working angle of 2.4 λ/D (Martinez et al. 2008). Figure 3 shows the apodizer and the pupil-plane stop for the APLC configuration.

Fourier transform of the field after the Lyot stop gives the electric field in the detector plane, and the squared modulus of this quantity yields the final on-axis instantaneous PSF. For the images of the test planets the electric field is propagated only through the Lyot stop (and the apodizer for APLC), giving the off-axis PSF. We scale and shift this off-axis PSF to simulate seven planets located at various separations from the star, within the range of 0.1 - 0.5″. The contrast ratios are between 10⁶ and 10⁷. These instantaneous images of planetary systems are stored for each channel and then co-added, resulting in a spectral data cube with 100 images, each 60sec long. Finally, we loop over this cube adding Poisson, background (15mag/arcsec²) and readout noise (10e-).

Spatial dependence of the coronagraphic transmission has to be taken into account when spatially-dependent metrics, such as contrast curves, are considered. For this, one has to simulate off-axis propagation with tilted wavefronts. We do this by adding a phase term in the Fourier space to a perfectly flat wavefront. This tilted wavefront is then propagated through a coronagraph. This is done for a range of tilts. Transmission as a function of the angular separation in the focal plane is then stored for both the Lyot and the APLC designs and it is used in the computation.
2.2.6 Post-processing

Spectral deconvolution is applied to the IFS data cube to increase the contrast by two to three orders of magnitude and recover the signal of the faintest test planets. Briefly: in the simulated data cube the residuals of the Airy pattern expand with wavelength, and the planet stays fixed. Each image in the cube is spatially scaled by Fourier interpolation to fit a common grid, so that the Airy rings are now aligned and the planet’s position changes through the cube: it gets closer to the star with increasing wavelength. A 2nd-order polynomial is fitted for each spaxel (spectral vector in the scaled cube). This particular choice of the polynomial order is justified by the following argument: The wavefront expressed in units of length, which to a first order can be assumed to have the same optical path difference across the 100 channels, is scaled by $2\pi/\lambda$ to give phase. Theoretically, response of intensity for one constant location on an Airy ring to change in wavelength of the wavefront should be $\propto (2\pi/\lambda)^2$, because image intensity is the power spectrum of the electric field. This argument ignores changes in PSF morphology which naturally occur for a wide wavelength range where the Strehl ratio changes significantly. Nevertheless these changes are smooth when one considers many channels, as we do here, and can be fitted with a moderately smooth polynomial. After the first fitting step, any outliers at the 2$\sigma$ level (possible planets) are removed from the vector, and this procedure is repeated until no value in the sample exceeds the fit by two standard errors. Once the fit is not skewed anymore by possible signals, it is subtracted from the spaxel under consideration - this step removes most of the stellar modulation, but leaves possible planetary signal intact. The images are then re-scaled to the original grid, so that now a planet would be aligned on the same pixel again. Finally the cube is collapsed to produce one final image.

2.3 Inputs

2.3.1 Scientific data

Young and hot exoplanets are the primary targets for the SPHERE instrument (Boccaletti et al. 2008). XAO at the E-ELT will be geared towards the detection of old and cold planets, similar to the ones in the Solar System. These targets are significantly more challenging. The contrasts to be expected are between $10^7$ and $10^{10}$, depending on the size of the planet and its physical distance from the star. Old planets do not emit their own light; they only reflect light generated by the stars. Therefore the planet/star flux ratio, $F_p / F_s$, is (Brown 2004):

$$\frac{F_p}{F_s} = p(\lambda) \Phi(\beta) \left(\frac{R}{r}\right)^2,$$

where $p$ is the geometric albedo of a planet, which depends on wavelength, $R$ stands for the radius of an exoplanet, $r$ is its separation from the star in AU, and $\Phi(\beta)$ is the phase function of the planet which depends on the phase angle $\beta$ (angle at the
Figure 4: Left: single 60sec-long exposure at 0.9µm. Right: image formed by collapsing the full spectral data-cube after processing using spectral deconvolution; all seven planets are now clearly visible. Star G2 (m_v = 4.7), distance 10pc, “perfect” coronagraph, 1min integration time. Images are displayed on a linear scale, and artificially saturated to bring out the faint details. Colours are inverted for better contrast.

Figure 5: Contrast curves before, and after spectral deconvolution (SD), for the case introduced in Figure 4 and scaled to 10 hours.
planet between star and observer):

\[ \Phi(\beta) = \frac{\sin(\beta) + (\pi - \beta) \cos(\beta)}{\pi}, \]  

(7)

Assuming an optimistic case of planet observed at opposition ($\Phi(\beta) = 1$) and an near-IR albedo of 0.3 we get for Jupiter ($R = 0.00048\text{AU}$, $r = 5.2\text{AU}$) the expected contrast of $2.5 \times 10^{-9}$. For Earth we get $5.6 \times 10^{-10}$. In both cases we assumed distances between planets and their parent star identical to the ones in our system. What happens when one considers smaller stars?

The habitable zone is expected to be found closer for smaller stars (Selsis et al. 2007). For M-type stars it could be located as close as 0.1AU. Assuming the same spatial scaling applies to other, heavier planets in this hypothetical planetary system around an M-type star we get contrasts of $2.5 \times 10^{-7}$ for “Jupiter” and $5.6 \times 10^{-8}$ for “Earth”. “Super-Earths”, i.e. planets with masses several times greater than the Earth’s mass would have contrasts around $2.6 \times 10^{-7}$ in this M-type star’s system.

### 2.3.2 Technical data

This DRM science case requires more technical data than there is provided in the DRM technical database. The additional parameters and their values are listed in Table 1.

### 2.4 Outputs

The single image resulting from spectral deconvolution (see right panel in Figure 4) is processed to yield azimuthally averaged contrast curves. Contrast curve is simply a linear function of the noise standard deviation, normalized by the peak of the on-axis un-obscured PSF integrated over all wavelengths. It gives 5σ detection limits as a function of angular separation from a star. Figure 5 shows contrast curves corresponding to the images shown in Figure 4.

### 3 Results of simulations

#### 3.1 Simulation runs

In Table 2 we summarize the various simulation runs performed as part of the DRM science case.

#### 3.2 Analysis

With the standard set of simulation parameters presented in Table 1 and the APLC coronagraph we obtain a contrast curve for a G-type star (Figure 6) and an M-type star (Figure 7) both located at a distance of 10pc. We over-plot expected fluxes from the three classes of planets discussed in Section 2.3.1. If a symbol representing an exoplanet is above a contrast curve this means this particular type of planet would be detected at the $S/N = 5$ level.
Table 1: Values of the parameters used in the simulations.

<table>
<thead>
<tr>
<th>Input parameters</th>
<th>XAO simulation</th>
<th>SCAO simulation</th>
</tr>
</thead>
<tbody>
<tr>
<td>Telescope</td>
<td>42m</td>
<td>-</td>
</tr>
<tr>
<td>Central obscuration</td>
<td>12.5m</td>
<td>-</td>
</tr>
<tr>
<td>Thickness of the spiders</td>
<td>0.4m</td>
<td>-</td>
</tr>
<tr>
<td>Number of segments</td>
<td>984</td>
<td>-</td>
</tr>
<tr>
<td>M1 aberrations due to wind(^b)</td>
<td>0.3(\mu)m rms piston - 0.006&quot; rms tip - 0.006&quot; rms tilt -</td>
<td>-</td>
</tr>
<tr>
<td>Atmosphere</td>
<td></td>
<td></td>
</tr>
<tr>
<td>Seeing at 0.5(\mu)m</td>
<td>0.7&quot;</td>
<td>-</td>
</tr>
<tr>
<td>Outer scale</td>
<td>25m</td>
<td>-</td>
</tr>
<tr>
<td>Sky</td>
<td></td>
<td></td>
</tr>
<tr>
<td>Background noise in H band</td>
<td>15\text{mag/arcsec}(^2) -</td>
<td>-</td>
</tr>
<tr>
<td>AO system</td>
<td></td>
<td></td>
</tr>
<tr>
<td>Type of WFS</td>
<td>Spatially filtered</td>
<td>Spatially filtered</td>
</tr>
<tr>
<td></td>
<td>Shack-Hartmann</td>
<td>Shack-Hartmann</td>
</tr>
<tr>
<td>Number of WFS sub-apertures</td>
<td>200 \times 200</td>
<td>84 \times 84</td>
</tr>
<tr>
<td>WFS integration time</td>
<td>0.33ms</td>
<td>2ms</td>
</tr>
<tr>
<td>AO lag(^c)</td>
<td>1ms</td>
<td>6ms</td>
</tr>
<tr>
<td>WFS read-out noise</td>
<td>0.5e-</td>
<td>3e-</td>
</tr>
<tr>
<td>WFS spectral range</td>
<td>0.6 - 0.9(\mu)m</td>
<td>0.6 - 0.9(\mu)m</td>
</tr>
<tr>
<td>Non-common-path aberrations(^d)</td>
<td>30nm rms</td>
<td>30nm rms</td>
</tr>
<tr>
<td>Integral field spectrograph</td>
<td></td>
<td></td>
</tr>
<tr>
<td>Spectral range</td>
<td>0.9 - 1.7(\mu)m</td>
<td>-</td>
</tr>
<tr>
<td>Number of spectral channels</td>
<td>100</td>
<td>-</td>
</tr>
<tr>
<td>Optical transmission to the detector</td>
<td>0.1</td>
<td>-</td>
</tr>
<tr>
<td>Read-out noise</td>
<td>10e-</td>
<td>-</td>
</tr>
<tr>
<td>Pixel scale</td>
<td>\lambda/2D at 0.9(\mu)m = 0.002&quot;</td>
<td>-</td>
</tr>
</tbody>
</table>

\(^a\)SCAO refers here to the AO correction provided by the telescope.

\(^b\)rms wavefront: these are correlated, low-order aberrations on the main mirror, left uncorrected by the M1 control system, see Section 2.2.2.

\(^c\)AO lag equals WFS read-out time + wavefront reconstruction time.

\(^d\)Simulated with an \(f^{-2}\) power spectrum: flat in the AO working range, and decreasing beyond the AO cut-off frequency.
Table 2: Simulation runs.

<table>
<thead>
<tr>
<th>Run number</th>
<th>Values of the investigated parameters (cf. Table 1)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>APLC, SCAO</td>
</tr>
<tr>
<td>2</td>
<td>APLC, XAO</td>
</tr>
<tr>
<td>3</td>
<td>Perfect coronagraph, XAO</td>
</tr>
<tr>
<td>4</td>
<td>Lyot coronagraph, XAO</td>
</tr>
<tr>
<td>5</td>
<td>APLC, G2, 10pc</td>
</tr>
<tr>
<td>6</td>
<td>APLC, G2, 30pc</td>
</tr>
<tr>
<td>7</td>
<td>APLC, G2, 50pc</td>
</tr>
<tr>
<td>8</td>
<td>APLC, F2, 10pc</td>
</tr>
<tr>
<td>9</td>
<td>APLC, K2, 10pc</td>
</tr>
<tr>
<td>10</td>
<td>APLC, M2, 10pc</td>
</tr>
<tr>
<td>11</td>
<td>APLC, G2, 10pc, $\tau_{qs} = 30$ sec</td>
</tr>
<tr>
<td>12</td>
<td>Perfect coronagraph, XAO, G2, M1 aberrations = 0.15µm rms</td>
</tr>
<tr>
<td>13</td>
<td>Perfect coronagraph, XAO, G2, M1 aberrations = 0.9µm rms</td>
</tr>
<tr>
<td>14</td>
<td>APLC, XAO, G2, M1 aberrations = 0.15µm rms</td>
</tr>
<tr>
<td>15</td>
<td>APLC, XAO, G2, M1 aberrations = 0.9µm rms</td>
</tr>
<tr>
<td>16</td>
<td>Perfect coronagraph, XAO, G2, post-AO aberrations = 10nm rms</td>
</tr>
<tr>
<td>17</td>
<td>Perfect coronagraph, XAO, G2, post-AO aberrations = 60nm rms</td>
</tr>
<tr>
<td>18</td>
<td>APLC, XAO, G2, post-AO aberrations = 10nm rms</td>
</tr>
<tr>
<td>19</td>
<td>APLC, XAO, G2, post-AO aberrations = 60nm rms</td>
</tr>
<tr>
<td>20</td>
<td>Perfect coronagraph, SCAO</td>
</tr>
<tr>
<td>21</td>
<td>Perfect coronagraph, XAO400</td>
</tr>
<tr>
<td>22</td>
<td>APLC, XAO400</td>
</tr>
<tr>
<td>23</td>
<td>Perfect coronagraph, XAO400, M2</td>
</tr>
</tbody>
</table>
Figure 6: Contrast curve for the G-type star at a distance of 10pc, APLC coronagraph, integration time = 10hrs. Expected fluxes from three types of planets are over-plotted.

Figure 7: Contrast curve for the M-type star at a distance of 10pc, APLC coronagraph, integration time = 10hrs. Expected fluxes from three types of planets are over-plotted.
3.3 Compliance with figures of merit

Figure 6 suggests that while it is likely that the E-ELT will make many discoveries of Jupiter-type, giant planets it seems that obtaining images of Earth-like planets will not be feasible with the XAO system, and especially with the APLC coronagraph, used throughout this report.

The same analysis repeated for an M-type star (Figure 7) shows the challenges associated with imaging of old and cold planets. It seems once again that direct imaging of rocky planets in the habitable zone would require a light-suppression system which is an order-of-magnitude more efficient than our combination of APLC and spectral deconvolution.

3.4 Sensitivity to input parameters

In the following sections we show how the change of a particular parameter impacts the final contrast curve. All contrast curves shown here correspond to an integration time of 10 hours.

3.4.1 XAO vs. SCAO

Here we look at the gain brought upon by increasing the actuator density. We ran one simulation with the SCAO system (cf. Table 1), and also two XAO simulations: one with the 200 × 200-actuator system, and one with XAO having 400 × 400 actuators. The other parameters of this “super XAO” system were the same as for the 200 × 200-actuator system. The simulations were then repeated for the “perfect” coronagraph.

Figure 8 suggests that at small separations - less than 0.05″ - APLC contrast curve seems to be independent of the level of AO correction. This is due to coronagraph chromaticity, which results in non-trivial spatial scaling of the images which cannot be removed by spectral deconvolution. Further away from the core the clearing of the halo by XAO brings an advantage of approximately one order of magni-

![Figure 8: Left: Effect of varying the number of actuators on the contrast curve for a G-type star at a distance of 10pc, “perfect” coronagraph. Right: Same analysis for the APLC coronagraph.](image-url)
tude. This is important for giant planets which are expected to have star/planet contrasts around $10^9$.

SCAO is not sufficient if one wants to detect planets in reflected light. This can be seen in curves for the “perfect” coronagraph and the APLC. On the other hand the effect of increasing the actuator density above $200 \times 200$ is very small. This can be understood if one remembers that decreasing the WFS lenslet size gives less photons for the wavefront reconstruction operation. While the fitting error will decrease, the WFS noise error will increase: the XAO400 system’s WFS consists of lenslets with size equal to 10cm, and records stellar sub-images at a frame rate of 3kHz. The Strehl ratios at the longest wavelength (1.7$\mu$m) are 96 and 98% for XAO200 and XAO400 respectively, and the difference in final contrast is minimal. Going from a $200 \times 200$-actuator system to a $400 \times 400$ XAO will not be helpful as far as detection of Earth-like planets is concerned.

3.4.2 Coronagraphs: “Perfect” vs. APLC vs. Lyot

Spectral deconvolution can in theory remove post-coronagraphic residuals as long as they are achromatic. The resulting image should be free of speckle noise. This is not the case with real, chromatic coronagraphs as illustrated in Figure 9. Chromaticity results in non-trivial spatial scaling of the images in the spectral data cube. This is visible primarily in the vicinity of the optical axis and this is where both the Lyot and APLC designs have significantly lower contrasts compared to the “perfect”, achromatic coronagraph.

Figure 10 emphasizes this point. Close to the optical axis chromaticity of the coronagraph is the “bottleneck”. The spatially-varying transmission of the coronagraph, i.e. the presence of the focal-plane mask and the resulting inner working angle, is insignificant. We compared Figure 10 to an equivalent plot but without taking into account the blocking by the focal-plane mask in computation of the contrast curve, and the differences between the figures were only visible at separations smaller than 0.01”. This means that at small separations the dominant factor is the chromaticity of the coronagraph. If spectral deconvolution is going to be the method of choice for high-contrast imaging, care has to be taken to use achromatic optics and to design “Fresnel-friendly” instruments (Verinaud 2010). Fresnel effects, i.e. phase aberrations translating to amplitude changes, cannot be removed by spectral deconvolution.

It should be mentioned here that EPICS has an apodizer-only design, i.e. no focal-plane mask is present (Verinaud 2010). This is precisely because of the effects outlined above. Nevertheless, APLC is considered very suitable for high-contrast imaging. Figure 10 shows that its performance, as quantified by the contrast curve, is almost ten times better than the classical Lyot coronagraph. This is due to the fact that APLC is significantly less chromatic than the Lyot design. Its chromaticity can be further suppressed by using a Gaussian focal mask, instead of the hard-edge stop (Verinaud 2010).
Figure 9: Results after spectral deconvolution for three different coronagraphs. Star G2 ($m_v = 4.7$), distance 10pc. Images are displayed on the same linear scale, with identical minimum and maximum values. Colours are inverted for better contrast.

Figure 10: Contrast curves for the three coronagraphs considered in this project. For the “perfect” coronagraph throughput equal to one has been assumed. Star G2 ($m_v = 4.7$), distance 10pc.
3.4.3 Distance to the star

When stars appear brighter three effects are expected on the contrast curves. Firstly, the relative level of Poisson noise should decrease. This can be explained as follows: Contrast curve is related to the standard deviation of stellar noise. When we consider two stars, the first twice as bright as the second, Poisson noise in the halo of the former should increase by a factor of square root of two, while the peak of the PSF will increase by a factor of two. Since contrast curve is normalized by the peak of the un-obscured PSF it will scale as \(2^{1/2}/2\), i.e. it will be steeper by a factor of 0.7 for the brighter star.

Secondly, a brighter star will provide more photons for the WFS and the AO correction will be better (assuming the AO error budget is dominated by WFS noise). The Strehl ratio will increase, that is the PSF will be sharper, and less photons will be scattered into the halo.

Thirdly, increased Strehl ratio will concentrate more light on the coronagraphic mask. This coherent light will be removed, and the suppressed Airy residuals will generate less photon noise and less “pinned” speckles - speckles which are modulated and amplified by the Airy pattern (Soummer et al. 2007).

In essence all the simulations corresponding to various distances and spectral types generate only the three effects mentioned above. The contrast curves will be steeper for brighter stars. To first-order, when the recorded flux ratio between two stars is say, two, the contrast curve will be steeper by a factor of 0.7.

Figure 11 also shows that contrast at small separations is again limited by the...
3.4.4 Star’s spectral type

Similarly to the previous section contrast is better for brighter stars (Figure 12).

3.4.5 Influence of M1 aberrations

To see the effect of higher M1 aberrations on the final contrast curve, we ran one simulation with M1 phase divided by two, and another where it was multiplied by three. The resulting rms wavefront values on M1 were 0.15μm and 0.9μm before XAO, and 4nm and 24nm after correction, respectively. Wavefront phase after AO (rms = 55nm) was dominated by post-AO aberrations, kept at the level of 30nm, and atmospheric residuals. Figure 13 shows that the effect of having very tight M1 specifications is negligible if instrumental errors remain. Both: “perfect” and APLC coronagraph’s contrast curves are not susceptible to varying M1 aberrations by almost one order of magnitude. Earth-like planets are still not detectable but super-Earths could be imaged with a less-chromatic coronagraph compared to the APLC.
3.4.6 Influence of post-AO aberrations

As can be seen in Figure 14, these aberrations - even though their magnitude is small compared to pre-AO M1 errors, determine the contrast curve. One could see that it is essential to keep post-AO aberrations below 60nm rms in order to have a chance of imaging a super-Earth. To detect an Earth-like planet these aberrations must be kept at the level of 10nm rms.

4 Concluding remarks

In this paper we reported on results of XAO simulations for the E-ELT. Simulations account for mirror segmentation, realistic segment aberrations, AO error budget, chromaticity of the coronagraph, and various sources of noise. The output of this extensive modelling suggests that with the combination of hardware and software described it will be possible to directly capture reflected light of super-Earth planets.
in the habitable zone of stars in the solar neighbourhood (d < 10pc) but only with a close to perfect coronagraph. The major limitation is imposed by the chromaticity of the coronagraph chosen for realistic simulations (APLC), post-AO aberrations being the next most important factor determining the contrast that can be achieved. This conclusion confirms the results of the simulations from the EPICS consortium (Verinaud 2010).

Imaging a true Earth twin within 10pc from the Sun would be feasible but will require keeping the instrumental post-AO aberrations at the level of 10nm rms, in addition to a perfect coronagraph (Figure 14). The E-ELT should open up a new discovery space when it comes to reflected light by old Jupiter-like planets and smaller giant gaseous planets. We have shown through our simulations that with a realistic coronagraph (APLC) and moderate integration times these planets could be detected up to the distances of 30pc if instrumental aberrations are kept at the level of around 30nm.

Acknowledgments

The authors would like to thank Bruno Femenia Castella, Christophe Verinaud, Patrice Martinez, Arkadiusz Swat, Babak Sedghi, Laurent Jolissaint and Isabelle Surdej.

References

Bonnet, H., 2009, E-ELT report E-SPE-ESO-313-0202
Bonnet, H., 2009, E-ELT report E-TRE-ESO-313-0387
Bonnet, H., 2009, E-ELT report E-TRE-ESO-313-0637
Brandl, B., et al., 2008, SPIE Proceedings, 7014, 70141N
Crossfield, I., Troy, M., 2007, Applied Optics, 46, 4533
Fusco, T., et al., 2006, Optics Express, 14, 7515
Gonte, F., 2009, E-ELT report ELT-TRE-ESO-04600-0099
Kasper, M., et al., 2008, SPIE Proceedings, 7015, 70151S
Müller, M., 2008, E-ELT report E-TRE-ESO-303-0273
Poyneer, L., Macintosh, B., 2006, Optics Express, 14, 7499
Tecza, M., Thatte, N., Clarke, F., Freeman, D., 2009, Science with the VLT in the ELT Era, 267
Troy, M., Crossfield, I., Chanan, G., Dumont, P., Green, J., Macintosh, B., 2006, SPIE Proceedings, 6272, 62722C
Verinaud, C., 2010, EPICS report E-TRE-LAO-556-0002