12th ESSENA School, Catania, 17-22 June 2024

Inferring stellar compositions and related open questions

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gaia

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Aim of this introductory lecture

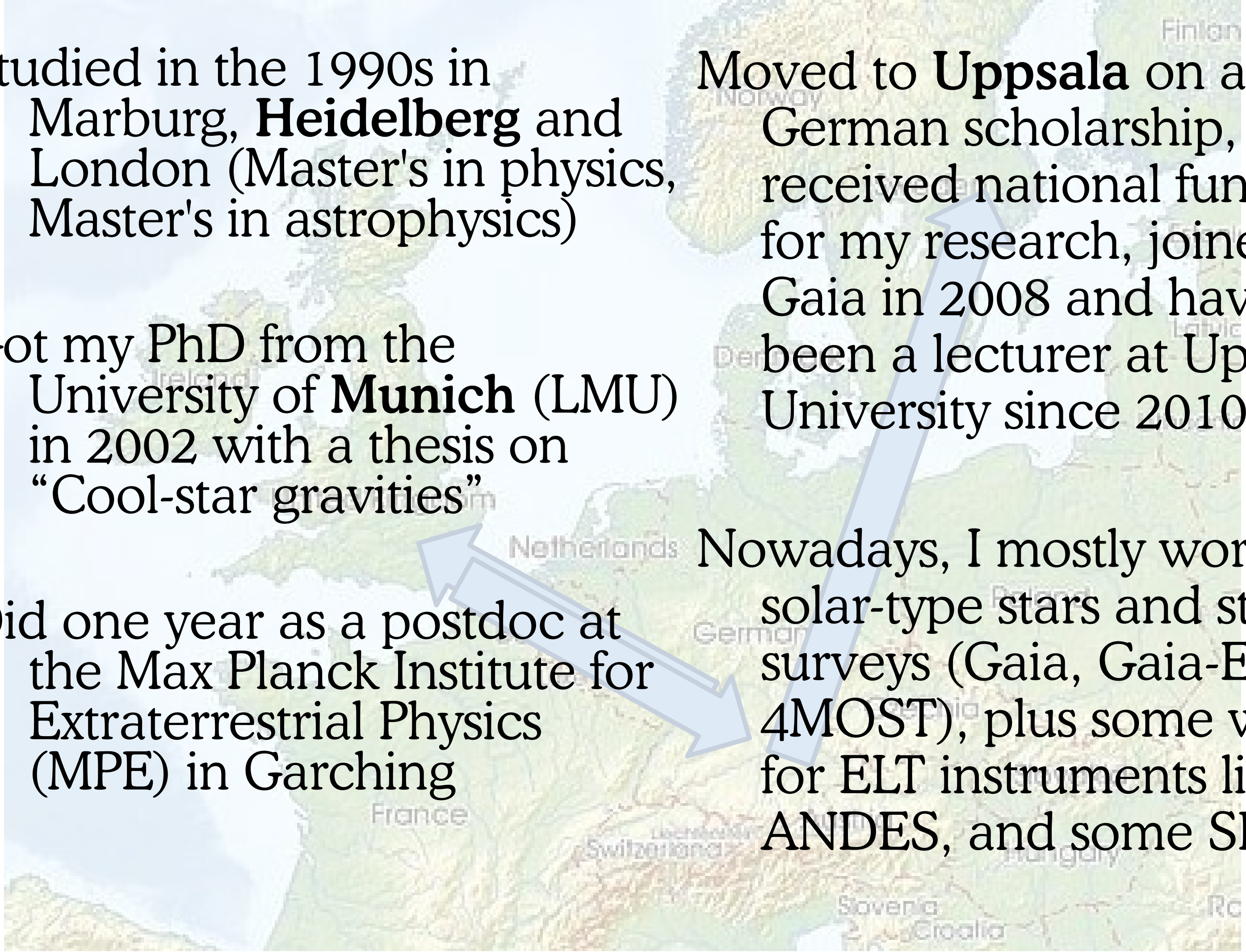
Get to know the "*bright side*"⭐, i.e. stellar spectroscopy and its assumptions, challenges and biases

Familiarize yourself with relevant astrophysical concepts and nomenclature

Get inspired to read on…

Other lectures later today/this week will give much more details on individual nuclear processes, astronomical messengers and astrophysical sources. And of course the nuclear part of the story…

My career in a nutshell



Studied in the 1990s in Marburg, **Heidelberg** and London (Master's in physics, Master's in astrophysics)

Got my PhD from the University of **Munich** (LMU) in 2002 with a thesis on “Cool-star gravities”

Did one year as a postdoc at the Max Planck Institute for Extraterrestrial Physics (MPE) in Garching

Moved to **Uppsala** on a German scholarship, received national funding for my research, joined Gaia in 2008 and have been a lecturer at Uppsala University since 2010

Nowadays, I mostly work on solar-type stars and stellar surveys (Gaia, Gaia-ESO, 4MOST), plus some work for ELT instruments like ANDES, and some SETI…

 # stars

10s

…

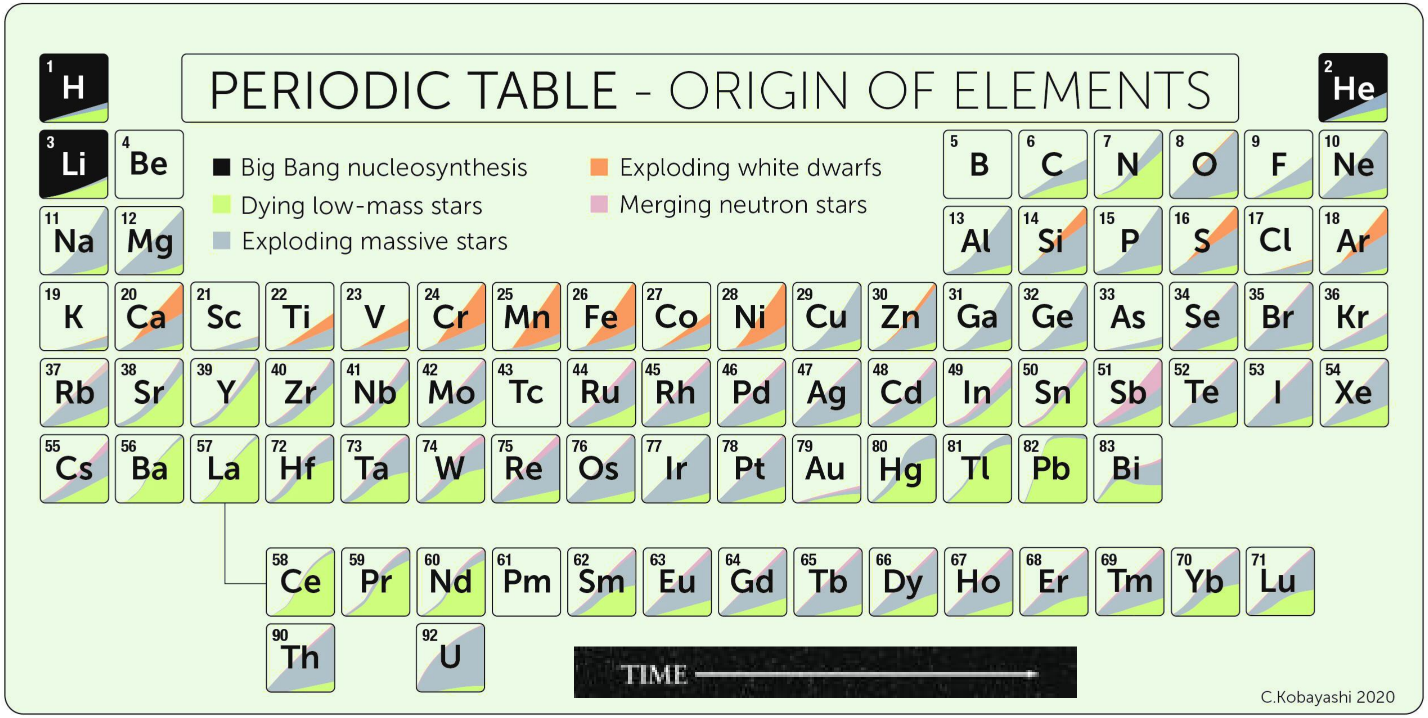
1000s

…

$5 \cdot 10^8$

Nuclear astrophysics 101

Observing the ~90 stable elements, inferring their abundances and unravelling their origin.



Observing isotopes in stellar spectra is only possible when the isotopic shift is large: this is the case in the lightest elements (e.g. Li) and when the element of interest is visible in molecular form (e.g. C_2, MgH), mostly in cool (giant) stars. Hfs splitting of strong lines sometimes works (Gallagher *et al*. 2015).

Stars: typical figures

The Sun (G2 V)

$M = 2 \times 10^{33}\ \mathrm{g} = \mathrm{M}_\odot$

$R = 7 \times 10^{10}\ \mathrm{cm} = \mathrm{R}_\odot$

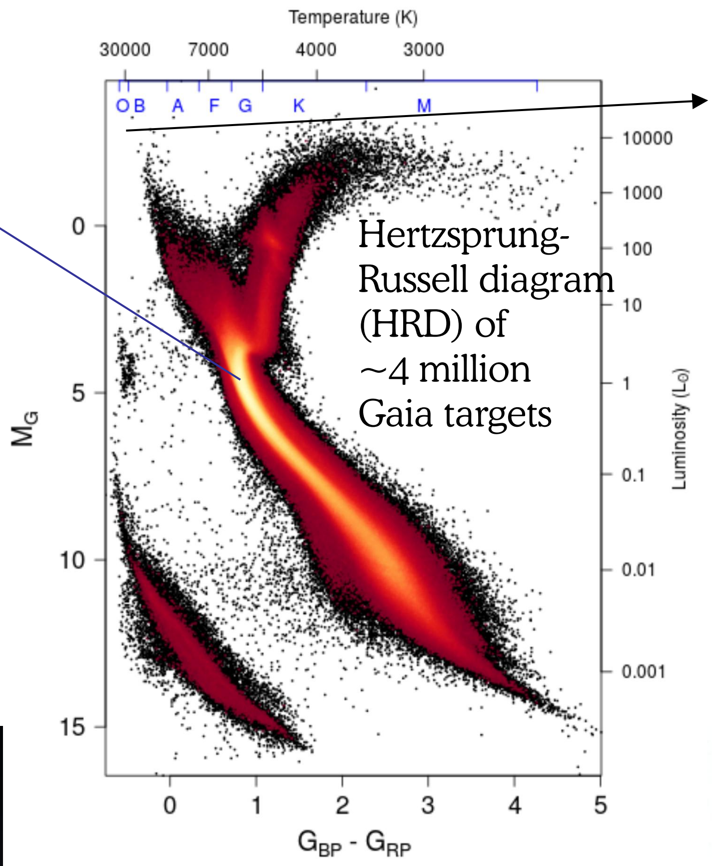
$L = 4 \times 10^{33}\ \mathrm{erg/s} = \mathrm{L}_\odot$

photosphere:

$\Delta R \approx 200\ \mathrm{km} < 10^{-3}\ \mathrm{R}_\odot$

$n \approx 10^{15}\ \mathrm{cm}^{-3}$

$T_{\mathrm{surface}} \approx 6000\ \mathrm{K}$



Hertzsprung-Russell diagram (HRD) of ~4 million Gaia targets

an O star

$M \approx 50\ \mathrm{M}_\odot$

$R \approx 20\ \mathrm{R}_\odot$

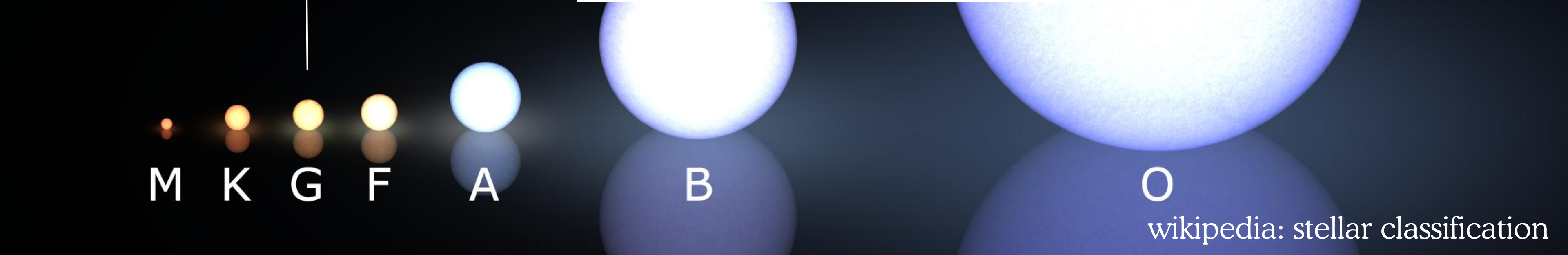
$L \approx 10^6\ \mathrm{L}_\odot (\propto M^3)$

photosphere:

$\Delta R \approx 0.1\ \mathrm{R}_\odot$

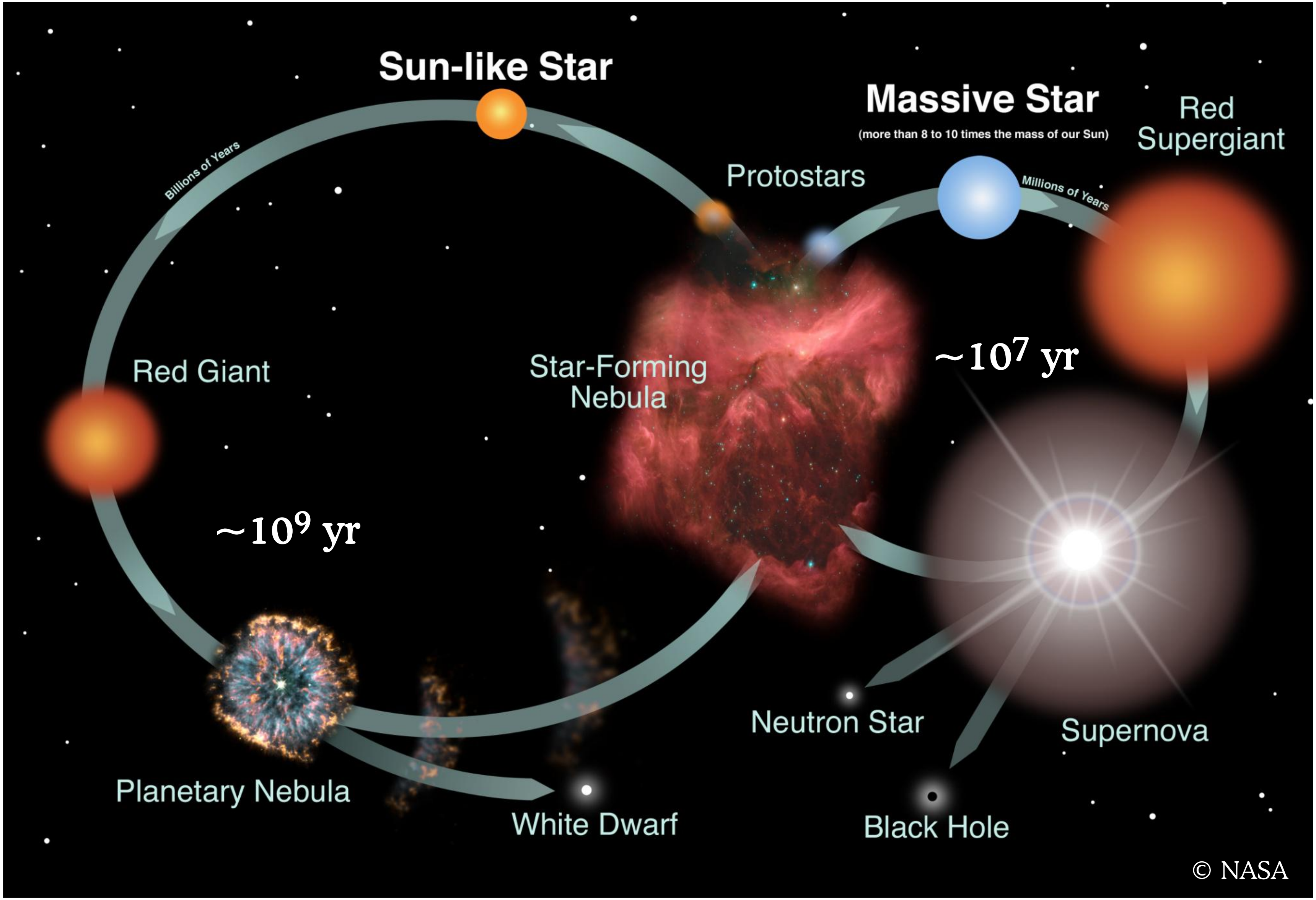
$n \approx 10^{14}\ \mathrm{cm}^{-3}$

$T_{\mathrm{surface}} \approx 40\,000\ \mathrm{K}$



wikipedia: stellar classification

Cosmic Matter Cycle



© NASA

Different stars and their USPs

OBA (so-called "early-type"): short-lived (10s of millions of years), thus probe current star formation. Progenitors of SN type II (e.g. SN 1987A). Can be observed in other galaxies as individual stars (literally megasuns!).
Relatively clean spectra: H, noble gases, CNO (often altered by stellar evolution), α-capture and iron-peak elements

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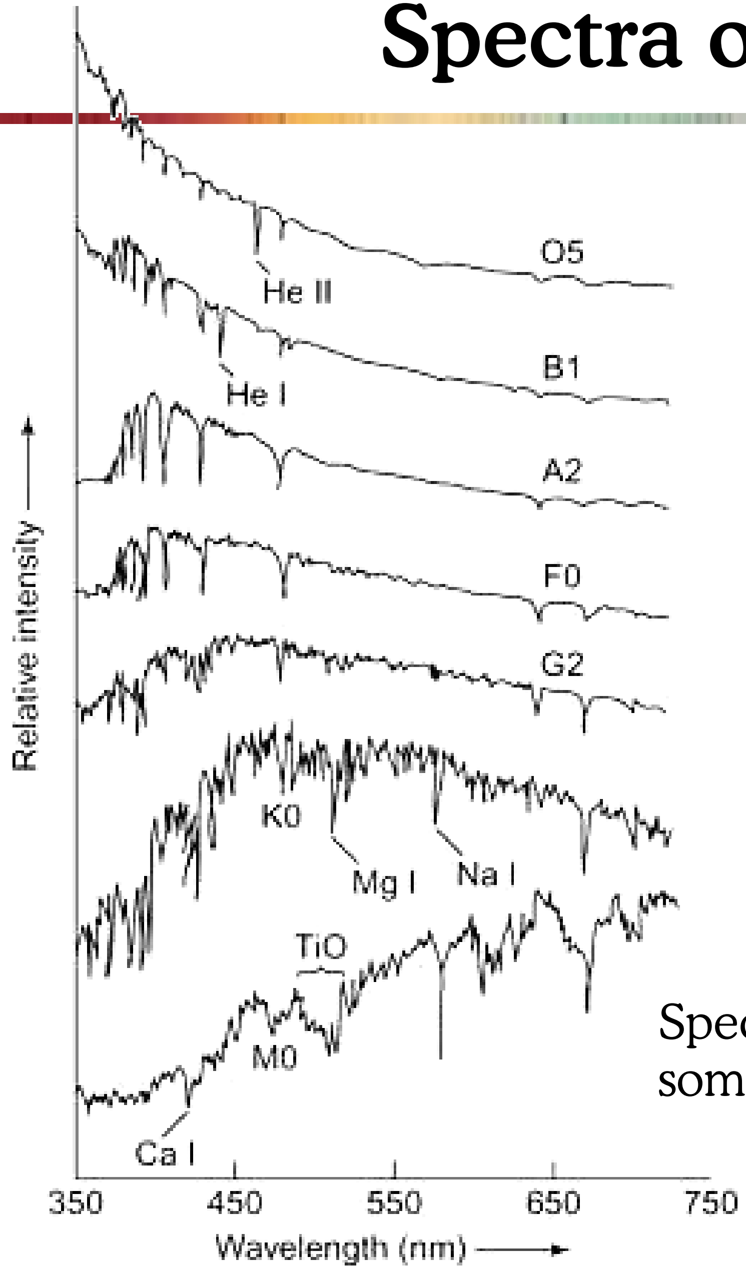
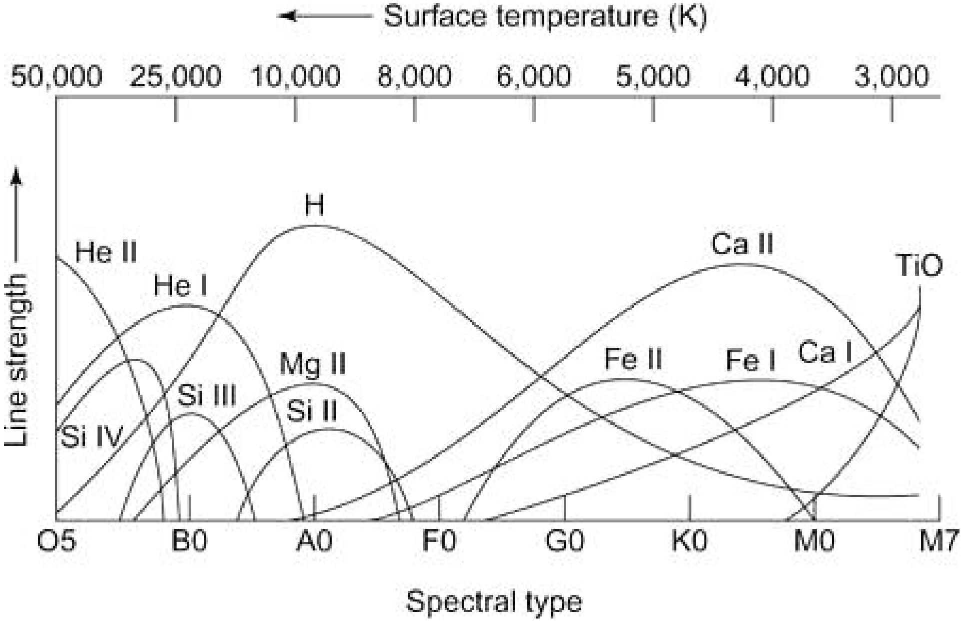
Relatively clean spectra: H, noble gases, CNO (often altered by stellar evolution), α-capture and iron-peak elements

FGKM (so-called "late-type"): long-lived (Sun's total livetime until it becomes a white dwarf ~10 Gyr), thus probe the whole chemical-evolution timeline of the MW. Not easy to observe beyond the Magellanic Clouds (work-around: integrated-light studies of e.g. clusters of stars).
Rich and complex spectra: H, many atomic and molecular lines of **all sorts of metals**, but no noble gases; isotopes in favourable cases

Spectra of different stars

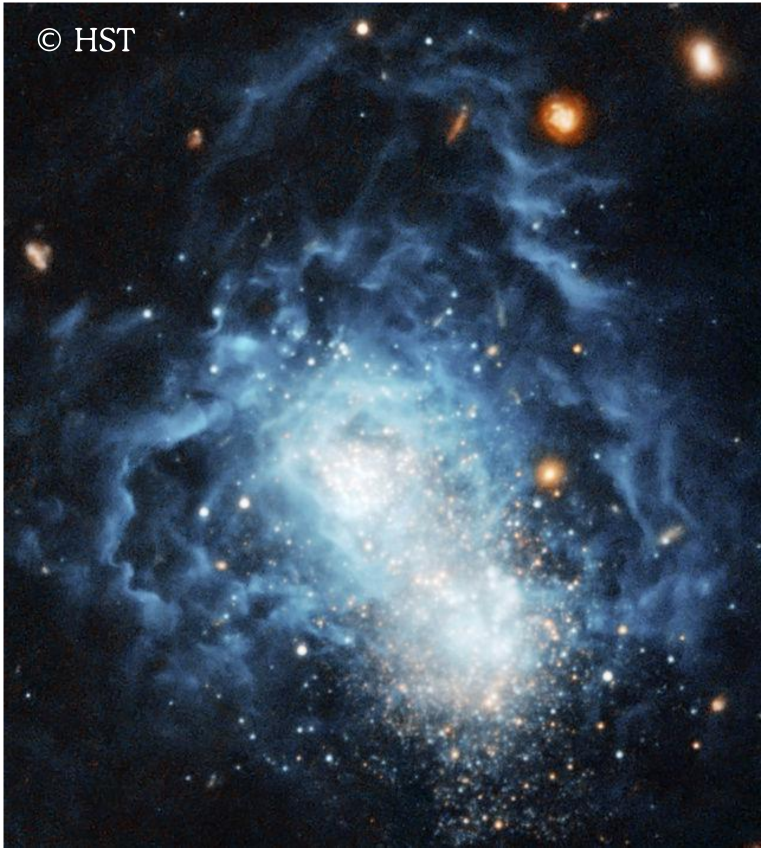




Spectra of low-mass stars (KM) allow to probe some isotopic ratios, e.g. ^{24,25,26}Mg (Yong *et al.* 2003)

© CliffsNotes: Spectral Types

Objects dominated by OB- vs GK-type stars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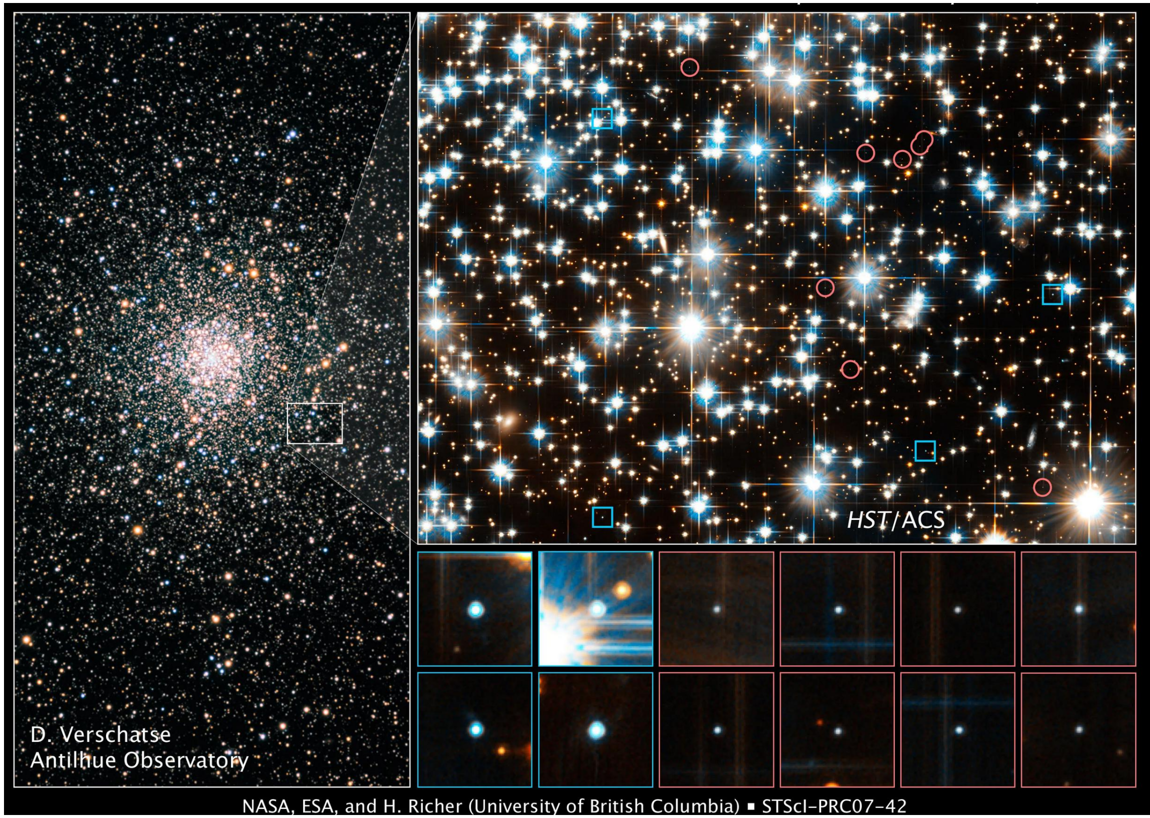
I Zwicky 18, a star-bursting
metal-poor galaxy
(a unusually late bloomer in the
local universe)

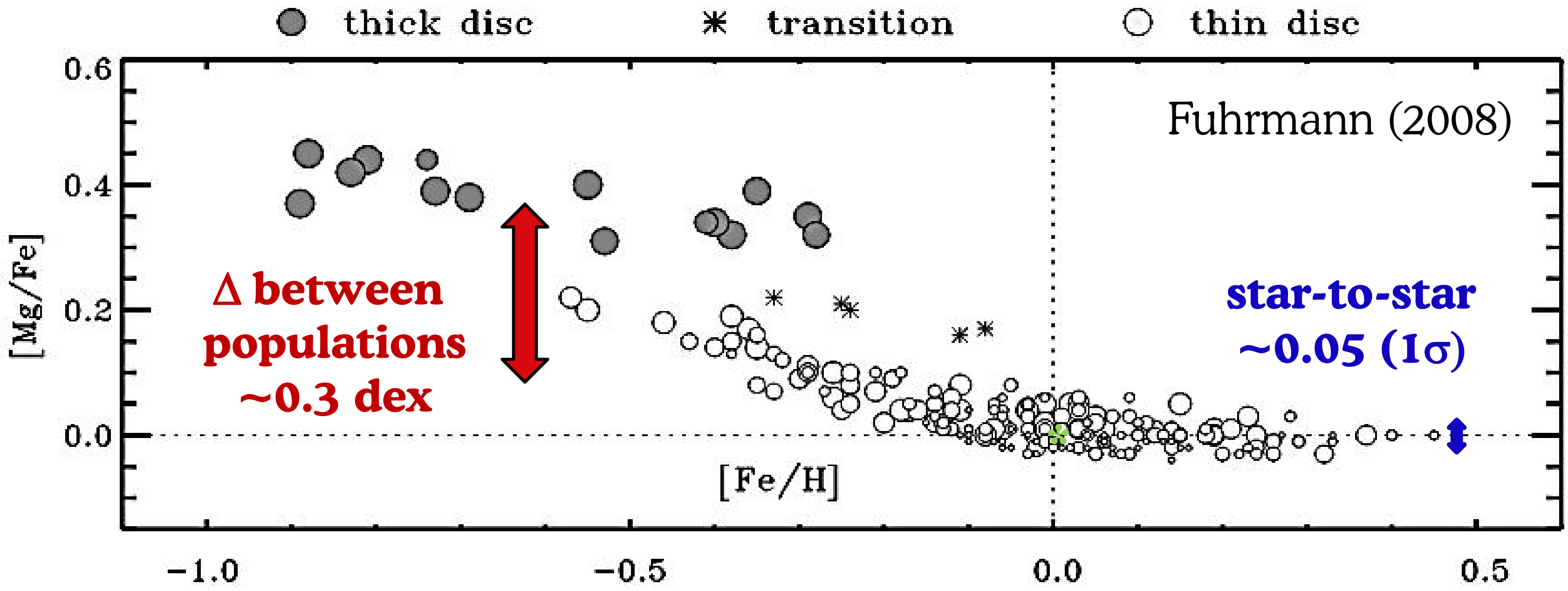
NGC 6397, an old metal-poor
globular cluster
(in thus cluster, the white-dwarf cooling
sequence has been followed to its end)



D. Verschatse
Antilhue Observatory

NASA, ESA, and H. Richer (University of British Columbia) ▪ STScI-PRC07-42

Achievable abundance accuracies



Δ between populations ~0.3 dex

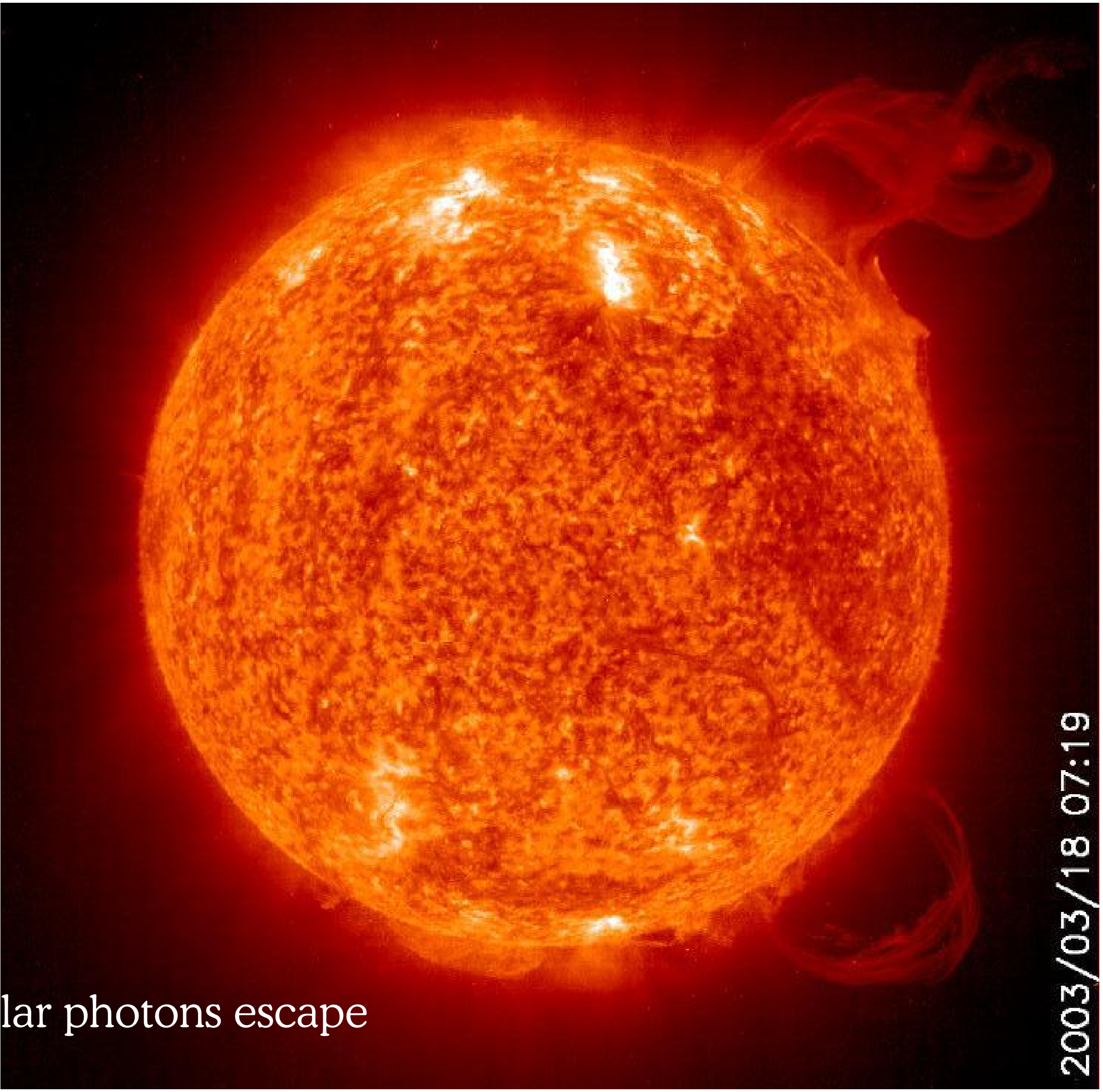
star-to-star ~0.05 (1σ)

The solar neighbourhood (25 pc) as seen in F- & G-type stars.

The Sun (*) is a normal, albeit fairly high-mass, thin-disk star.

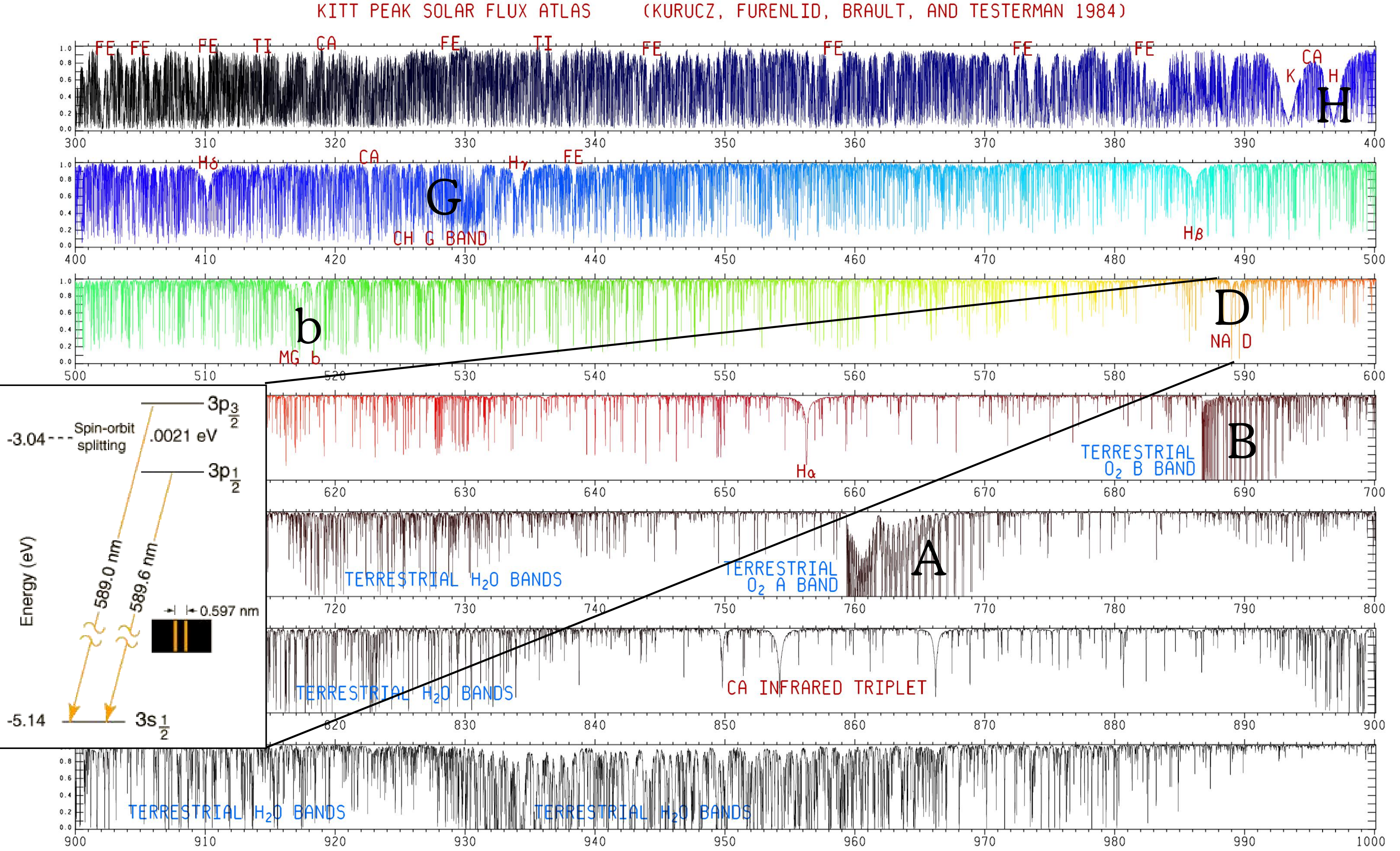
(No bulge and very few halo stars (not shown here) within 25 pc of the Sun.)

Stellar atmospheres…



… where stellar photons escape

The solar spectrum

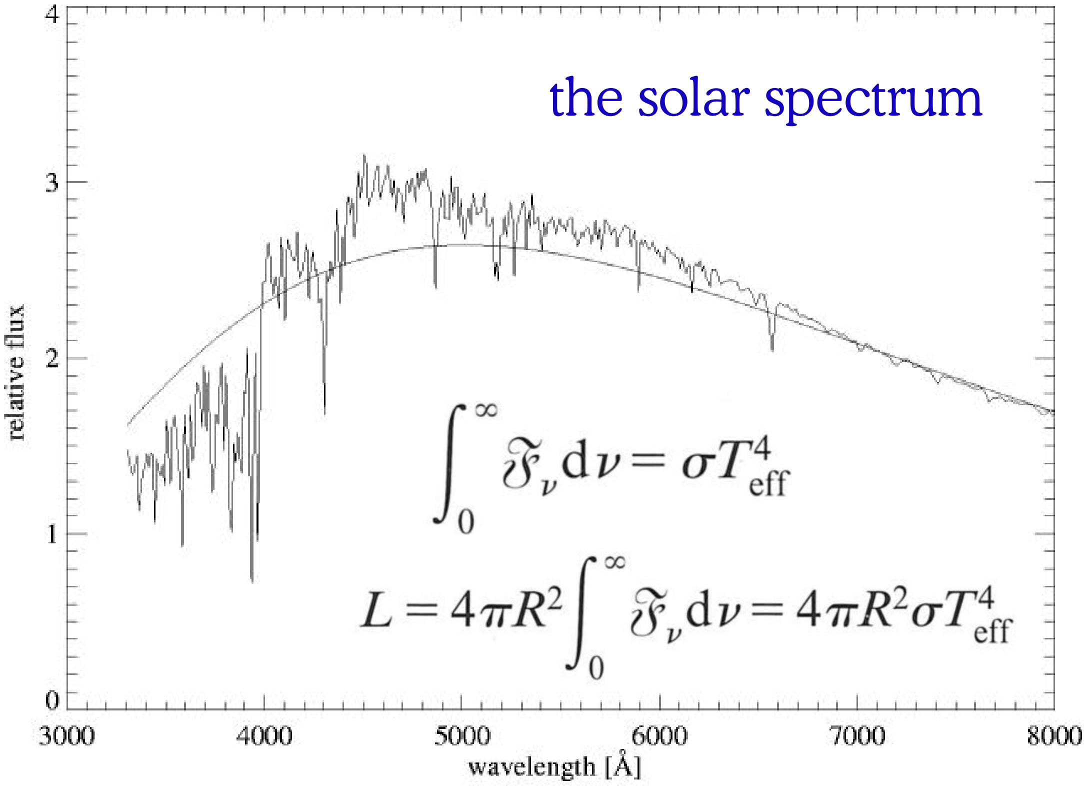


The spectra of stars

Luckily, stars (and other celestial bodies) are not in thermodynamic equilibrium (TE) and do not shine like blackbodies.

(Astronomy would be the dullest of all sciences!)

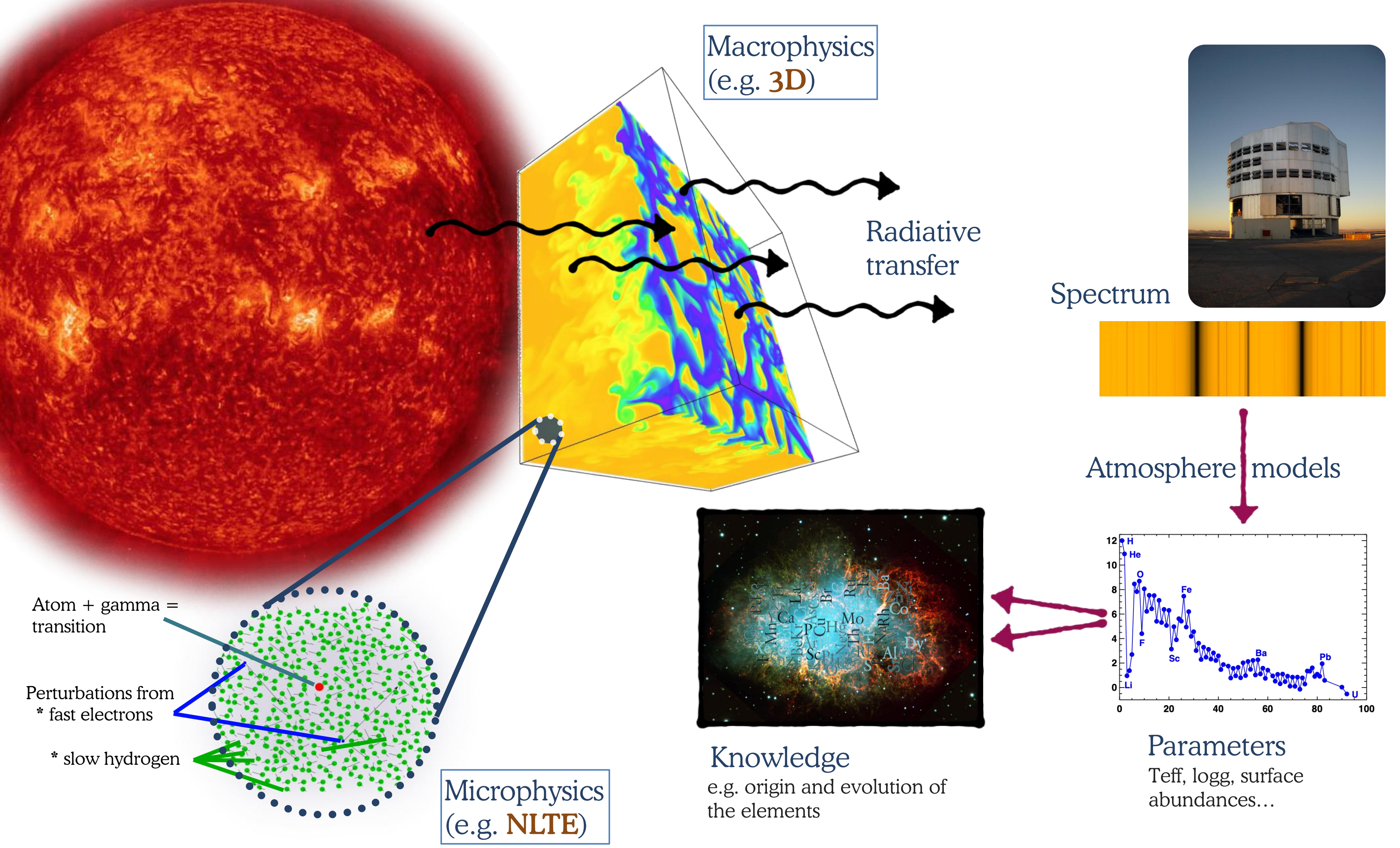
In constrast to B_ν, I_ν depends on plasma properties and the viewing angle. One cannot use TE to describe starlight.



$$\int_0^\infty \mathcal{F}_\nu \, \mathrm{d}\nu = \sigma T_{\mathrm{eff}}^4$$

$$L = 4\pi R^2 \int_0^\infty \mathcal{F}_\nu \, \mathrm{d}\nu = 4\pi R^2 \sigma T_{\mathrm{eff}}^4$$

Where micro- and macrophysics meet…



Macrophysics
(e.g. 3D)

Radiative
transfer

Spectrum

Atmosphere models

Parameters
Teff, logg, surface
abundances…

Knowledge
e.g. origin and evolution of
the elements

Atom + gamma =
transition

Perturbations from
* fast electrons

* slow hydrogen

Microphysics
(e.g. NLTE)

adapted from A. Amarsi

Model atmospheres 101

The cool tenuous layer of stars we call stellar atmospheres absorb and emit photons. We model this by solving the **radiative transfer equation**:

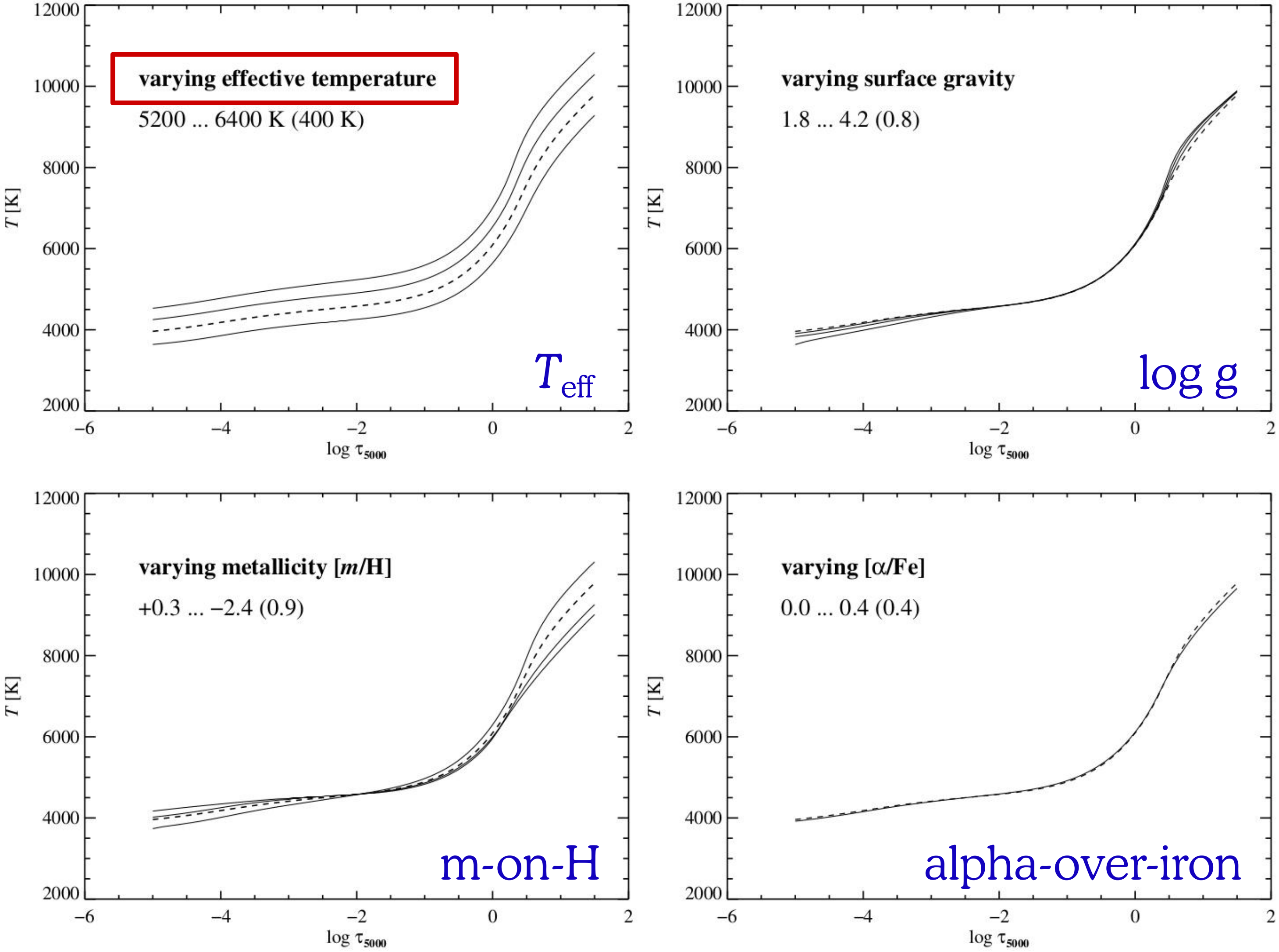
$$\frac{\mathrm{d}I_\nu}{\mathrm{d}\tau_\nu} = -I_\nu + S_\nu$$

A central goal is to find the temperature gradient that establishes itself in order to conserve the total flux (flux-constancy models).

A model atmosphere tabulates two or more thermodynamic variables as a function of (optical) depth. It is one of the main inputs to calculations of synthetic stellar spectra.

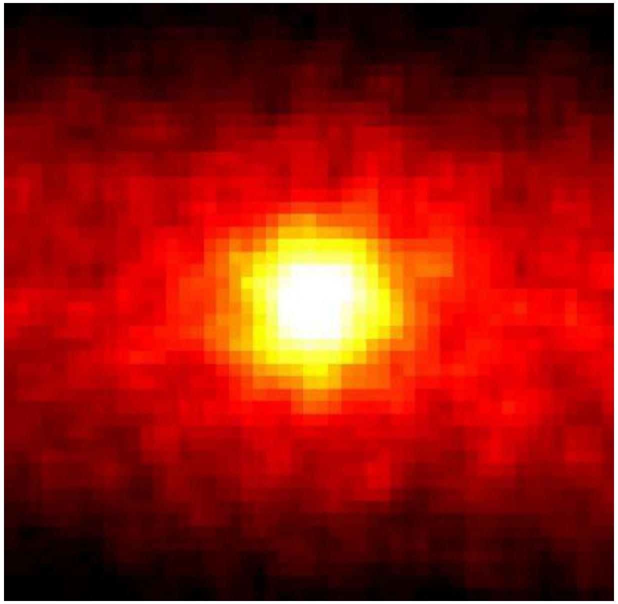
In order to derive reasonably realistic models, one needs to consider 100s of 1000s of opacity sources from atoms and molecules. See e.g. Gustafsson *et al.* (2008) for one set of models ("MARCS").

Grids of model atmospheres: *T* vs τ



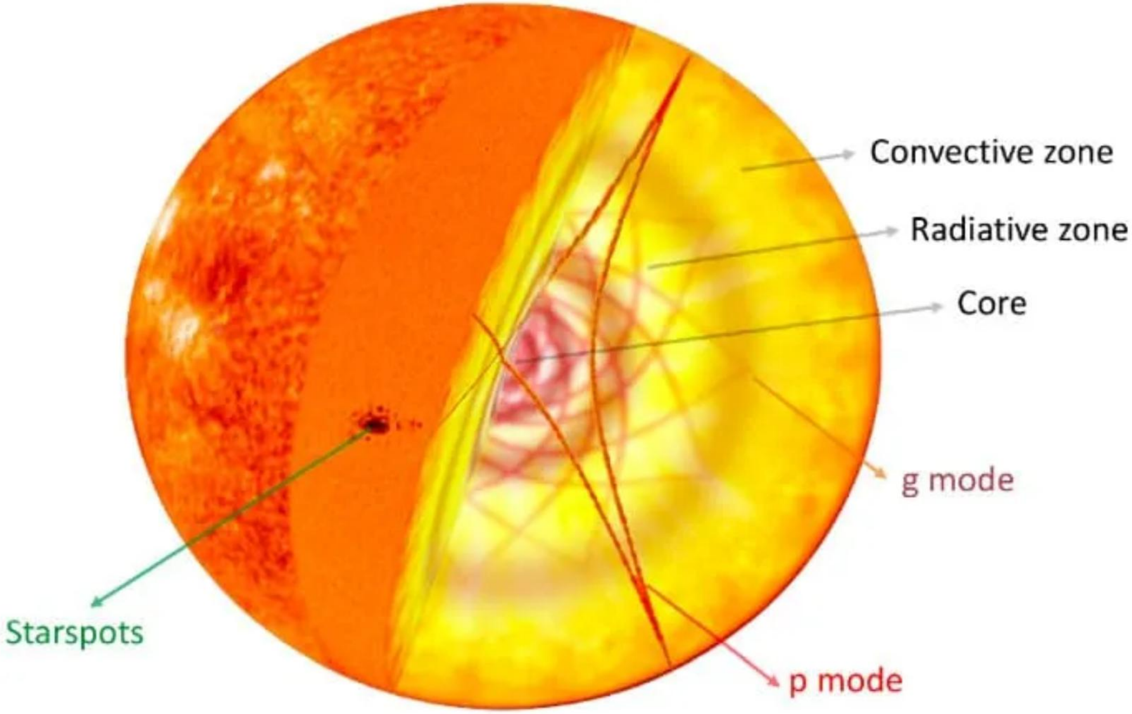
⇦ towards surface towards interior ⇨

The interior of the Sun



Neutrinos from the fusion reactions in the Sun's core as seen by **Superkamiokande** (Japan) looking right through Earth!

Probing acoustic (p-mode) waves through helioseismology also lets us study the Sun's interior structure. These are excited by near-surface convection.



Asteroseismology of solar-type stars. Image by link.springer.com

Line formation 101

The flux coming from subphotospheric layers (where the mean free path of photons is small) is Planckian, i.e. a blackbody.

Based on the run of temperature and pressure as a function of (optical) depth, you can study how the electronic transitions in atoms and molecules lead to absorption lines at characteristic wavelengths.

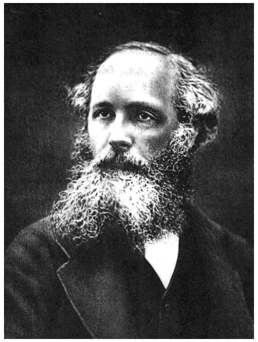
The strength of a spectral line is proportional to the ratio of the line vs the continuous absorption coefficient. It also depends on the gradient of the source function throughout the depths of line formation.

In the classical LTE approximation, the formation of every lines is taken as an isolated process following *local* equilibrium (Saha-Boltzmann) statistics.

TE statistics

Particle velocities are assumed to be Maxwellian:

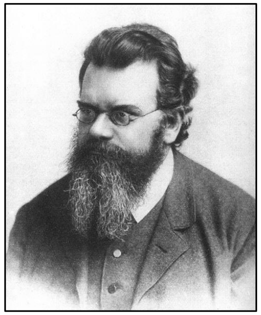
$$\frac{n(v)}{n_{\mathrm{tot}}}\mathrm{d}v = (\frac{m}{2\pi kT})^{\frac{3}{2}} e^{-\frac{mv^2}{2kT}}\,\mathrm{d}v\,\frac{,^{2}}{T}\,\mathrm{d}v$$



Excitation follows the Boltzmann distribution:

$$\frac{n_u}{n_{\mathrm{tot}}} = \frac{g_u}{u(T)}e^{-\frac{\chi_u}{kT}}$$

χ: excitation energy



Ionization can be computed via the Saha equation:

$$\frac{n_{\mathrm{II}}}{n_{\mathrm{I}}}P_e = \frac{(2\pi m_e)^{3/2}kT^{5/2}}{h^3}\,\frac{2u_{\mathrm{II}}(T)}{u_{\mathrm{I}}(T)}e^{-\frac{I}{kT}}$$

I: ionization energy

In **local thermodynamic equilibrium** (LTE) often used in line-formation calculations, these are applied ***locally***.

Line strength dependencies

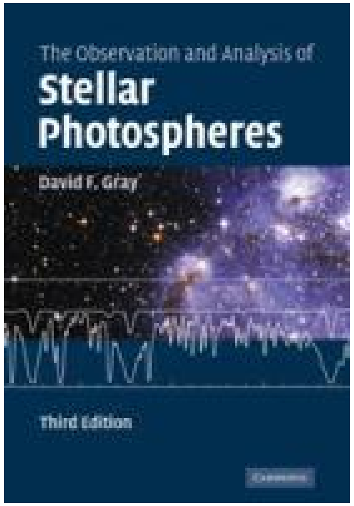
governed by the line
formation (LTE vs. NLTE)

$$\frac{\mathfrak{F}_c - \mathfrak{F}_\nu}{\mathfrak{F}_c} \approx \tau_1 \frac{\mathrm{d}\ln S_\nu}{\mathrm{d}\tau_c}\bigg|_{\tau_1} \frac{\ell_\nu}{\kappa_\nu}$$

governed by the model
atmosphere (1D vs. 3D)

The left-hand side of the above equation is a measure of the monochromatic line flux (subscript *ν*) eaten out of the continuum (subscript c). Integrate this and you get the line strength.
The right-hand side shows important dependencies:
* gradient of the source function with optical depth and
* the ratio of line to continuous absorption coefficient.

With this, one can basically understand how different lines (transitions) behave. See Gray's excellent book.



Opacities

Continuous opacity

Caused by *bf* or *ff* transitions

In the optical and near-IR of cool stars, H^- ($I = 0.75\,\mathrm{eV}$) dominates:

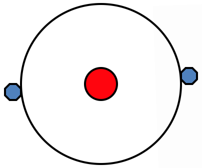
$$\kappa_\nu(\mathrm{H}^-_{\mathrm{bf}}) = \text{const.}\ T^{-5/2} P_e \exp(0.75/kT)$$

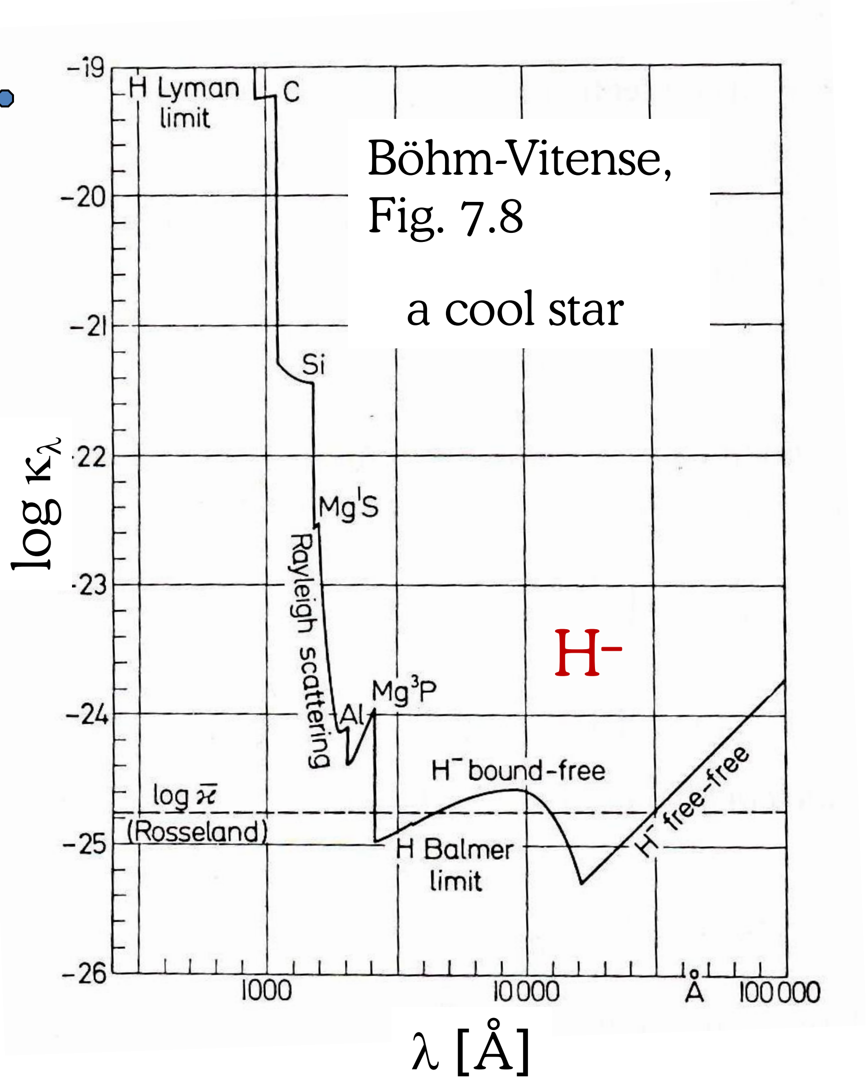
NB: There is only 1 H^- per 10^8 H atoms in the Solar photosphere.

Line opacity *(all the lines you see!)*

Caused by *bb* transitions

Need to know log *gf*, damping and assume an abundance





Böhm-Vitense, Fig. 7.8

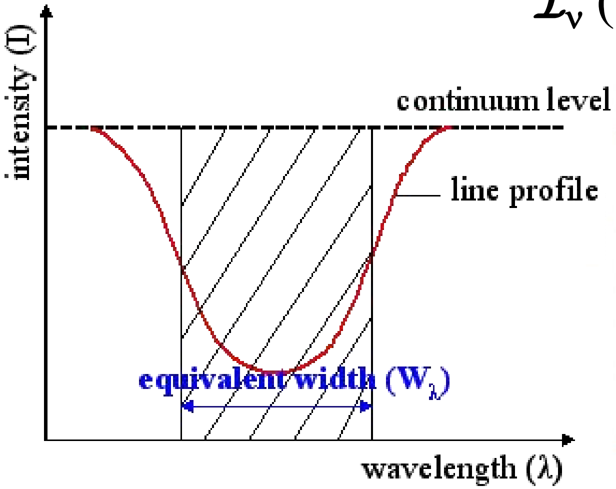
a cool star

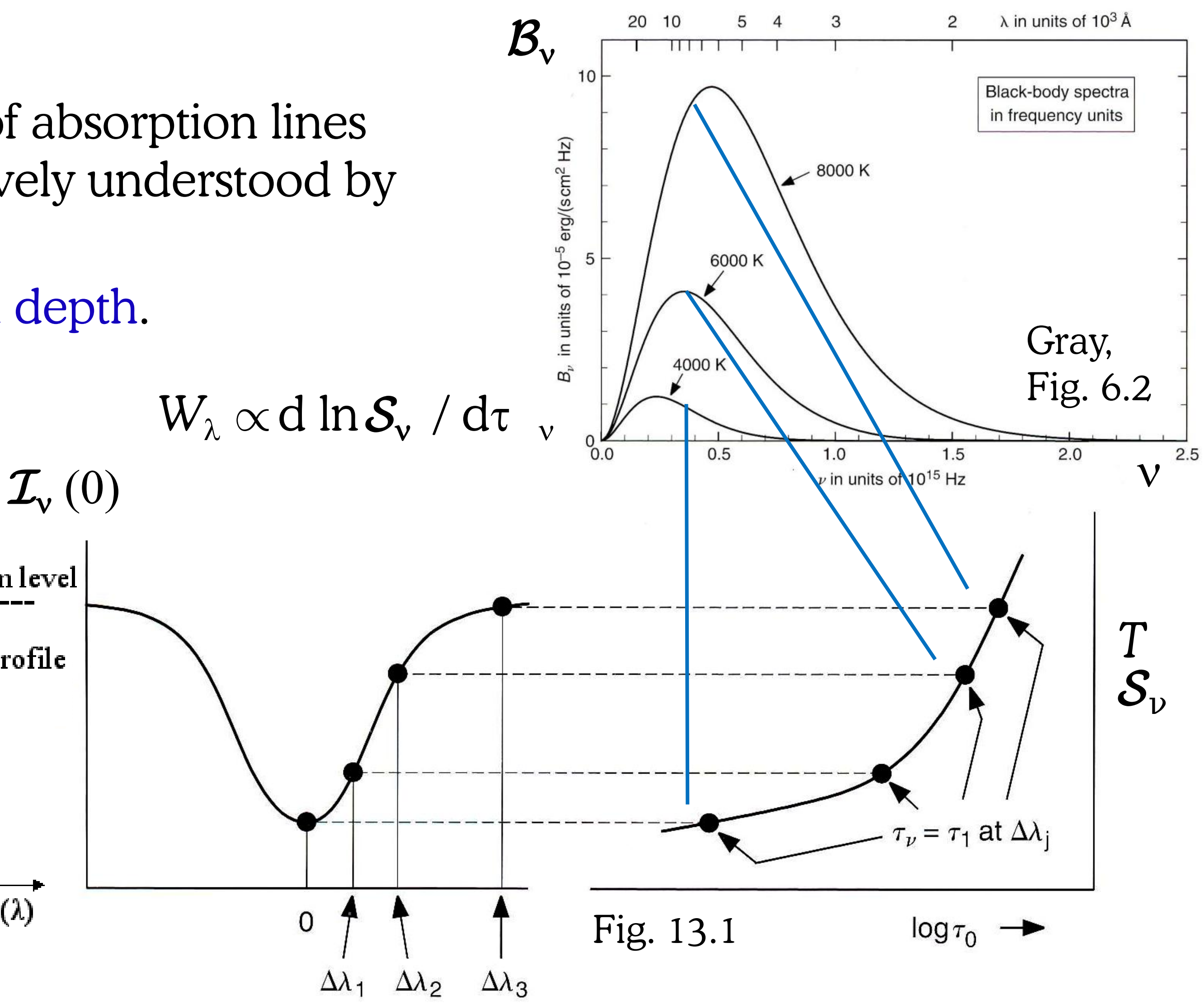
How spectral lines originate

The formation of absorption lines can be qualitatively understood by studying how

$\mathcal{S}_\nu$ changes with depth.

$$W_\lambda \propto d \ln \mathcal{S}_\nu \, / \, d\tau_\nu$$





Gray, Fig. 6.2

Fig. 13.1

Spectral lines as a function of abundance

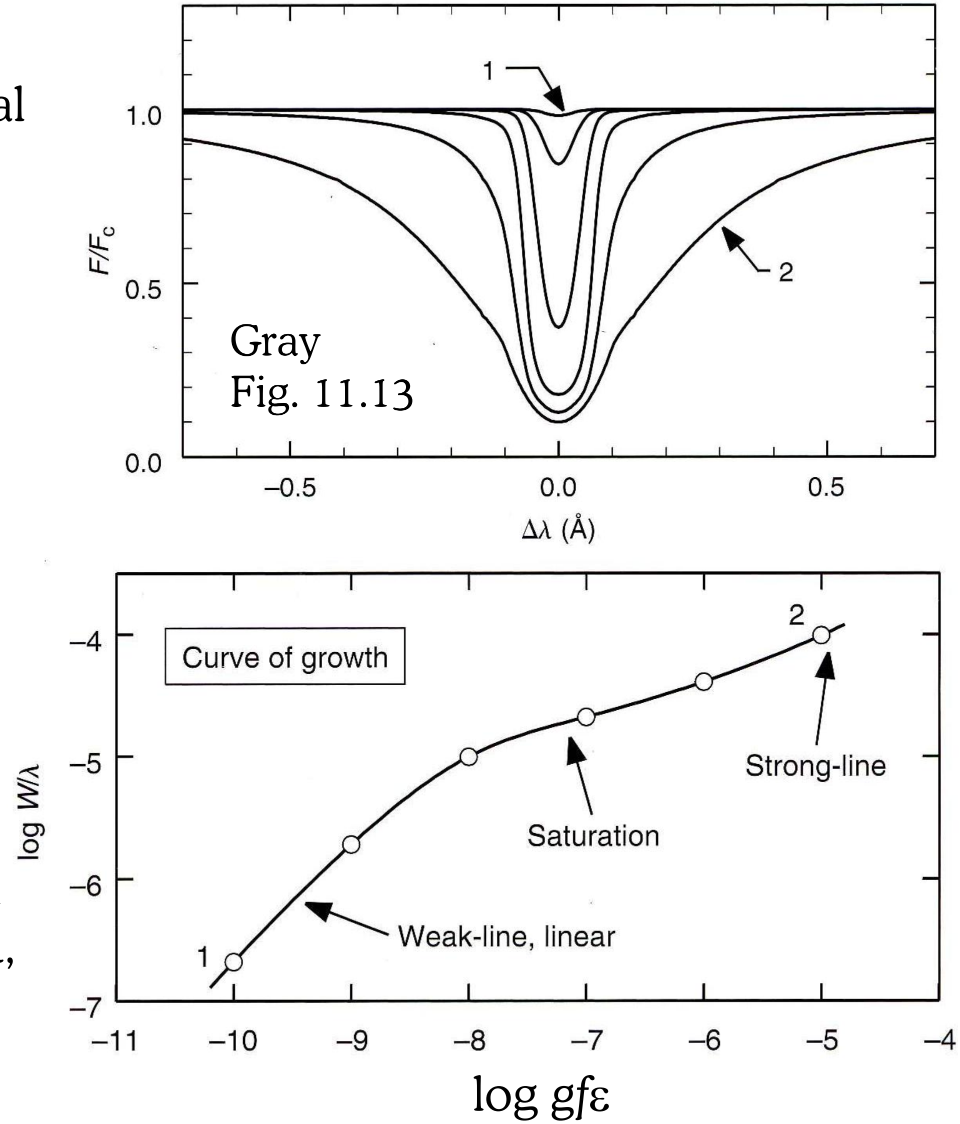
Starting from low log ε (low log *gf*), the line strength is directly proportional to *f* and n_X:

$$W_\lambda \propto gf\, n_X$$

When the line centre becomes optically thick, the line begins to saturate. The dependence on abundance lessens. Only when damping wings develop, the line can grow again in a more rapid fashion:

$$W_\lambda \propto \mathrm{sqrt}(\, gf\, n_X\,)$$

Weak lines are thus best suited to derive the stellar composition, given that they are well-observed (high SNR, little or known blending)



Gray
Fig. 11.13

Understanding abundance nomenclature

Mass fractions: let X, Y, Z denote the mass-weighted abundances of H, He and all other elements ("metals"), respectively, normalized to unity ($X + Y + Z = 1$).

example: $X = 0.744$, $Y = 0.242$, $Z = 0.014$ for the *present* Sun

The 12 scale: $\log \varepsilon(\mathrm{X}) = \log (n_\mathrm{X} / n_\mathrm{H}) + 12$

example: $\log \varepsilon(\mathrm{O})_\odot \approx 8.7$ dex, i.e., oxygen is 2000 times less abundant than H in the Sun (the exact value is still debated!)

($\log \varepsilon(\mathrm{H}) \equiv 12$)

Square-bracket scale: $[\mathrm{X/H}] = \log (n_\mathrm{X} / n_\mathrm{H})_\star - \log (n_\mathrm{X} / n_\mathrm{H})_\odot$

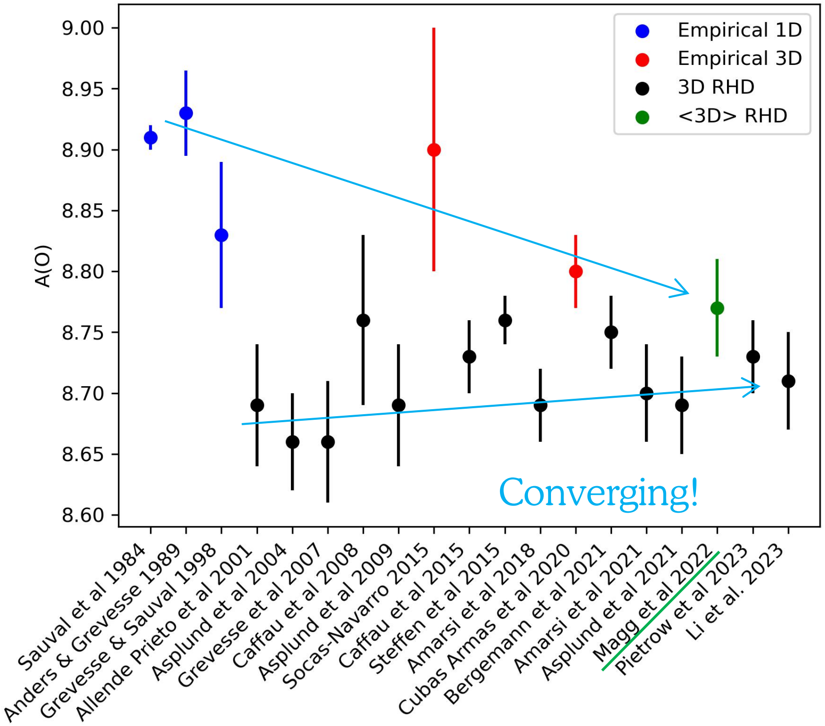
example: $[\mathrm{Fe/H}]_\mathrm{HE0107-5240} = -5.3$ dex, i.e., this star has an iron abundance (metallicity) a factor of 200 000 below the Sun (Christlieb *et al.* 2002)

($[\mathrm{X/H}]_\odot \equiv 0$)

Oxygen, the most abundant metal

The solar abundance of oxygen has come down very significantly due to better modelling of the different indicators.

The authorative reference for the solar composition is **Asplund, Amarsi & Grevesse (2021).**



Lind & Amarsi (2024)

Indicator	Non-LTE				LTE			
	3D	⟨3D⟩	MARCS	HM	3D	⟨3D⟩	MARCS	HM
[O I]	8.703	8.688	8.664	8.705	8.703	8.688	8.664	8.705
O I	8.686	8.682	8.657	8.678	8.805	8.779	8.732	8.789
OH (dv = 0)	–	–	–	–	8.689	8.792	8.782	8.837
OH (dv = 1)	–	–	–	–	8.707	8.800	8.729	8.883
OH (dv = 2)	–	–	–	–	8.690	8.756	8.673	8.845

Important (questionable) approximations

Classical models rest on a number of assumptions that for certain classes of stars have been shown to be problematic for quantitative results:

1. Plane-parallel atmospheres
2. Hydrostatic equilibrium with a local theory for convective energy transport
3. Line formation in Local Thermodynamic Equilibrium (LTE)
4. A well-mixed atmosphere representing the bulk stellar material (in part a stellar-structure question)
5. No chromospheres, no coronae, no magnetic fields, no rotation, no mass loss

LTE *vs.* NLTE

Occupation, excitation & ionization are assumed to be local properties

⇨ **Saha-Boltzmann statistics**

Assuming the *T*-*P*-τ relation to be known, all you need to to calculate a line strength is

(a) the level energies and statistical weights involved
(b) the transition probability
(c) broadening mechanisms (microturbulence, van-der-Waals damping)

Photons carry non-local information!

Occupation, excitation & ionization depend on the microphysics (radiation field, collisions etc.)

One needs to know (and master!) a whole lot of atomic physics.

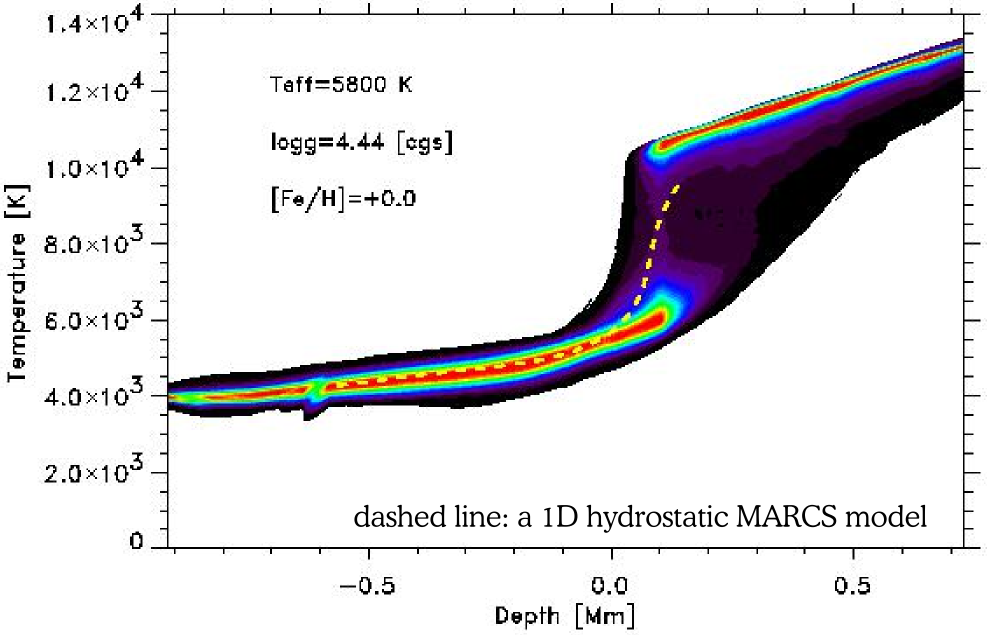
One also needs to solve the involved numerical problem of radiative transfer plus **rate equations**:

$$n_i \sum_{j \neq i} (R_{ij} + C_{ij}) = \sum_{j \neq i} n_j \, (R_{ji} + C_{ji})$$

While LTE may be an acceptable approximation for a cool-star photosphere on the whole, it can be very wrong for specific lines.

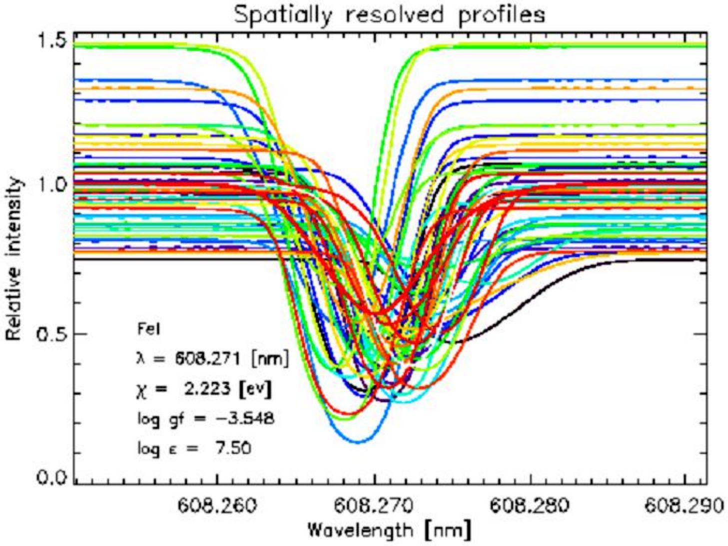
RHD model atmospheres

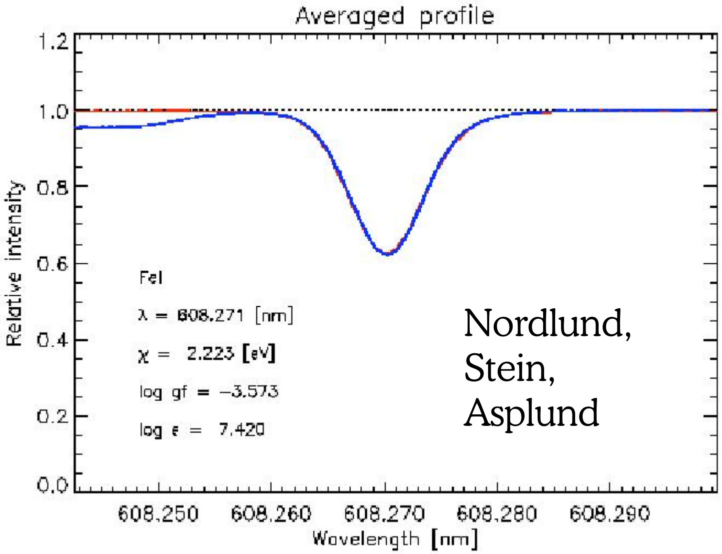
… all of the above in radiation-hydrodynamic (**3D**) models!



dashed line: a 1D hydrostatic MARCS model

Collet, PhD thesis, UU (2006)

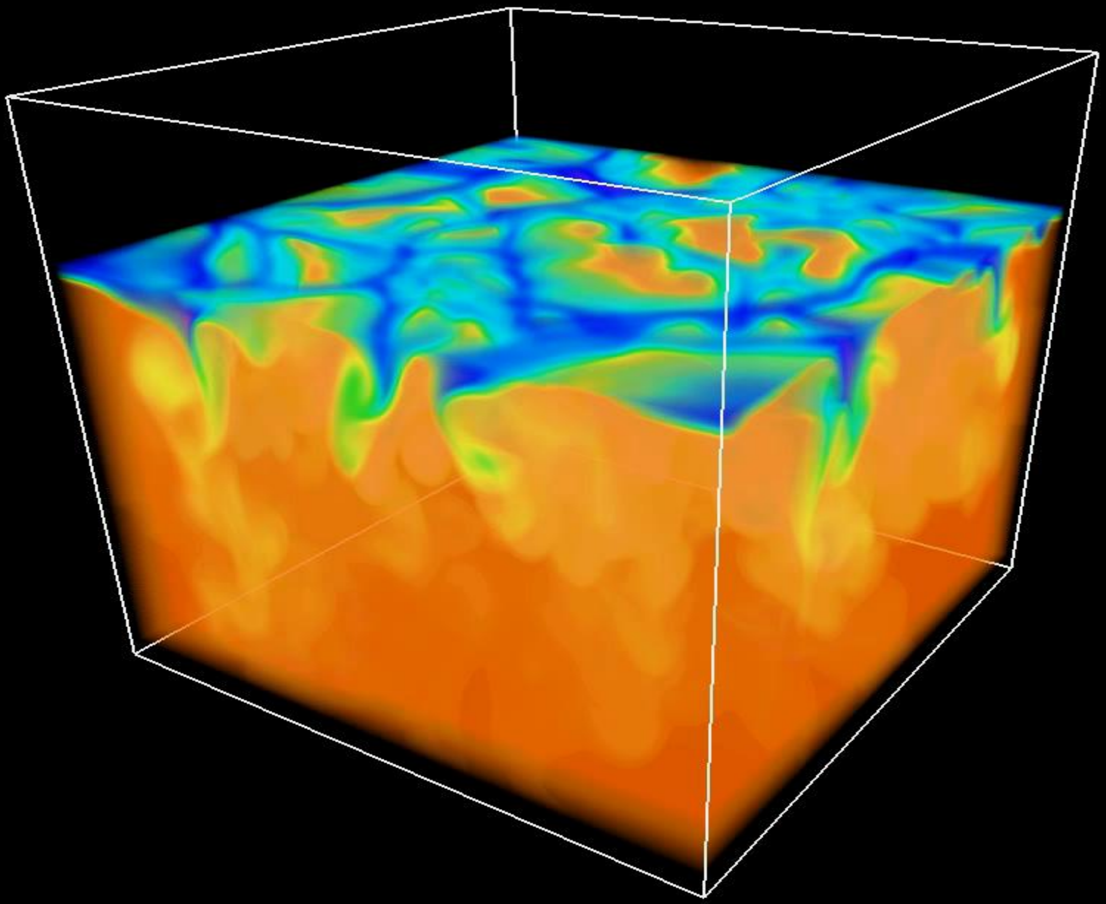




Nordlund,
Stein,
Asplund

At solar metallicity, 3D atmospheres resemble their 1D counterparts well, apart from the *T* fluctuations. But radiative transfer is not linear in *T*.

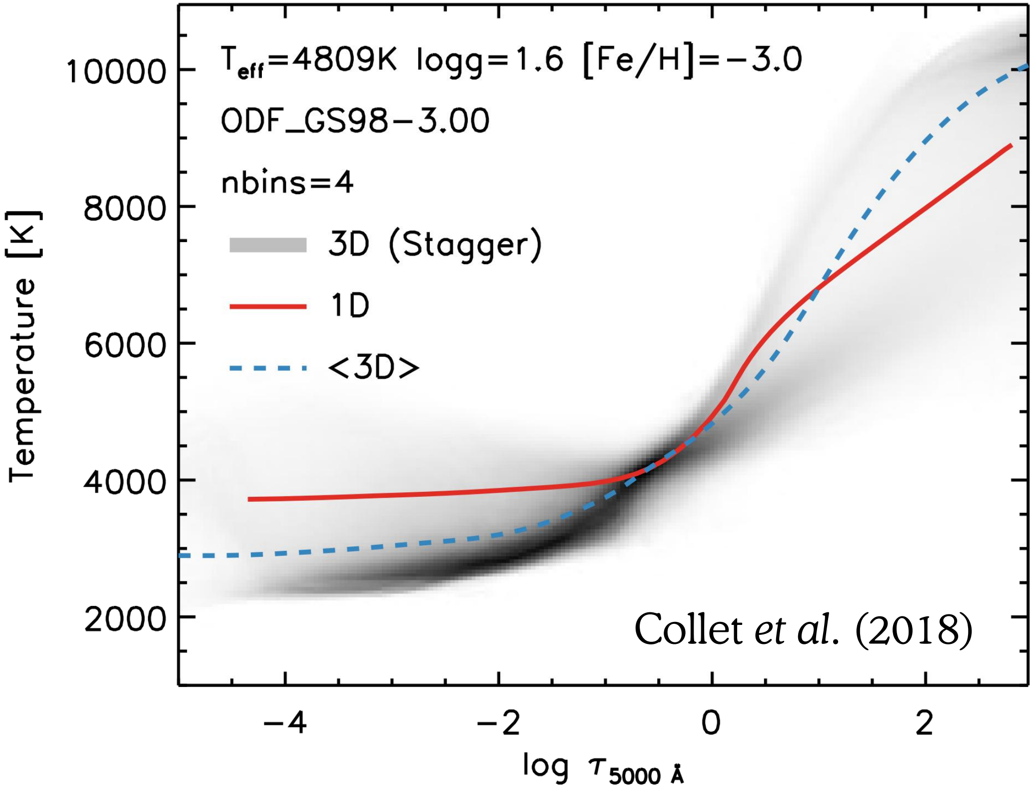
State-of-the-art RHD simulations



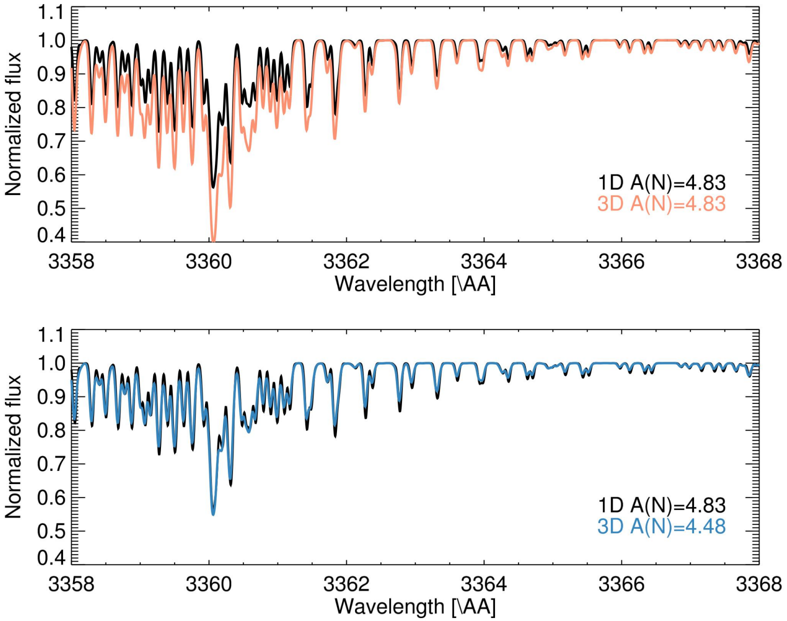
Highly realistic,
as analyses of
solar data has
shown.

Entropy fluctuations in STAGGER-code simulations, computed by Remo Collet

3D atmospheres for a metal-poor giant



Collet *et al.* (2018)



A 0.35 dex (55%) drop in inferred N abundance!

3D atmospheres are significantly cooler at the surface (due to adiabatic cooling of expanding upflows), with important consequence for the inferred abundances of some elements, in particular CNO.

The holy grail of quantitative stellar spectroscopy

To combine hydrodynamic (3D) modelling of stellar atmospheres with non-equilibrium (NLTE) modelling of line formation of complex atoms (like Fe and neutron-capture elements).

This has been achieved for many atoms (from Li to Ba), with important consequences for the inferred abundances.

For example, 3D-NLTE modelling of the Li-6+7 blend in metal-poor stars could convincingly refute earlier claims of a Li-6 plateau among Galactic halo stars (Lind *et al*. 2013).

Consequence: **There is no cosmological Li-6 problem!**

I claim that the cosmological Li-7 problem is significantly alleviated given known stellar physics (see Korn 2024).

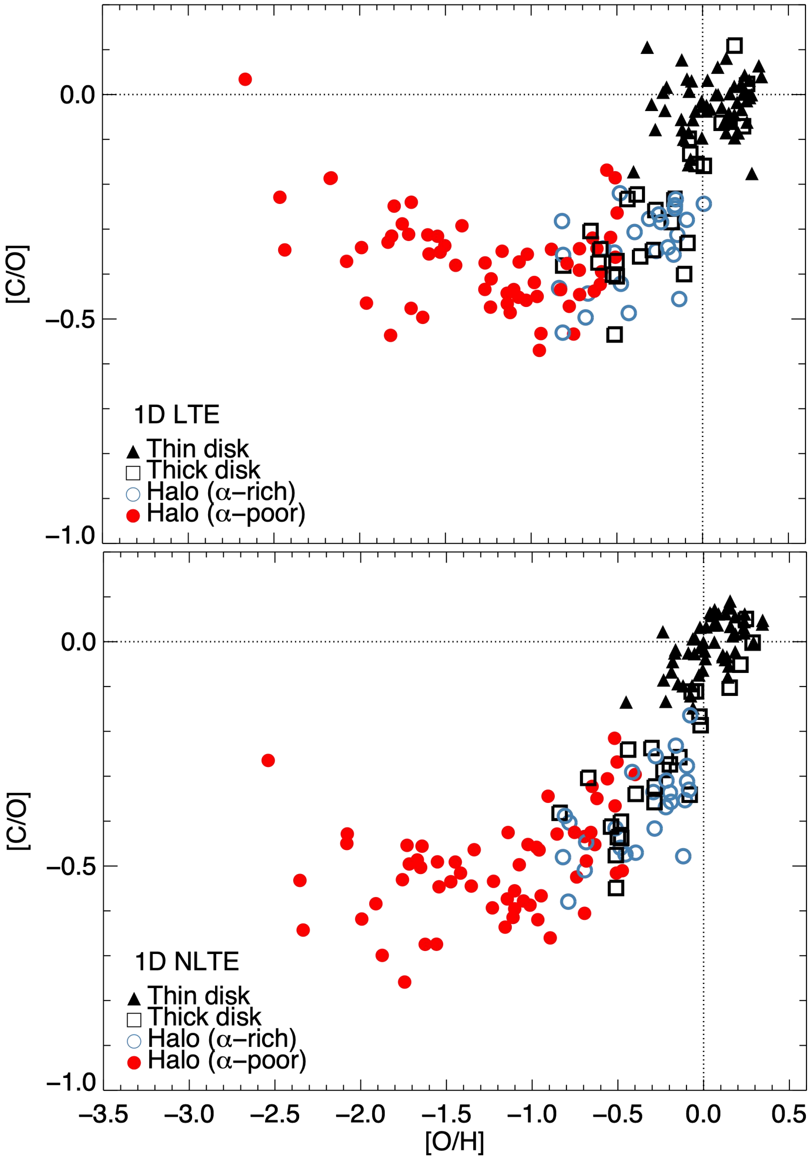
3D-NLTE Li corrections are now available for post-processing of 1D-LTE abundances (Wang *et al*. 2021). They have already been applied to the ~600,000-star GALAH survey (Wang *et al*. 2024). This is a major achievement!

C/O in metal-poor stars

Amarsi, Nissen & Skúladóttir (2019) have systematically explored C/O ratios in thin-/thick-disk and halo stars.

In 1D-LTE, an upturn in C/O ratios seems to be present in low-metallicity halo stars ([Fe/H]<−1.5).

In 1D-NLTE, the upturn rather looks like a plateau with some spread and individual outliers.



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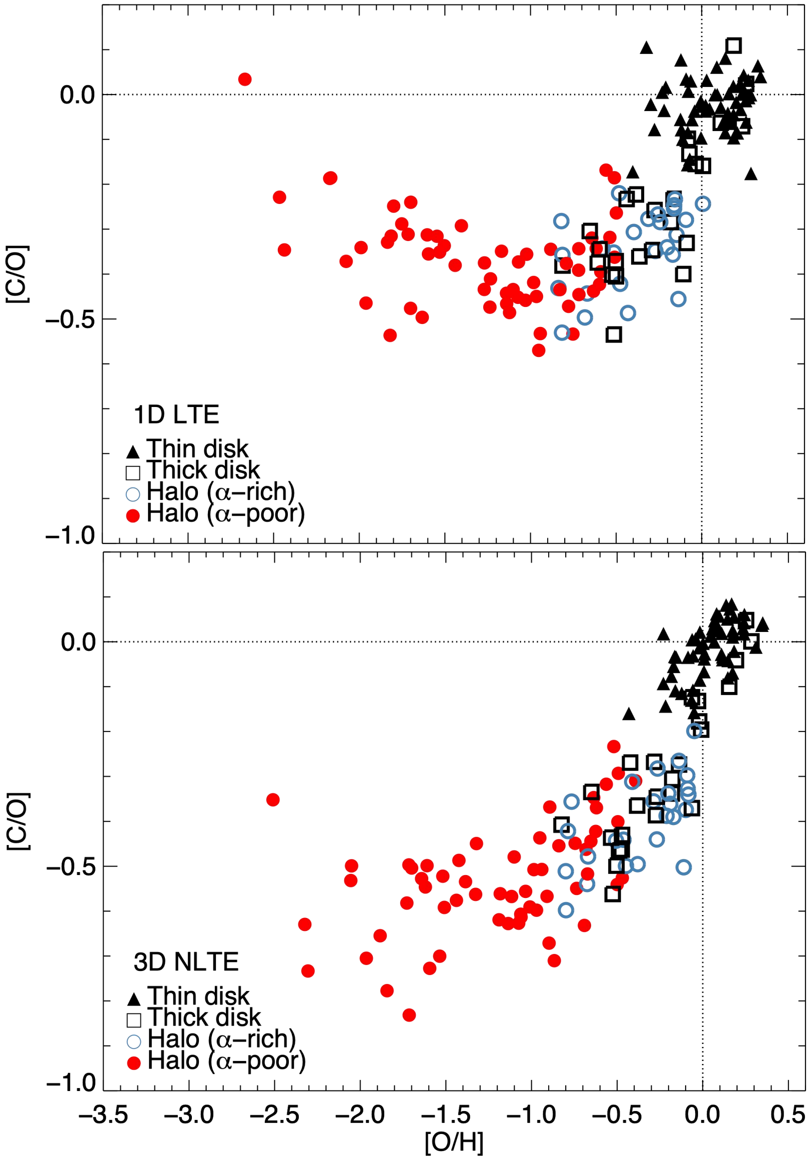
Amarsi, Nissen & Skúladóttir (2019) have systematically explored C/O ratios in thin-/thick-disk and halo stars.

In 1D-LTE, an upturn in C/O ratios seems to be present in low-metallicity halo stars ([Fe/H]<−1.5).

In 1D-NLTE, the upturn rather looks like a plateau with some spread and individual outliers.

In 3D-NLTE, the plateau-like signature is even more apparent, with consequences for the interpretation in terms of Pop III yields. (The star-to-star scatter is similar to the 1D-NLTE case.)



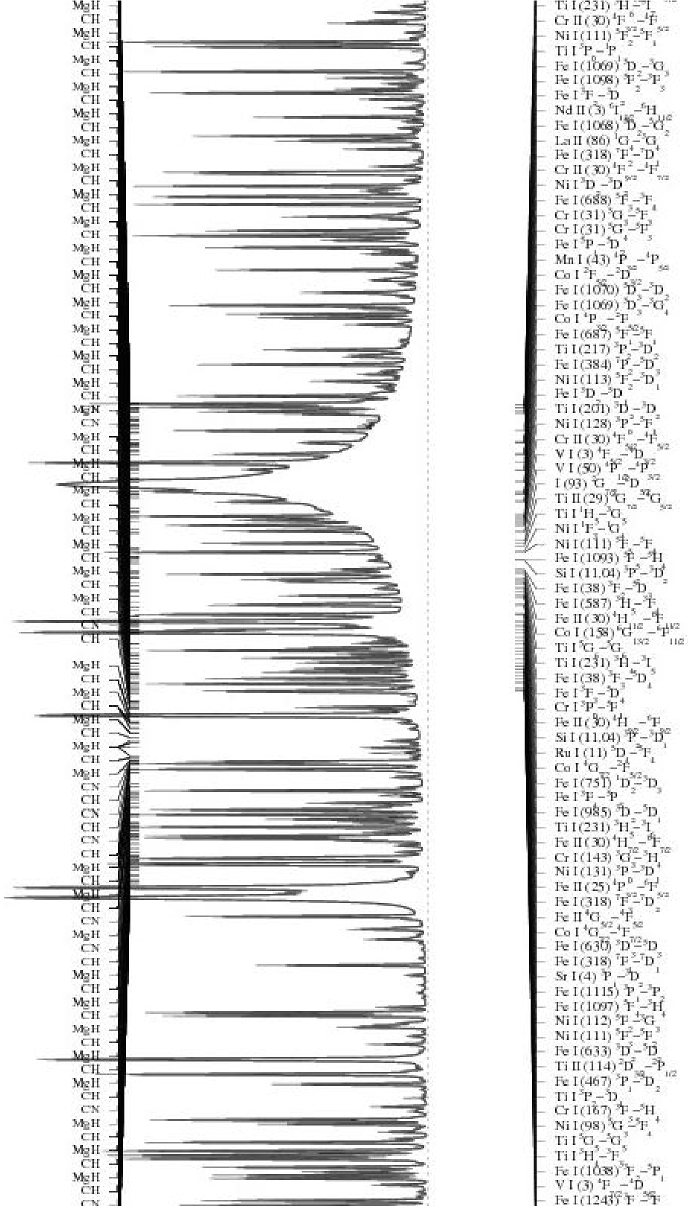
Abundances from H to U

Assuming that the stellar parameters are not biased, it is relatively easy to determine chemical abundances for your favourite element(s).

Caveats

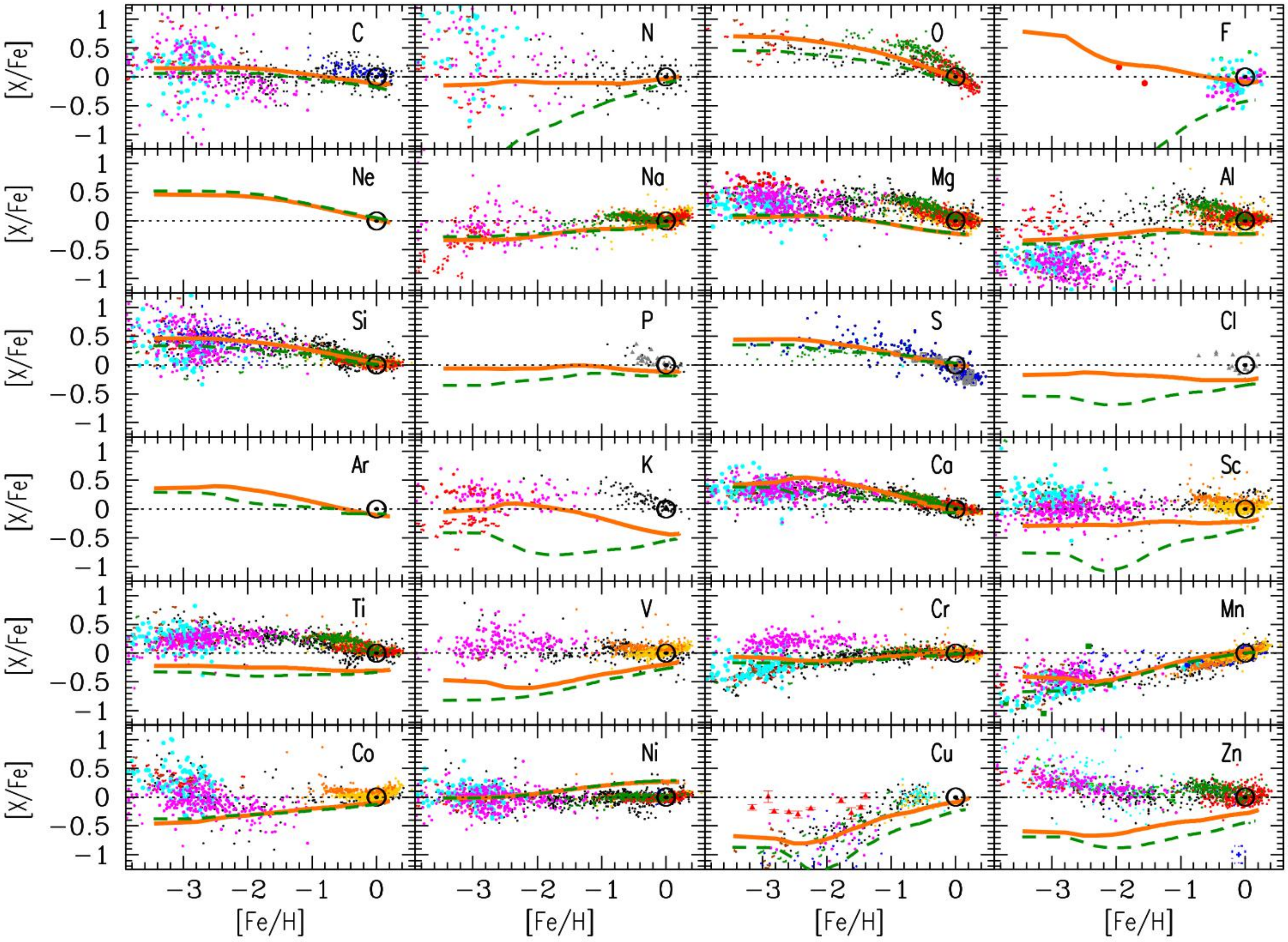
- ☐ some elements are not visible, e.g. noble gases in cool stars
- ☐ lines may lack or have inaccurate atomic data (differential abundances can help)
- ☐ lines can be subject to unmodelled effect, e.g. blending, 3D and NLTE, hfs, isotopic and Zeeman splitting

When in doubt, look for the most trustworthy 1D-NLTE or 3D-NLTE abundances.



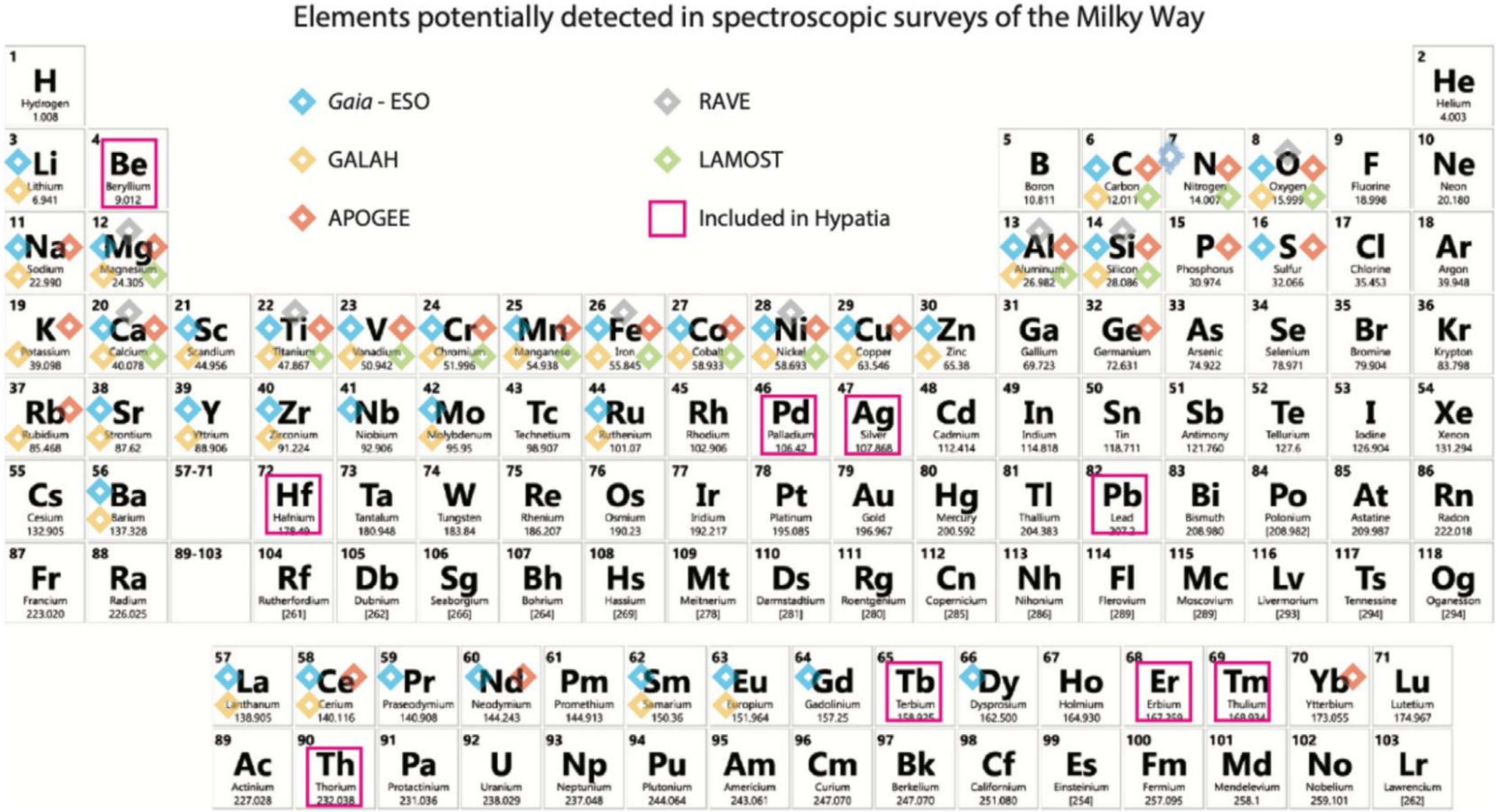
Following the elements' evolution

Abundances mot yet in 3D-NLTE…



Prantzos *et al.* (2018)

Deep and wide stellar surveys



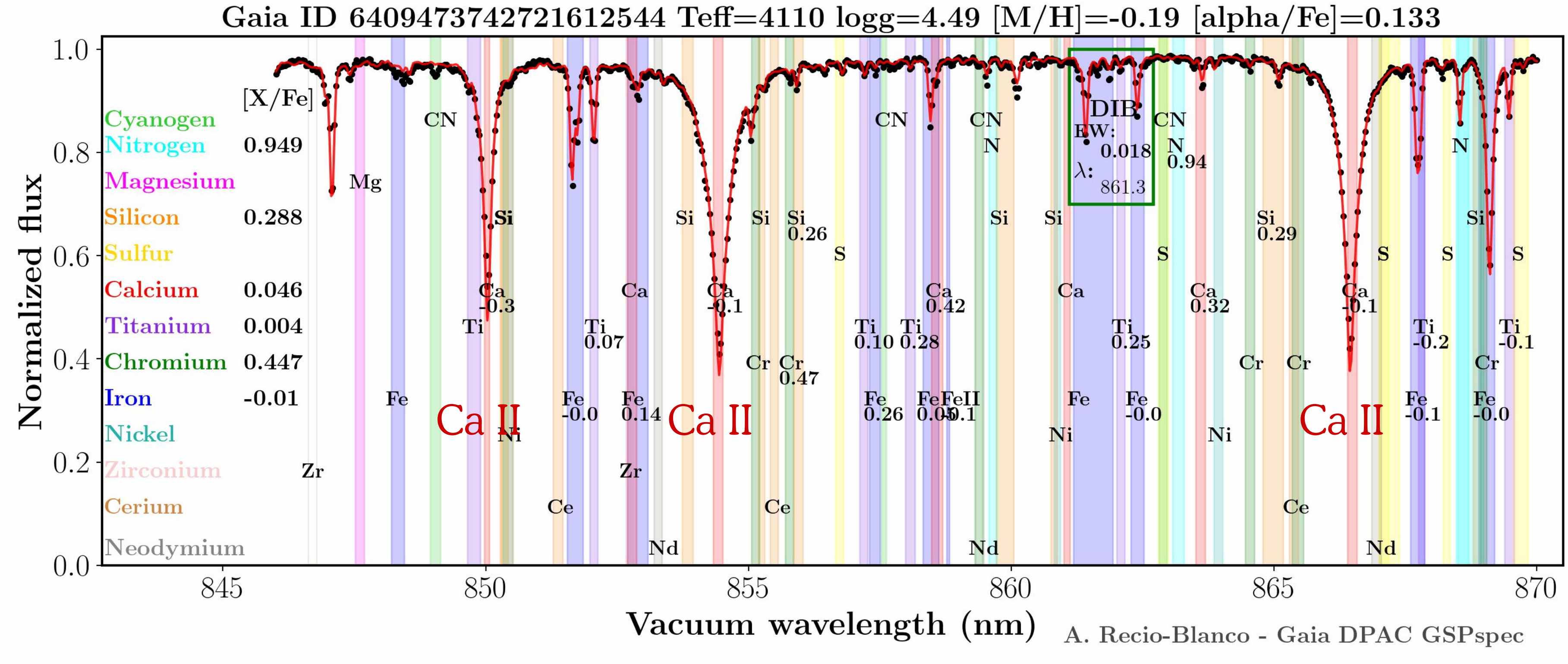
Jofré *et al.* (2019)

And the near future will see surveys with *tens of millions* of spectra: **WEAVE** (N) and **4MOST** (S). Plus **Gaia-RVS**!

Gaia RVS: spectra by the million

With Data Release 3 (June 2022), the Gaia archive contains individual chemical abundances for a few million stars. Up to a ~dozen elements are derived from the RVS spectra (R=11500, see below).

Gaia RVS is a growing treasure trove of stellar surface abundances!



A. Recio-Blanco - Gaia DPAC GSPspec

The future of stellar abundances...

ELT with ANDES
and MOSAIC...

...and Gaia RVS, and APOGEE,
and PLATO, and MSE...

... is bright indeed!

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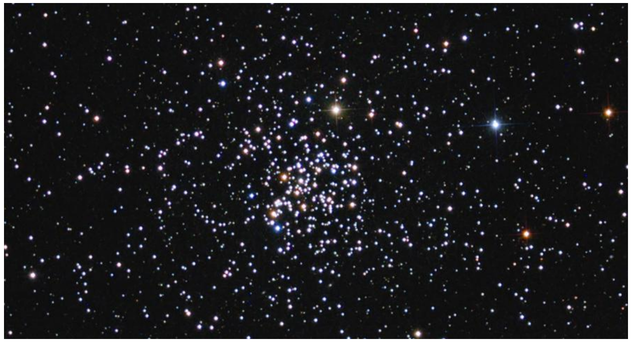
The role of environment

Open clusters: youngish (0-6 Gyr) and fairly simple. Excellent objects to test one's analysis methods *and* study stellar evolution. There are **many details left to understand** (dredge-up, role of rotation…)!



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Globular clusters: very old, very complex, ***unexplained***! See e.g. Milone *et al*. (2017) for chromosome-map analysis. HRD construction: https://youtu.be/HWQslu4S5eQ

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MW satellite galaxies, e.g. dSph galaxies: remnants of the hierarchical structure formation of the MW. Some have totally **unique abundance patterns** constraining individual enrichment events (see e.g. Ji *et al*. (2016) on Reticulum II).