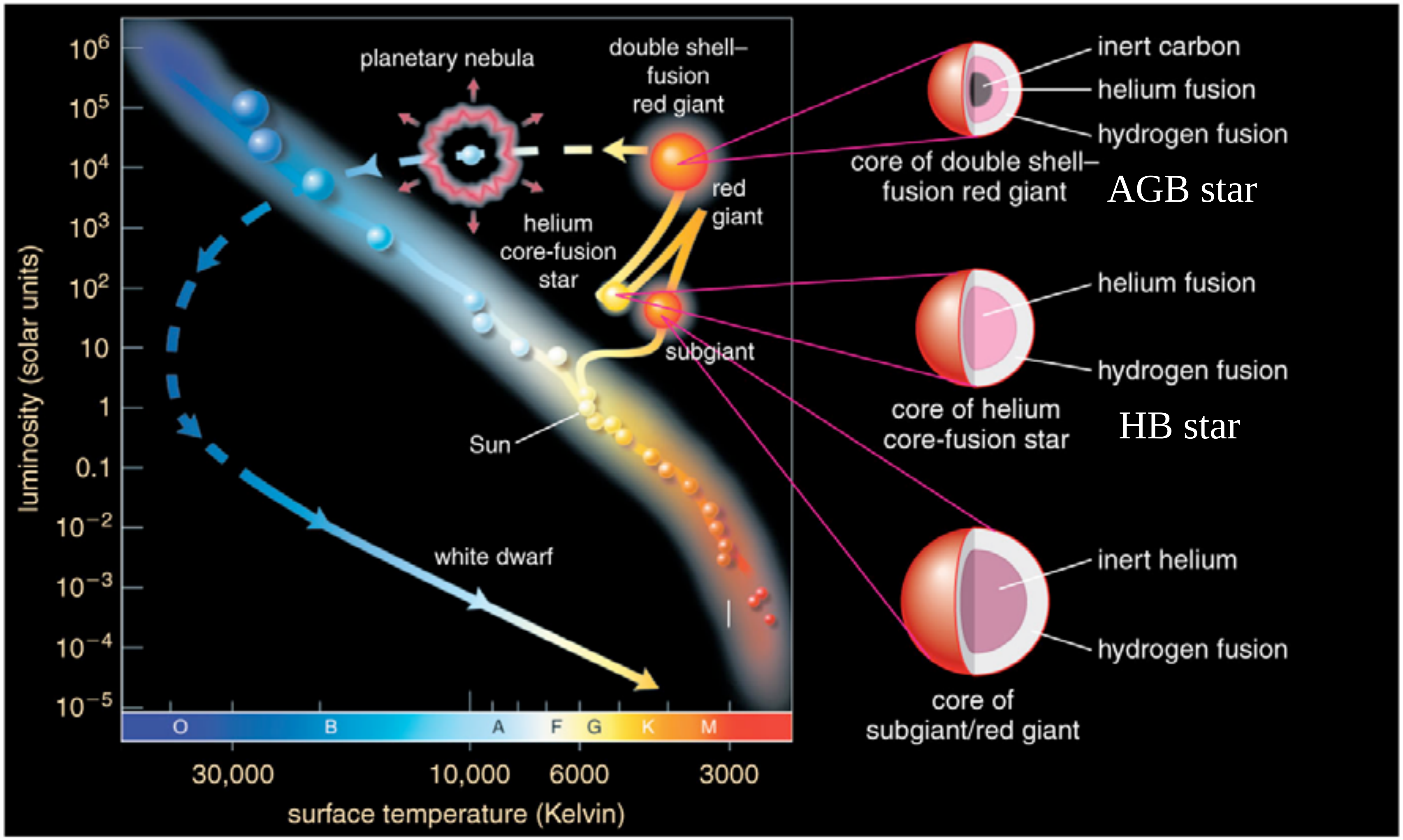
LECTURE 15
– Jerome Fang -

- Making heavy elements in low-mass stars: the s-process (review)
- White dwarfs: diamonds in the sky
- Evolution of high-mass stars ($M > 8\ M_{\odot}$); post-helium burning fusion processes

Review of evolution of low-mass stars ($M < 8\ M_{\odot}$)



Making heavy elements: the s-process

During the AGB phase (double shell burning), additional elements are produced via the s-process in the outer shells of the star

Heavier elements are produced via neutron capture onto nuclei (add neutrons to nuclei one-by-one)

The neutrons are byproducts of helium and carbon burning (see next slide)

S-process can make elements from cobalt up to bismuth

The “s” stands for “slow” because products decay easily, making it difficult to form stable nuclei

Additional Nucleosynthesis – The s-process.

- During helium burning:

$$^{14}N + {}^{4}He \rightarrow {}^{18}F + \gamma$$

$$^{18}F \rightarrow {}^{18}O + e^{+} + \nu$$

$$^{18}O + {}^{4}He \rightarrow {}^{22}Ne + \gamma$$

$$^{22}Ne + {}^{4}He \rightarrow {}^{25}Mg + n \qquad M \gtrsim 8\,M_{\odot}$$

In $M \lesssim 8\,M_{\odot}$ during "AGB" stage, a little H may get mixed into the thin He-burning shell. Then

$$^{12}C + p \rightarrow {}^{13}N \rightarrow {}^{13}C + e^{+} + \nu$$

$$^{13}C + {}^{4}He \rightarrow {}^{16}O + n$$

Where do the neutrons go?

Iron nuclei are the "seeds" of the s-process

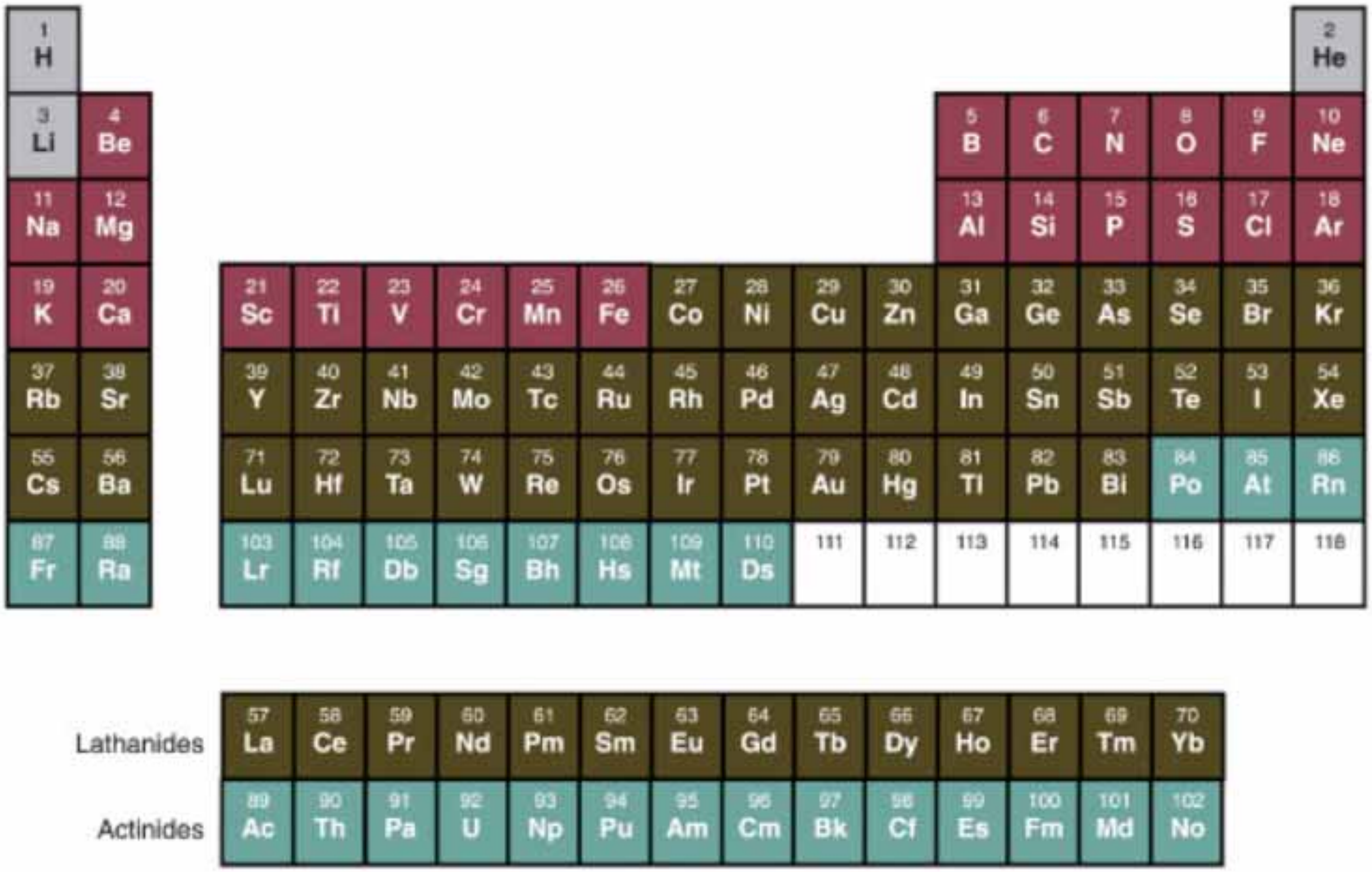
$$\left.\begin{array}{l} {}^{56}Fe + n \rightarrow {}^{57}Fe + \gamma \\ {}^{57}Fe + n \rightarrow {}^{58}Fe + \gamma \\ {}^{58}Fe + n \rightarrow {}^{59}Fe + \gamma \end{array}\right\} \quad \text{years}$$

$${}^{59}Fe \rightarrow {}^{59}Co + e^{-} + \bar{\nu} \qquad \text{44.5 days}$$

$${}^{59}Co + n \rightarrow {}^{60}Co + \gamma$$

etc., all the way to ^{209}Bi.

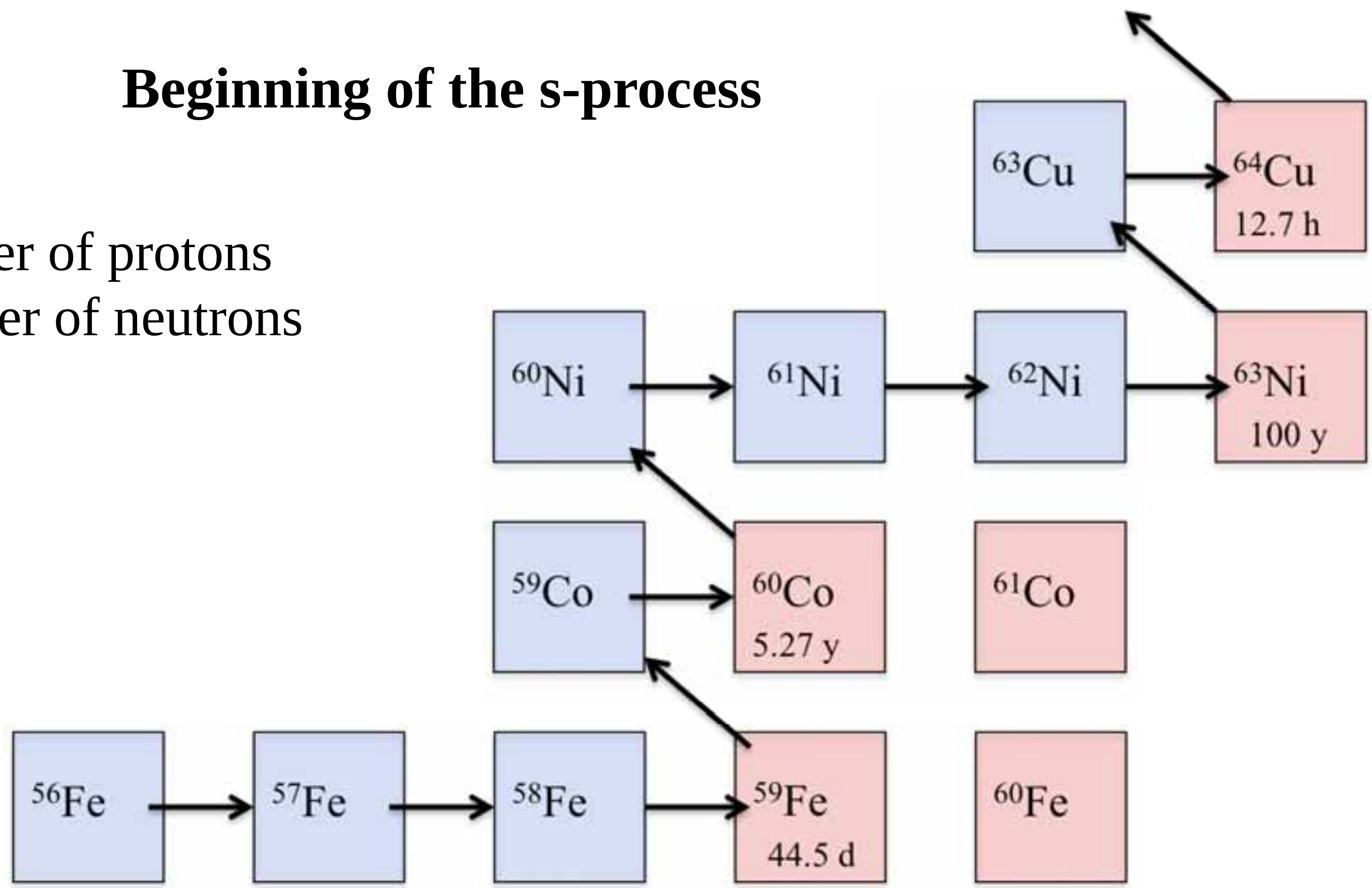
This is called the "slow" process of neutron addition or the "s-process". (There is also a "r-process")



Magenta = elements produced by fusion
Brown = elements produced by the s-process

Beginning of the s-process

Z = Number of protons
N = Number of neutrons



Z

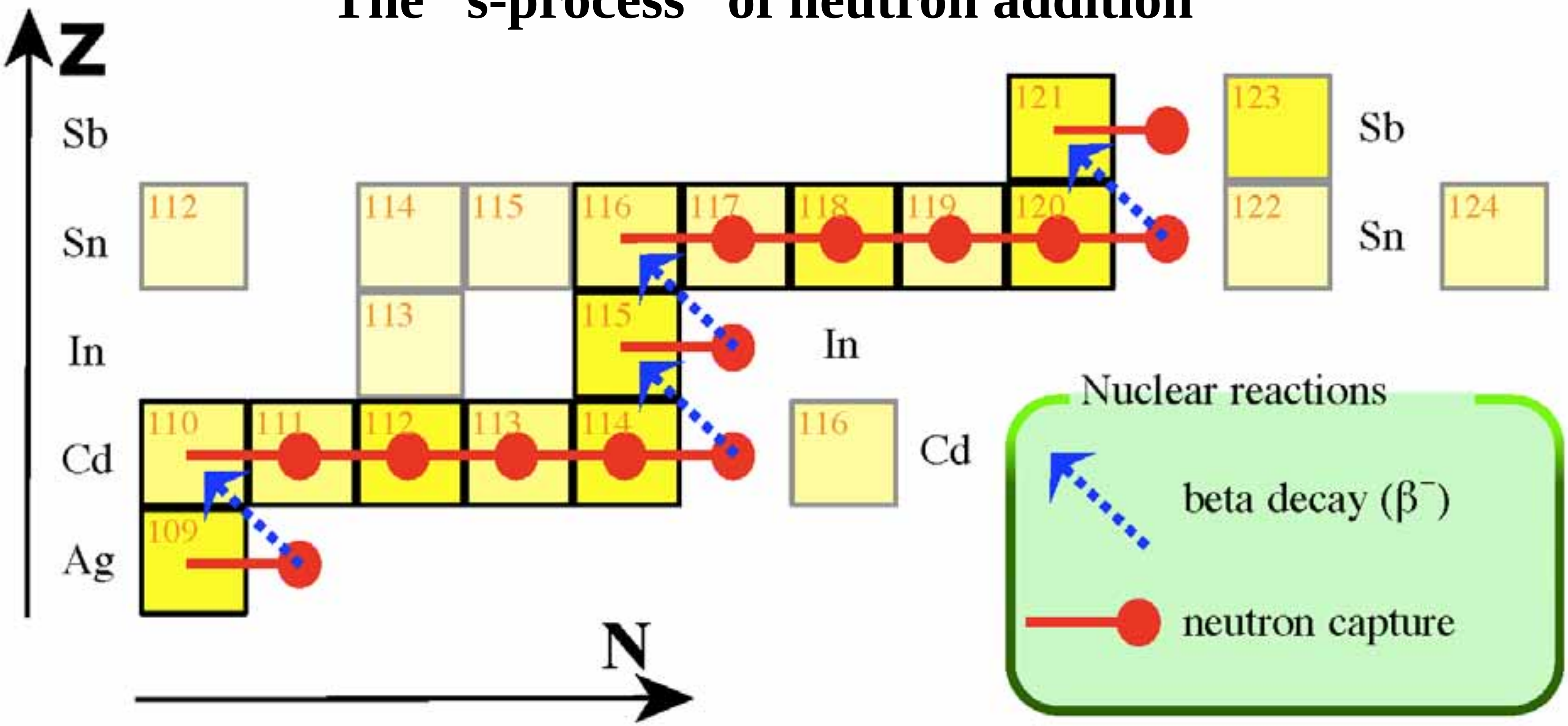
N

→ = (capture n, release γ)

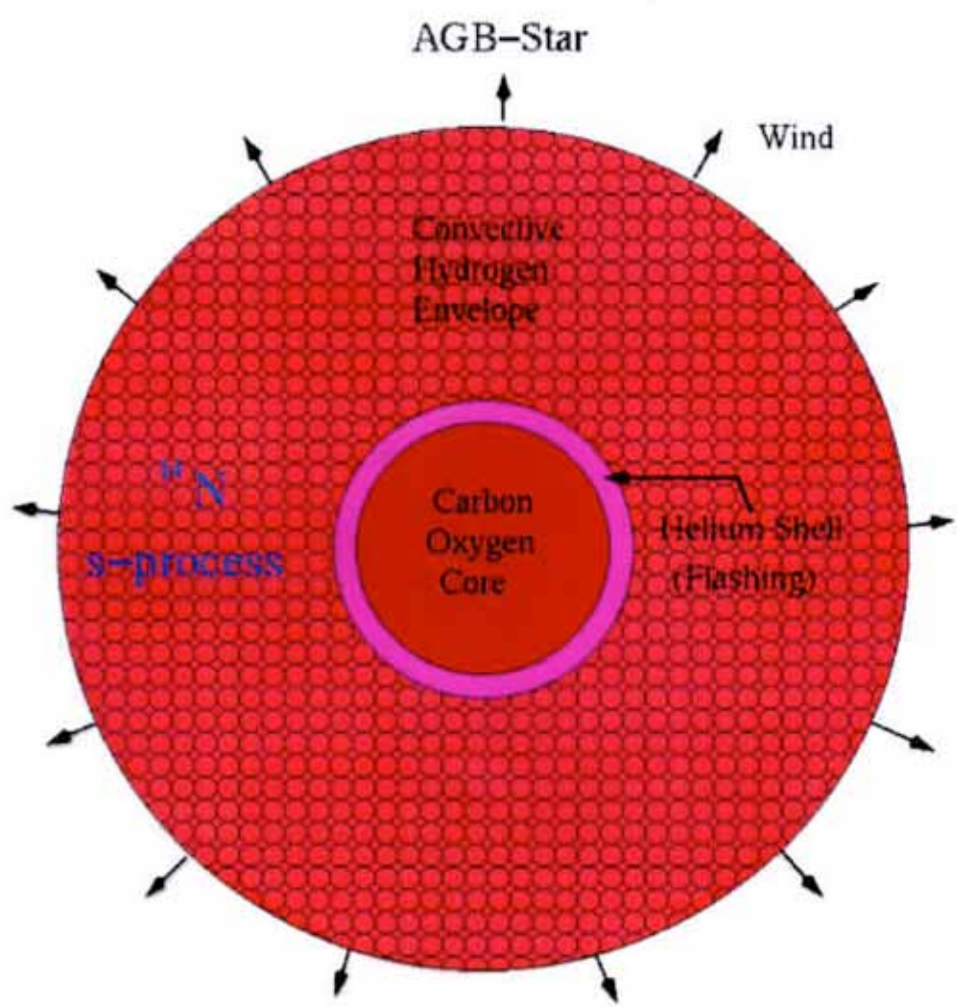
↖ = (release e^-,ν)

Each row = different isotopes of an element

The "s-process" of neutron addition



Each neutron capture takes you one step to the right in this diagram. Each decay of a neutron to a proton inside the nucleus moves you up a left diagonal.



CO–core temperature about 5×10^8 K
He–shell flash $T \approx 3 \times 10^8$ K

Convective hydrogen envelope dredges up ashes of helium shell flashes to the surface where they are lost to wind and planetary nebula formation.

On top of the He burning shell there is also a thin H burning shell

WHITE DWARFS

General properties of white dwarfs

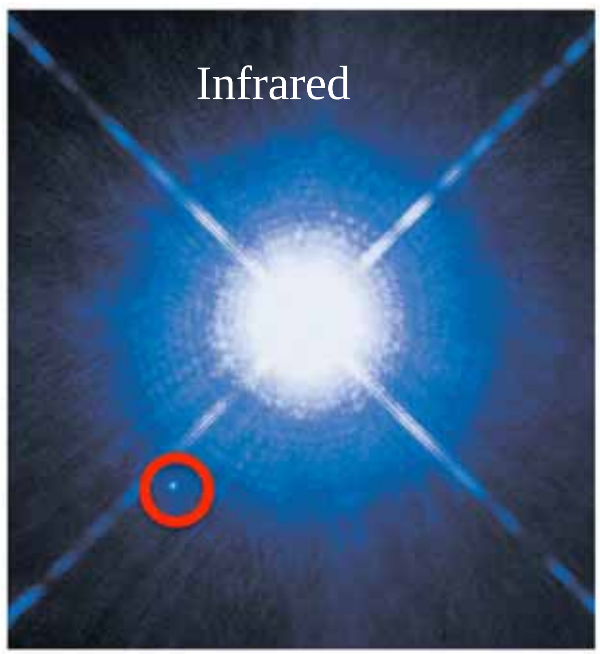
•The remnant core of a low-mass star, containing mainly carbon and oxygen (from helium burning)

•Supported by electron degeneracy pressure; no fusion occurring

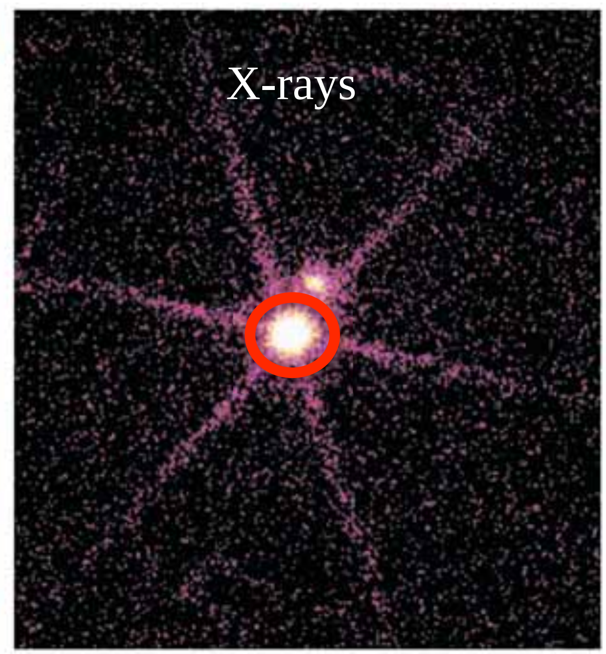
•High surface temperatures (~10^4-10^5 K), small radii (about the size of Earth)

Sirius is a binary system

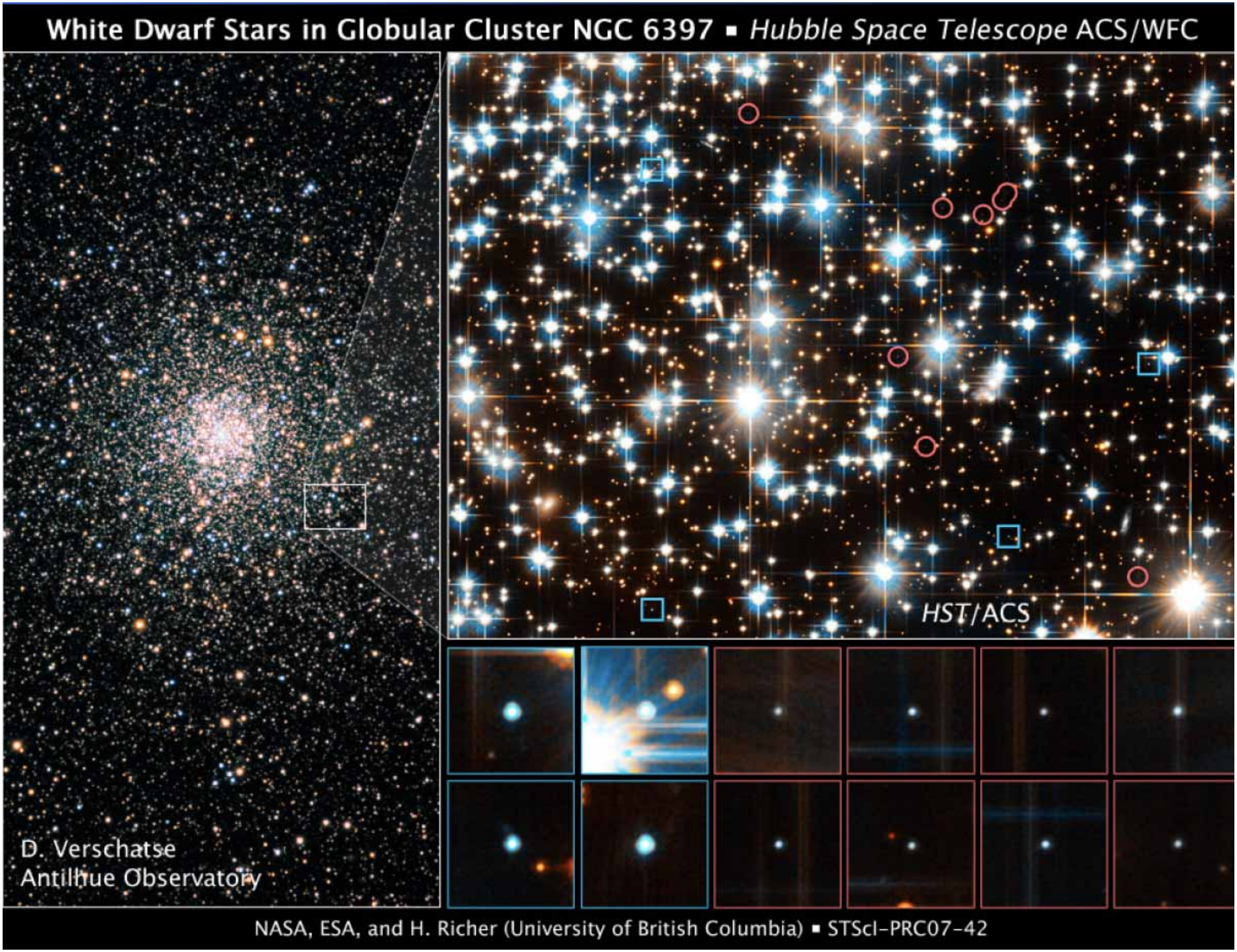
Sirius B is a WD
Sirius B is easily seen in X-rays



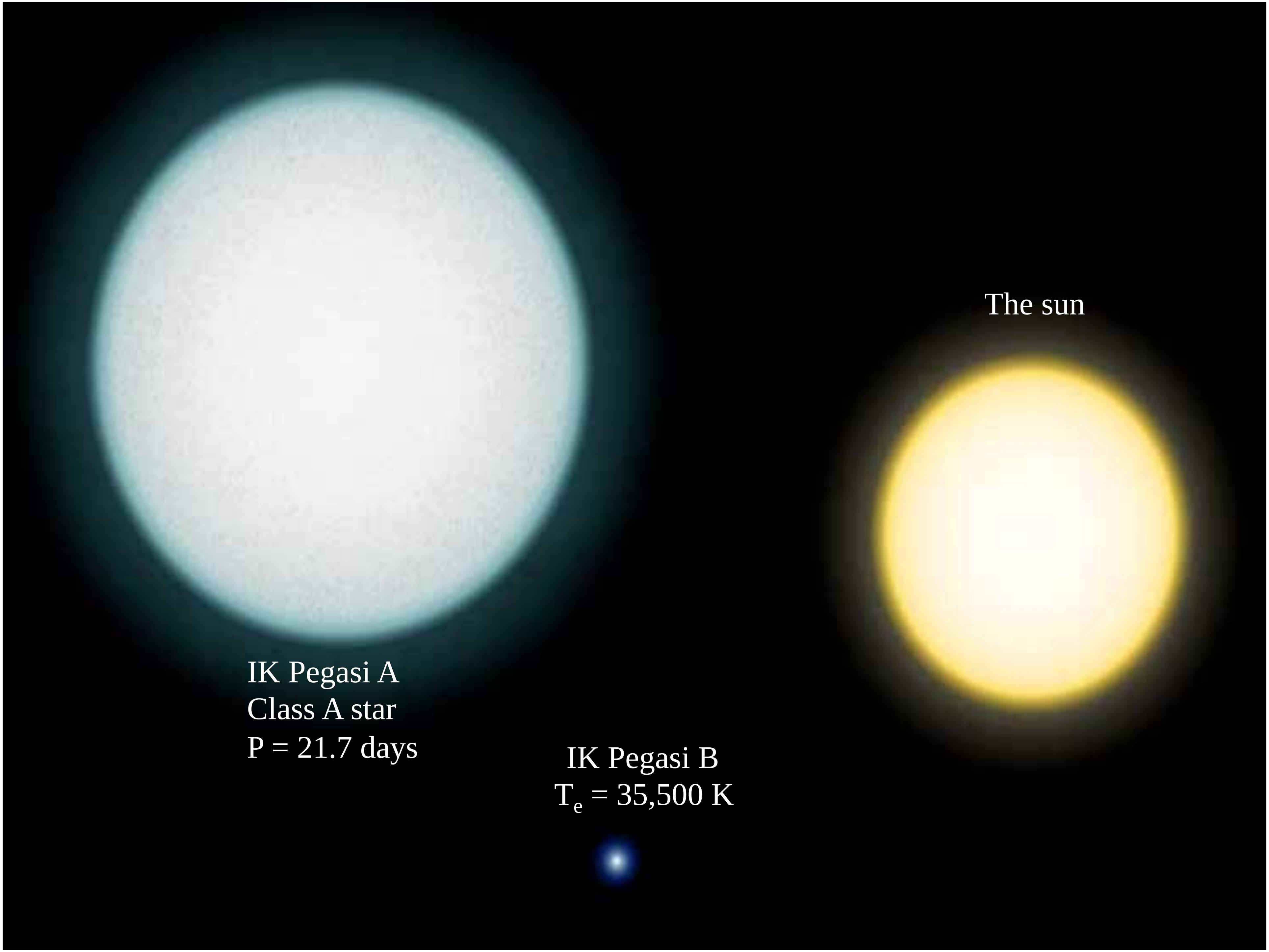
a Sirius as seen in infrared light by the Hubble Space Telescope.

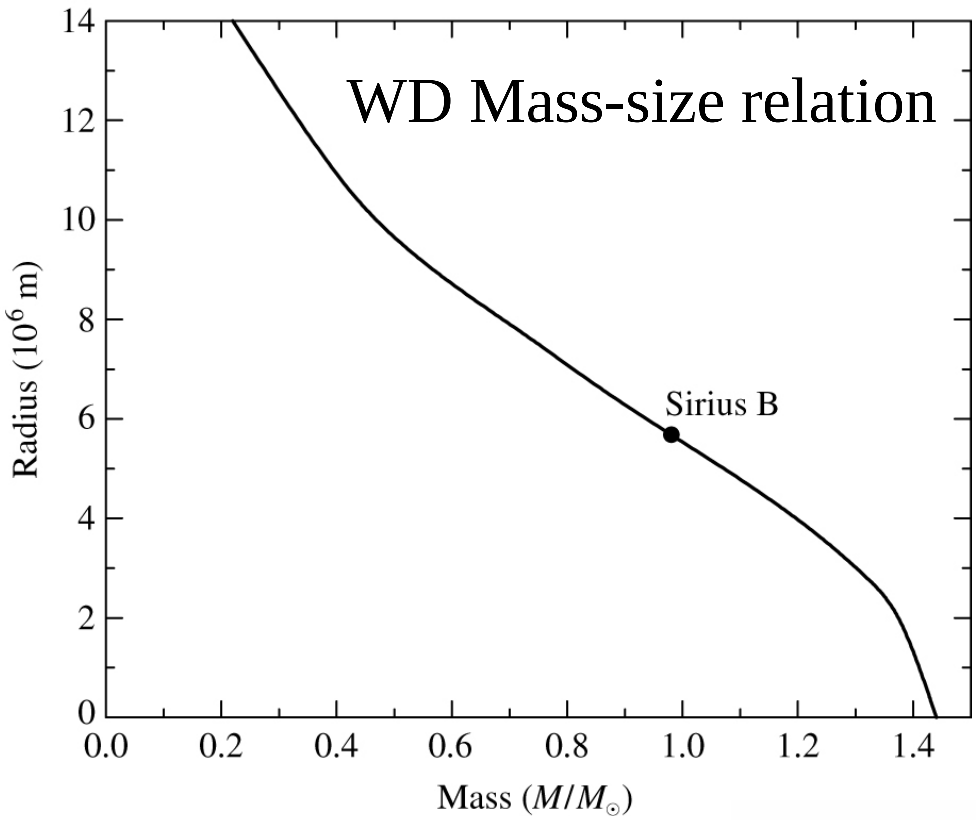


b Sirius as seen by the Chandra X-ray Telescope.



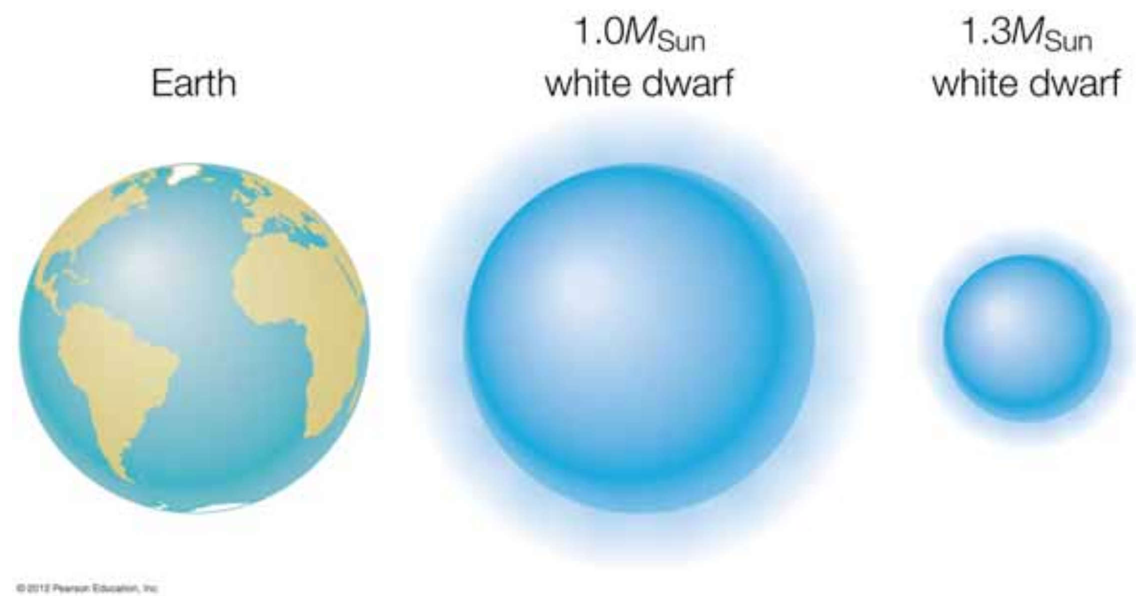
White Dwarf Stars in Globular Cluster NGC 6397 ▪ *Hubble Space Telescope* ACS/WFC
HST/ACS
D. Verschatse
Antilhue Observatory
NASA, ESA, and H. Richer (University of British Columbia) ▪ STScI-PRC07-42


The sun
IK Pegasi A
Class A star
P = 21.7 days
IK Pegasi B
T_e = 35,500 K



The more massive a WD, the <u>smaller</u> its size…

It's a consequence of degeneracy pressure…



MASS RADIUS RELATION FOR WHITE DWARFS

$$P_c = \frac{GM\rho}{2R} = 1.00 \times 10^{13} \ (\rho Y_e)^{5/3}$$

if supported by non-relativistic electron degeneracy pressure

$$\frac{GM\rho}{2R} \approx \frac{GM\left(\frac{3M}{4\pi R^3}\right)}{2R} = 1.00 \times 10^{13}\left(\frac{3M}{4\pi R^3}\right)^{5/3}\left(\frac{1}{2}\right)^{5/3}$$

$$\left(\frac{3G}{8\pi}\right)\frac{M^2}{R^4} = 1.00 \times 10^{13}\left(\frac{3}{8\pi}\right)^{5/3} \frac{M^{5/3}}{R^5}$$

$$M^{1/3} = 1.00 \times 10^{13}\left(\frac{3}{8\pi}\right)^{2/3} \frac{1}{GR}$$

$$R = 3.63 \times 10^{19} \ / \ M^{1/3}$$

$$R = 2.9 \times 10^{8}\left(\frac{M_\odot}{M}\right)^{1/3} \ \text{cm}$$

$$R \propto \frac{1}{M^{1/3}}$$

Actually $\boxed{5 \times 10^8\,(M_\odot/M)^{1/3}}$ cm is more accurate.

$$R \propto \frac{1}{M^{1/3}}$$

$$\begin{aligned}\rho &= \frac{3M}{4\pi R^3} \\ &= \frac{(3)(1.99 \times 10^{33})}{(4\pi)(5 \times 10^8)^3}\left(\frac{M}{M_\odot}\right)^2 \\ &= 4 \times 10^6 \left(\frac{M}{M_\odot}\right)^2 \text{ g cm}^{-3}\end{aligned}$$

- Note implication: As M goes up, R gets smaller and ρ gets larger.

APPEARANCE

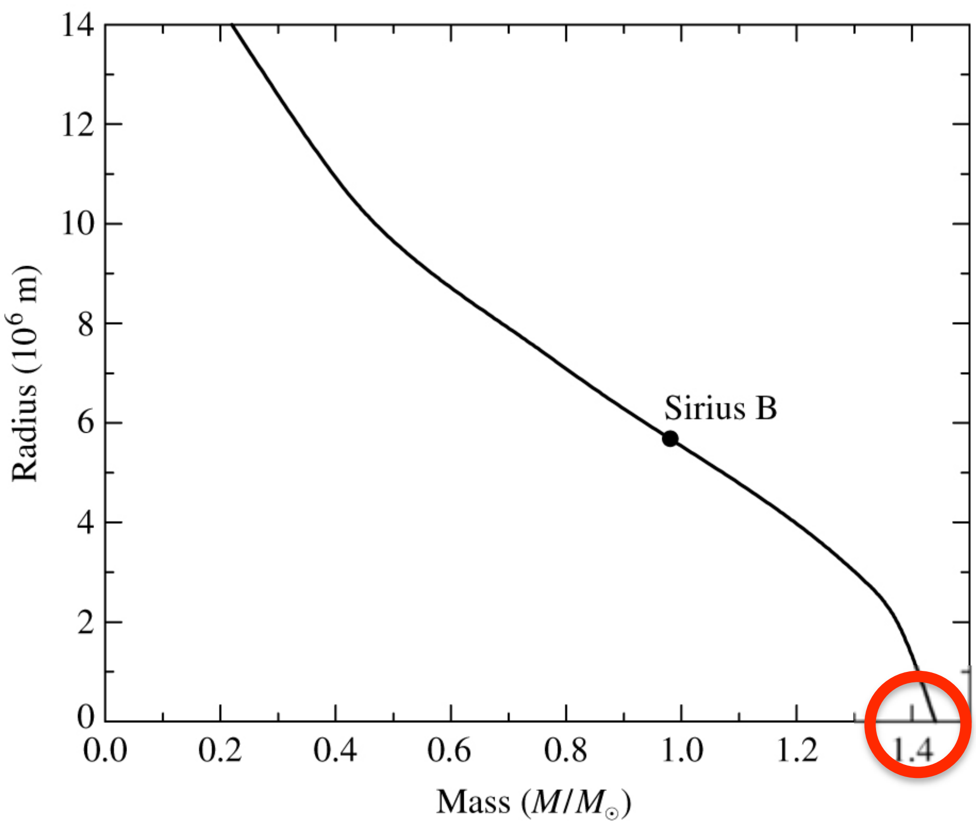
- $L \approx 0.01\, L_\odot$

$$\begin{aligned}T_e &= \left(\frac{L}{4\pi\sigma R^2}\right)^{1/4} \\ &\approx \left[\frac{4 \times 10^{31}}{(4\pi)(5 \times 10^8)^2(5.6 \times 10^{-5})}\right]^{1/4} \\ &= 20,000\text{ K}\end{aligned}$$

$$\begin{aligned}\lambda_{max} &= \frac{2.89 \times 10^7\text{ A}}{T_{eff}} \\ &= 1400\text{ A}\end{aligned}$$

White dwarfs are known with temperatures ranging from 4000 K to 200,000 K

The mass-radius relationship implies that there is a maximum mass for a WD (when its radius becomes zero)



This maximum WD mass is called the <u>Chandrasekhar mass</u>, M_{ch}, after its discoverer

The value of M_{ch} is ~1.4 $M_\odot$

Estimate of the Chandrasekhar mass

As WD mass increases, its radius shrinks → density increases → electrons move faster and faster in core (almost speed of light) →Relativistic degeneracy pressure becomes important at around a density of $\rho \sim 10^7$ g/cm^3

If we set central pressure equal to relativistic degeneracy pressure, we can find an expression for M_{ch}

$$P_c = P_{\text{deg}}^R$$
$$\frac{GM\rho}{2R} = 1.24 \times 10^{15} (\rho Y_e)^{4/3}$$

THE CHANDRASEKHAR MASS

As M gets larger and the radius decreases, the density rises

Eventually at ρ greater than about 10^7 g cm^{-3} electrons in the *central* part of the white dwarf start to move close to the speed of light. As the mass continues to grow, a larger fraction of the star is supported by relativistic electron degeneracy pressure. Consider the limit:

$$\mathrm{P}_{\mathrm{deg}}^{R} = 1.24 \times 10^{15} \left(\rho Y_e\right)^{4/3} = \frac{GM\rho}{2R}$$

As usual examine the constant density case for guidance

$$\rho \approx \left(\frac{3M}{4\pi R^3}\right)$$

$$1.24 \times 10^{15}\ \cancel{\rho}\, Y_e^{4/3} \left(\frac{3M}{4\pi R^3}\right)^{1/3} = \frac{GM\cancel{\rho}}{2R} = P_{central}$$

Nb. R drops out

$$M^{2/3} = 1.24 \times 10^{15}\ Y_e^{4/3} \left(\frac{3}{4\pi}\right)^{1/3} \frac{2}{G}$$

$$M^{2/3} = 2.3 \times 10^{22}\ Y_e^{4/3}$$

$$M = 3.5 \times 10^{33} Y_e^2 \text{ gm} = 1.75\ Y_e^2\ M_\odot$$

Actually $\boxed{M = 5.7\ Y_e^2\ M_\odot = 1.4\ M_\odot}$ if $Y_e = 0.5$

Aside:

This result extends beyond white dwarfs.

There can be no stable star whose pressure depends on its density to the 4/3 power

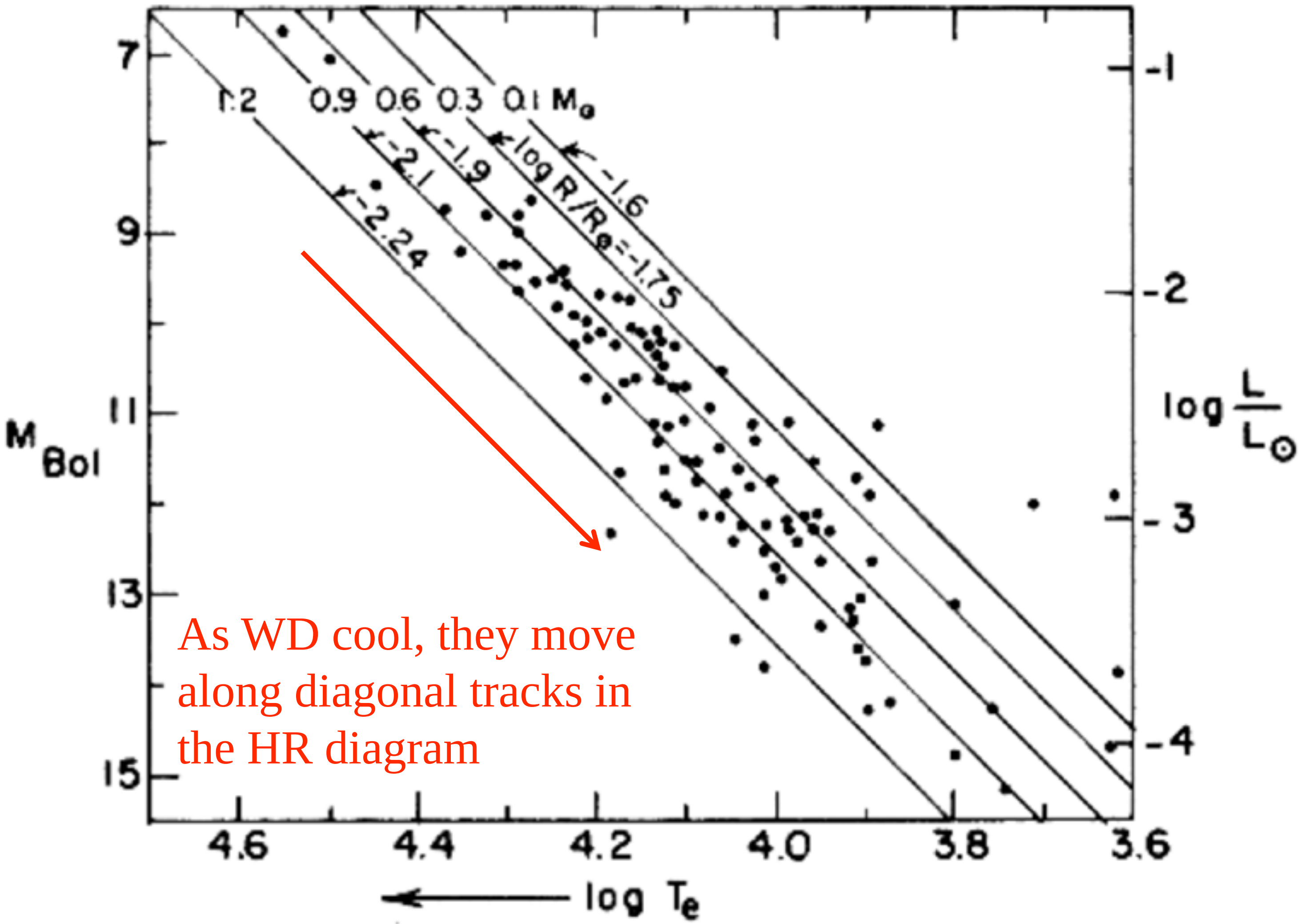
This is important later on when we talk about the maximum mass a star can have

Table 8.5 Central Densities, Total Mass, and Radius of Different White Dwarf Models, Taking $\mu_e = 2$ (Negligible Hydrogen Concentration)[a]

$\log \rho_c$		$M/M_\odot$	$\log R/R_\odot$
5.39		0.22	−1.70
6.03		0.40	−1.81
6.29		0.50	−1.86
6.56		0.61	−1.91
6.85		0.74	−1.96
7.20	↕ ↓ relativistic	0.88	−2.03
7.72	e^-	1.08	−2.15
8.21		1.22	−2.26
8.83		1.33	−2.41
9.29		1.38	−2.53
∞		1.44	$-\infty$ ⇒ R = 0

[a]See text for comments. (After M. Schwarzschild Sc58b.) From *Structure and Evolution of the Stars* (copyright © 1958 by Princeton University Press) p. 232.

EVOLUTION OF WHITE DWARF STARS



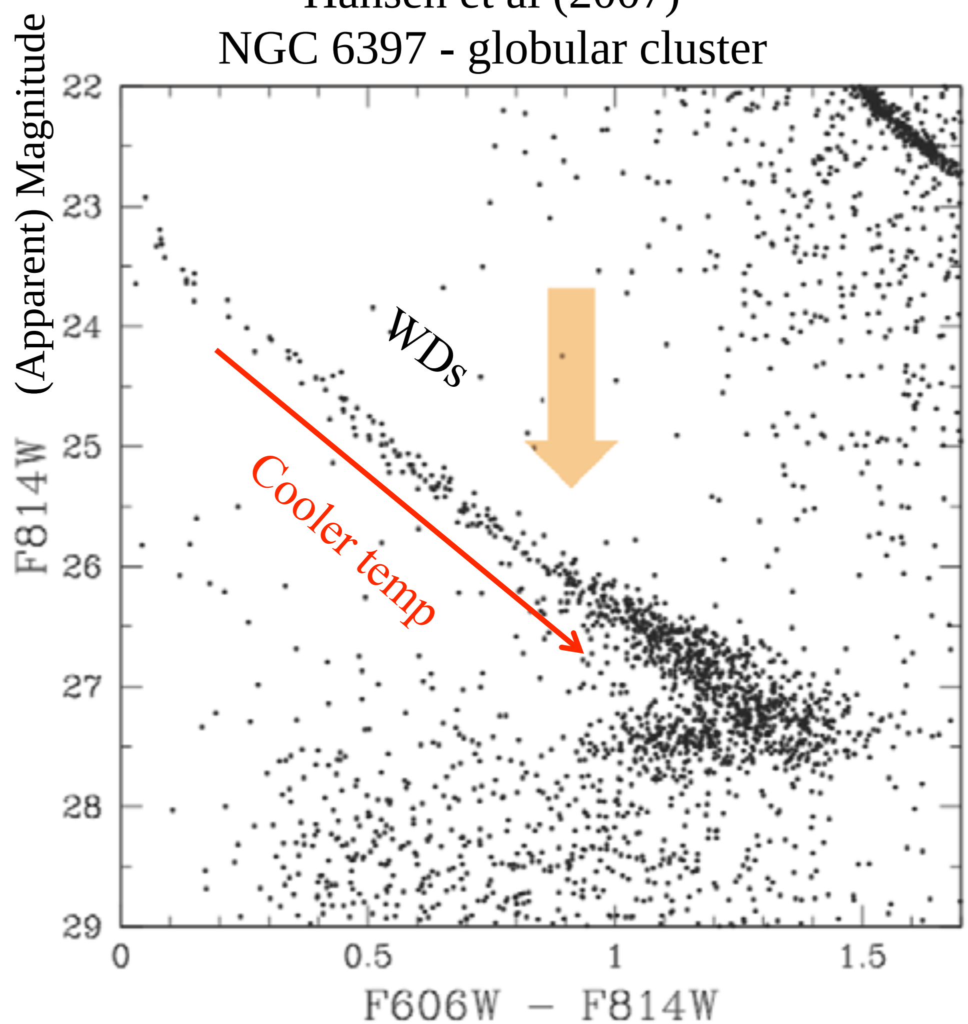
For a WD of constant mass, R = constant

Cooling and crystallization in white dwarfs

When the interior temperature declines to ~5000 K, the carbon and oxygen start to crystallize into a lattice. This crystallization releases energy and provides a source of luminosity that slows the cooling.

WDs "pile up" at the end of the cooling track.

Hansen et al (2007)
NGC 6397 - globular cluster



Color

http://en.wikipedia.org/wiki/White_dwarf

The coolest, faintest white dwarfs still have a surface temperature of ~4000 K. The universe is not old enough for “black dwarfs” to have formed yet.

E.g., 0.59 solar mass WD - like the sun will make - takes about 1.5 billion years to cool to 7140 K and another 1.8 billion years to cool to 5550 K.

Critical Masses

$0.08\ M_{\odot}$

Contracting protostars below this mass do not ignite hydrogen burning on the main sequence. They become brown dwarfs or planets.

$0.50\ M_{\odot}$

Stars below this mass are completely convective on the main sequence
“ “ “ “ do not ignite helium burning

$2.0\ M_{\odot}$

Stars below this mass (and above .5) experience the helium core flash
Stars above this mass are powered by the CNO cycle (below by the pp-cycles)
Stars above this mass have convective cores on the main sequence (and radiative surfaces)

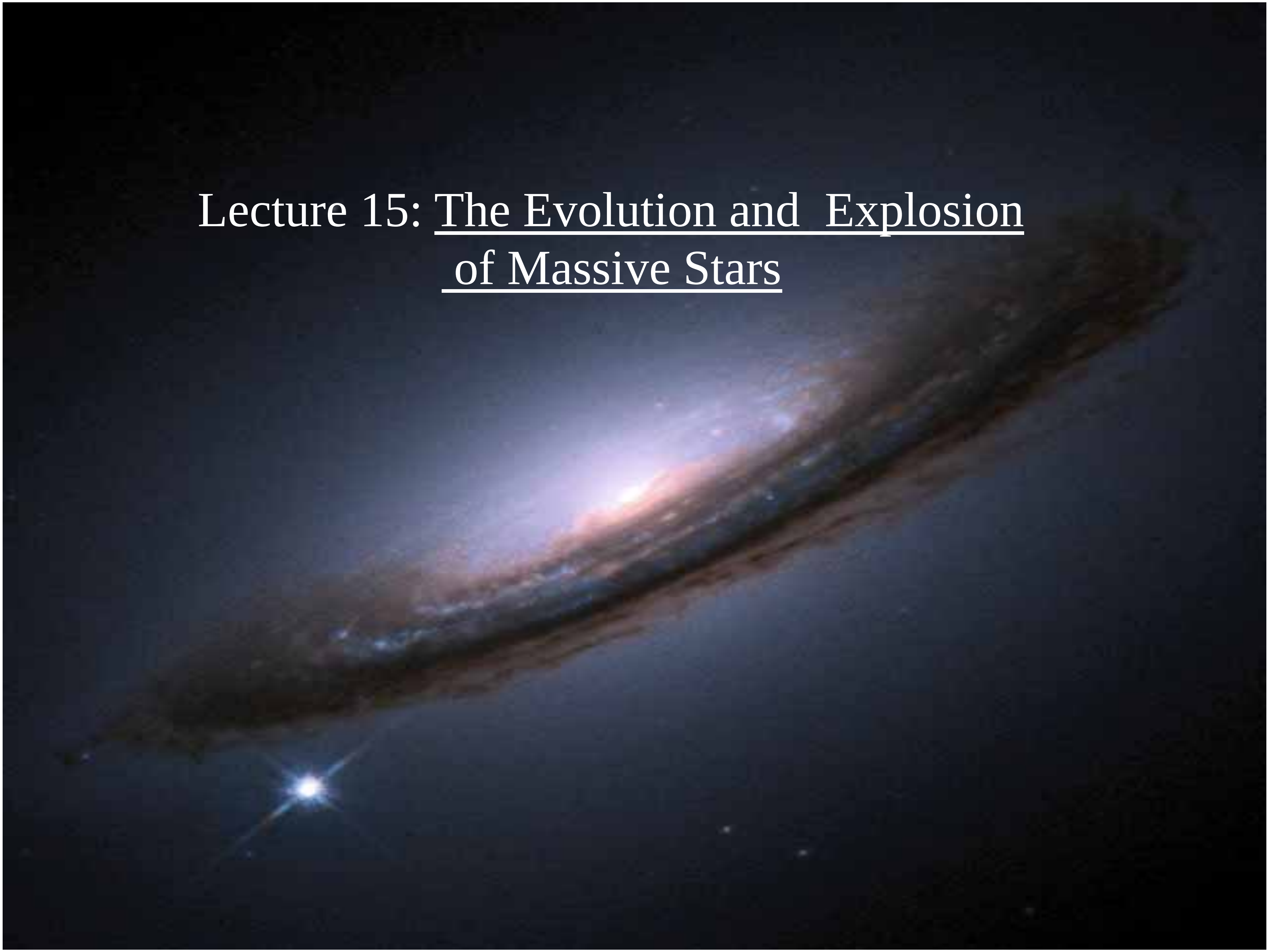
$8\ M_{\odot}$

Stars below this mass do not ignite carbon burning. They end their lives as planetary nebulae and white dwarfs. Stars above this mass make supernovae.

$\sim 150\ M_{\odot}$

Population I stars much above this mass pulse apart on the main sequence. No heavier stars exist.

Lecture 15: The Evolution and Explosion of Massive Stars



MAXIMUM MASS STAR

Because of the increasing dominance of radiation pressure, stars much above 100 solar masses become *pulsationally unstable* and experience episodes of *violent mass ejection* (not Cepheids nor supernovae or planetary nebulae, but a lot of fast mass loss).

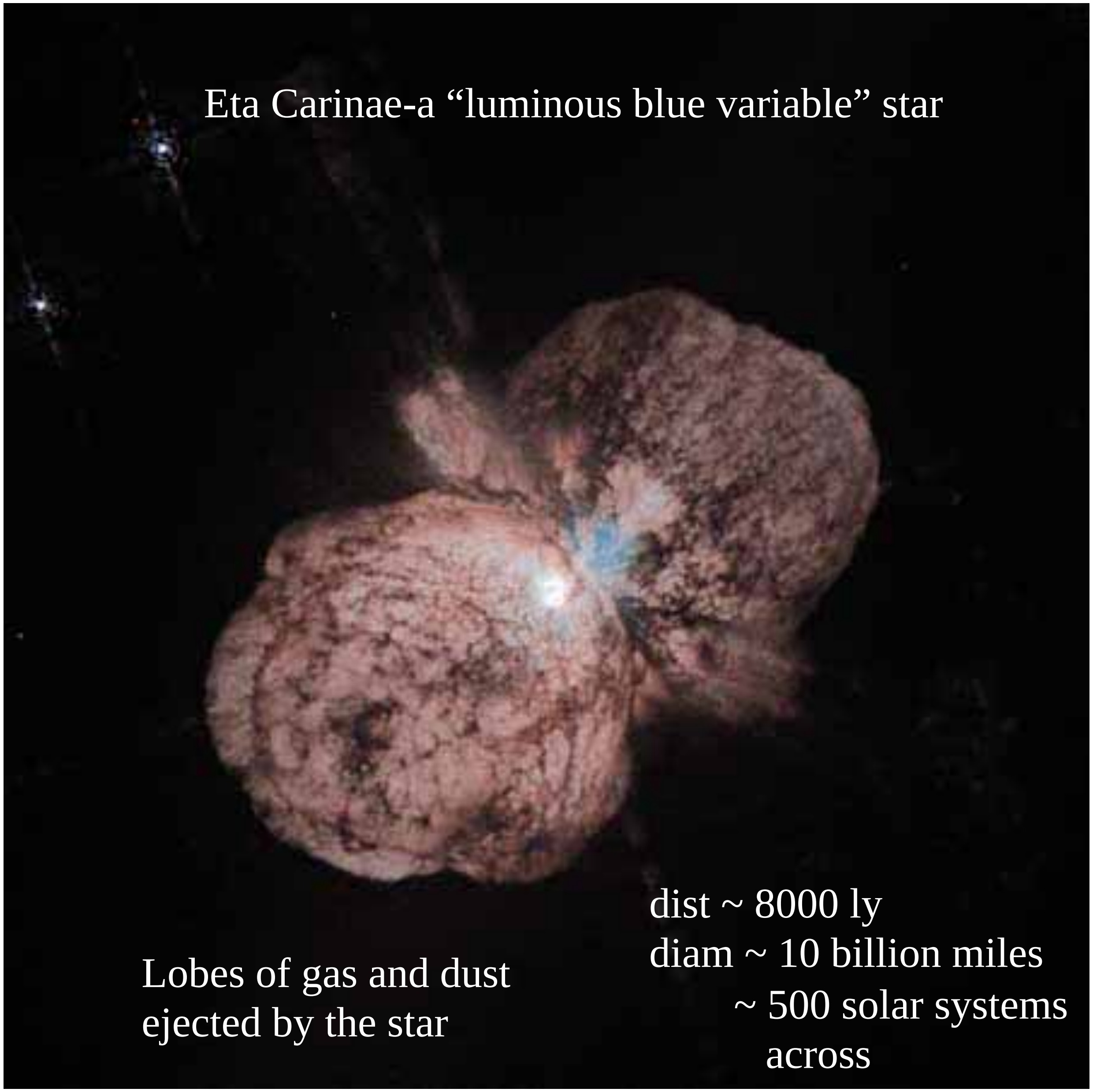
No star can be supported by 100% radiation pressure:

$$P_c = \frac{GM\rho}{2R} \approx \frac{1}{3}\, aT^4 \quad \text{if supported by } P_{radiation}$$

$$\text{but} \quad \rho \sim \left(\frac{3M}{4\,\pi R^3}\right) \Rightarrow \frac{3GM^2}{8\pi R^4} \sim \frac{1}{3}\, aT^4$$

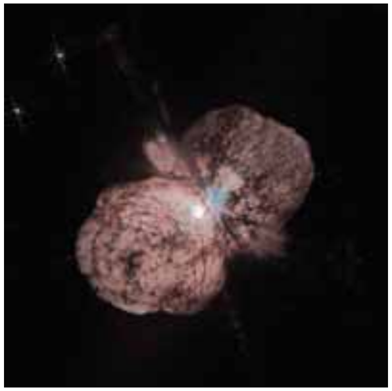
$$\text{so for a fixed M,} \;\; T^4 \propto \left(\frac{1}{R}\right)^4 \propto \rho^{4/3}$$

$P \propto \rho^{4/3}$ which is known to have no stable solution



Most luminous star in our galaxy (that we can study well), several million times more luminous than the sun, bigger than the solar system.

Peculiar star *Eta Carinae* in Carina



1677 – discovered by Edmond Halley – 4^{th} magnitude star
1730 – brightness had reached 2^{nd} magnitude
1801 – brightened again then faded back to 4^{th} magnitude by 1811
1820 – began to brighten again
1822 – reached 2^{nd} magnitude
1827 – reached 1^{st} magnitude began to fade back to 2^{nd} magnitude for about 5 years, then rose to magnitude 0 faded slightly then rose again
1843, April – magnitude -0.8 second brightest star in sky after Sirius, then faded continuously
1868 – became invisible
1900 – had faded to 8^{th} magnitude, stayed there til 1941, then began to brighten again
1953 – 7^{th} magnitude
early 1990's – 6^{th} magnitude
1998-99 – brightened by a factor of 2

Eta Carinae is about 8,000 light years away and one of the most massive stars in the sky (120 to 150 times the mass of the sun).99% of its luminosity is in the infrared.
Probably a supernova in the next 100,000 years, maybe sooner.

Observations suggest a cutoff around 150 $M_\odot$.
Controversial claims of heaver stars come and go.

Star Name	Mass (solar masses)
R136a1	265?
WR101e	150 - 160
HD 269810	150
Peony Nebula Star	150
LBV 1806 - 20	130

July 2010. R136a1 265 $M_\odot$ - controversial

Overview of Evolution (150 > M > 8 Solar Masses)

- Above 8 solar masses, stars ignite carbon burning stably after helium depletion. They avoid becoming degenerate in their centers and go on to burn heavier fuels culminating in the production of an iron core

- Such massive stars have very high luminosities and short lives. They are all (presently) of Population I.

- During the red giant stage the very high luminosities of these stars (and their large radii) imply that the surface layers are very loosely bound. Extensive mass loss occurs

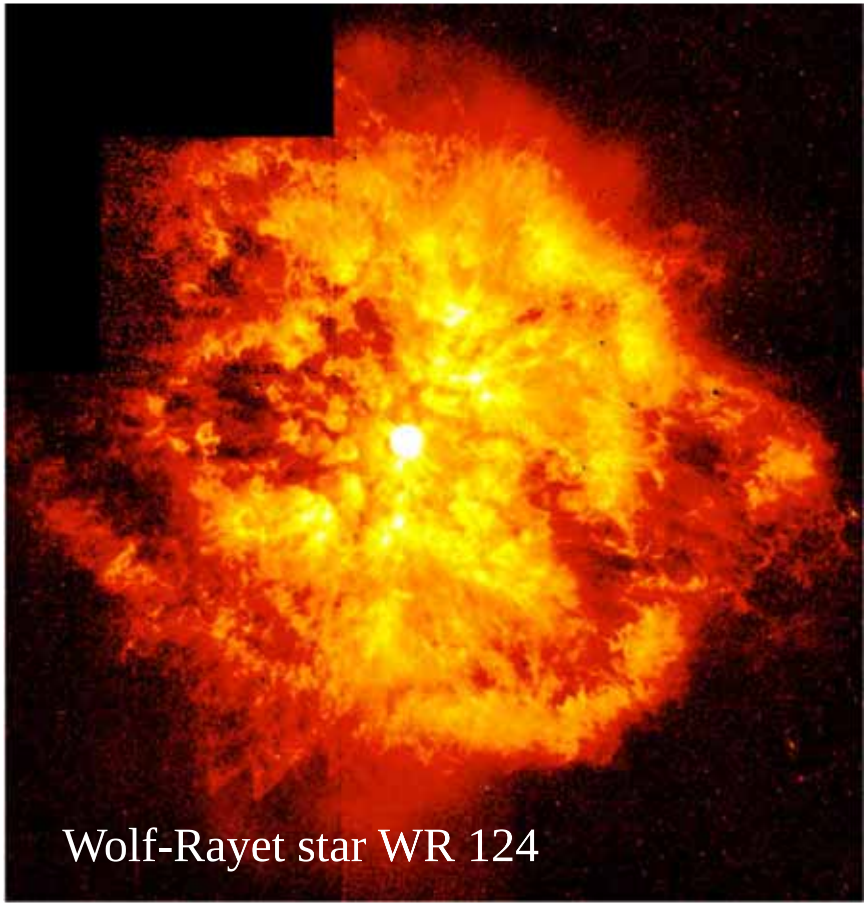
- For stars above about 40 solar masses the entire hydrogen envelope is lost during helium burning. The star becomes a Wolf-Rayet star and mass loss continues at a rapid pace.

Wolf-Rayet stars are massive stars undergoing strong mass loss after leaving the main sequence

- $M > 40\ M_\odot$
- mass is ejected at a rate of $\sim 10^{-6}\ M_\odot$/yr (compare with the Sun: $10^{-14}\ M_\odot$/yr)

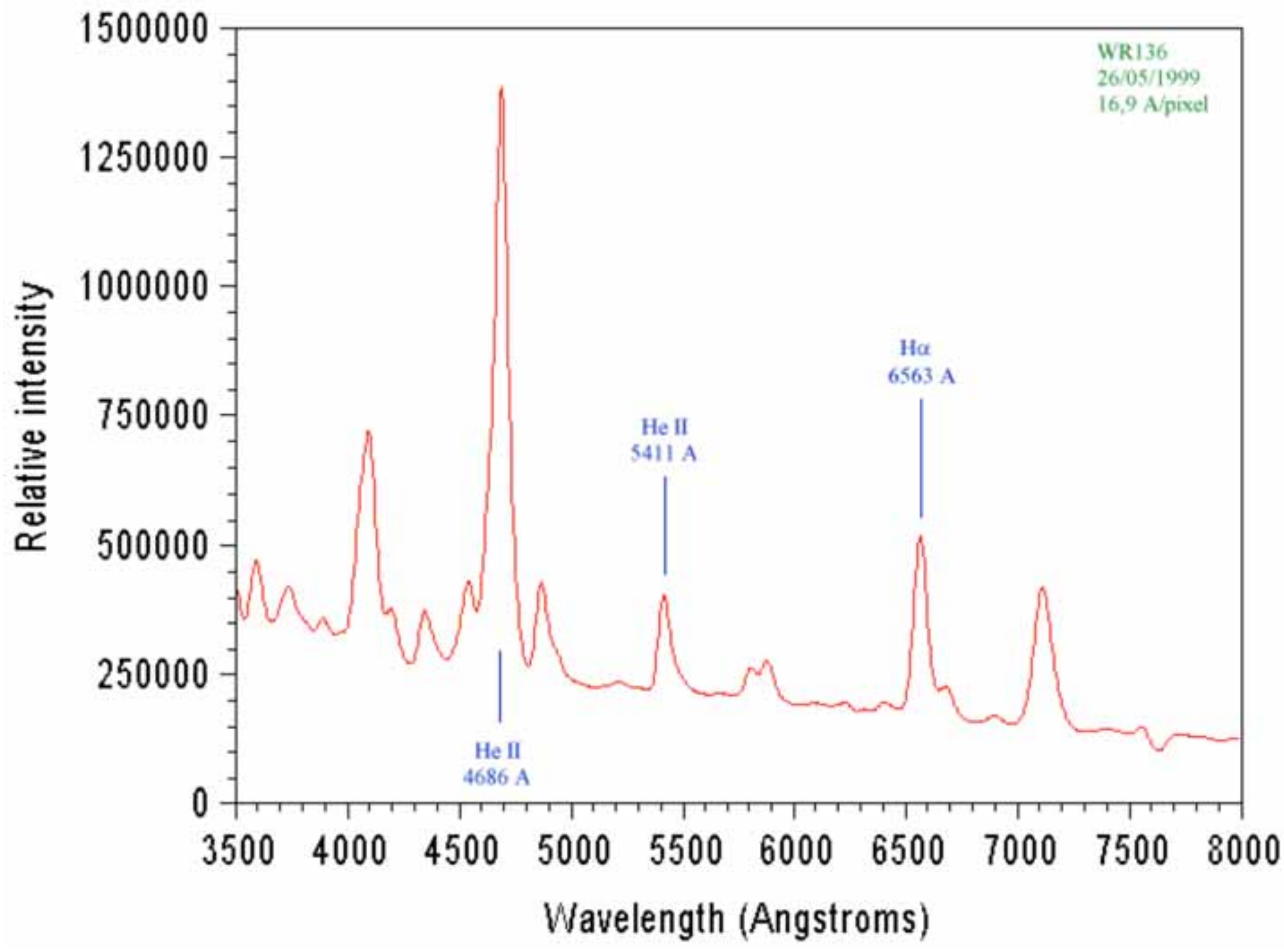
What is the spectral type of a Wolf-Rayet star?

The expanding shells of gas are heated up by radiation from the star. What would a spectrum of the star and gas look like?

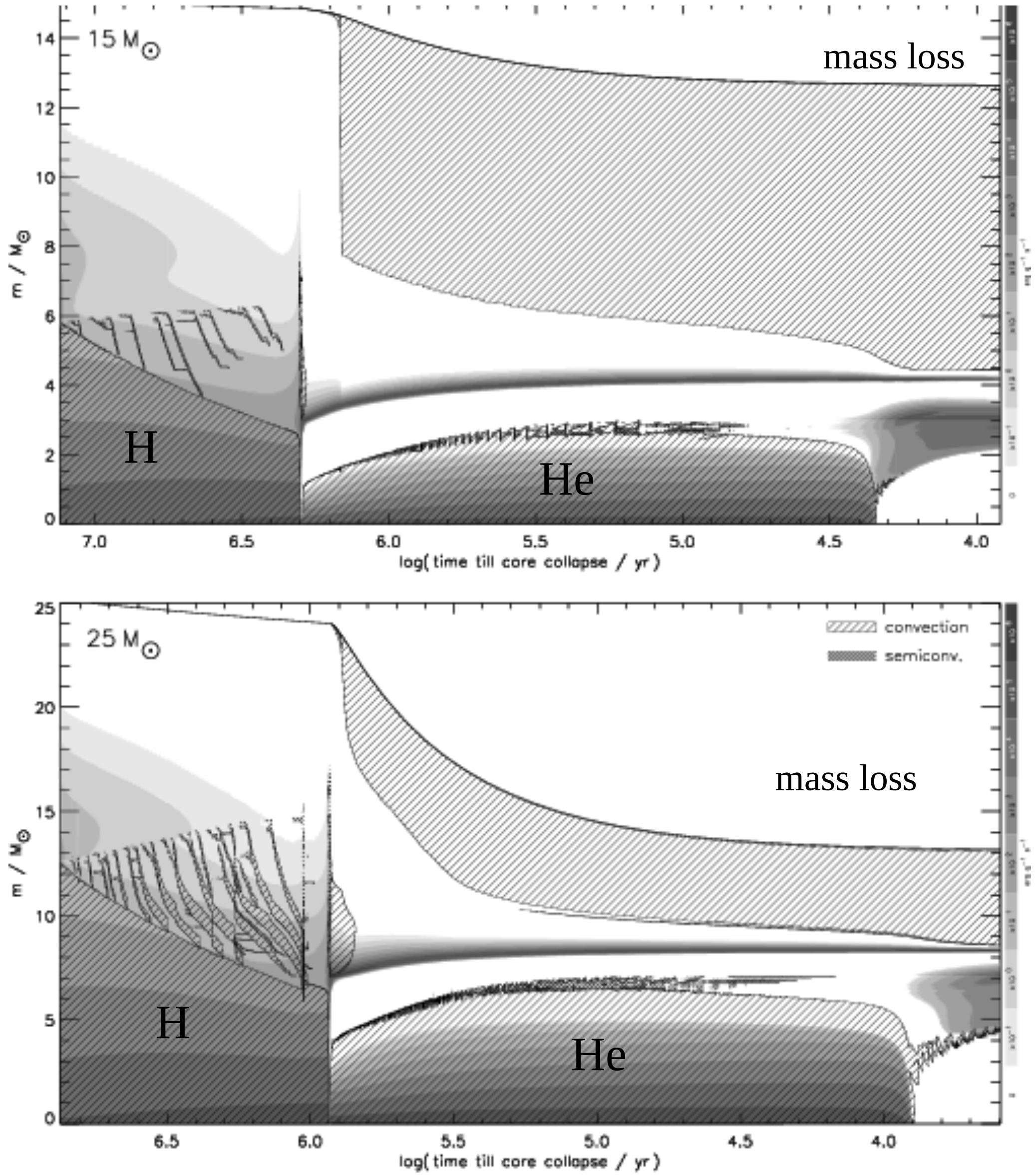


Spectrum of a Wolf-Rayet star

strong emission lines are produced by the hot gas surrounding the star



Convective history 15 $M_\odot$ and 25 $M_\odot$ stars



note: radiate surfaces and convective centers on the main sequence. Time axis is log time until death as a supernova.

Above about 40 solar masses, everything outside the helium core is lost. This makes a Wolf-Rayet star.

Overview of Evolution (150 > M > 8 Solar Masses)

- On the main sequence such massive stars have convective cores and are powered by the CNO cycle. Their surfaces are not convective. After burning hydrogen they ignite helium burning non-degenerately (no "helium flash")

- Evolution beyond helium burning is greatly accelerated by thermal neutrino losses, especially from electron-positron pair annihilation.

- The massive stars that keep part of their hydrogen envelope become Type II supernovae. Those that lose their envelope (either in binaries or single stars above 40 solar masses) become Type Ib or Ic supernovae

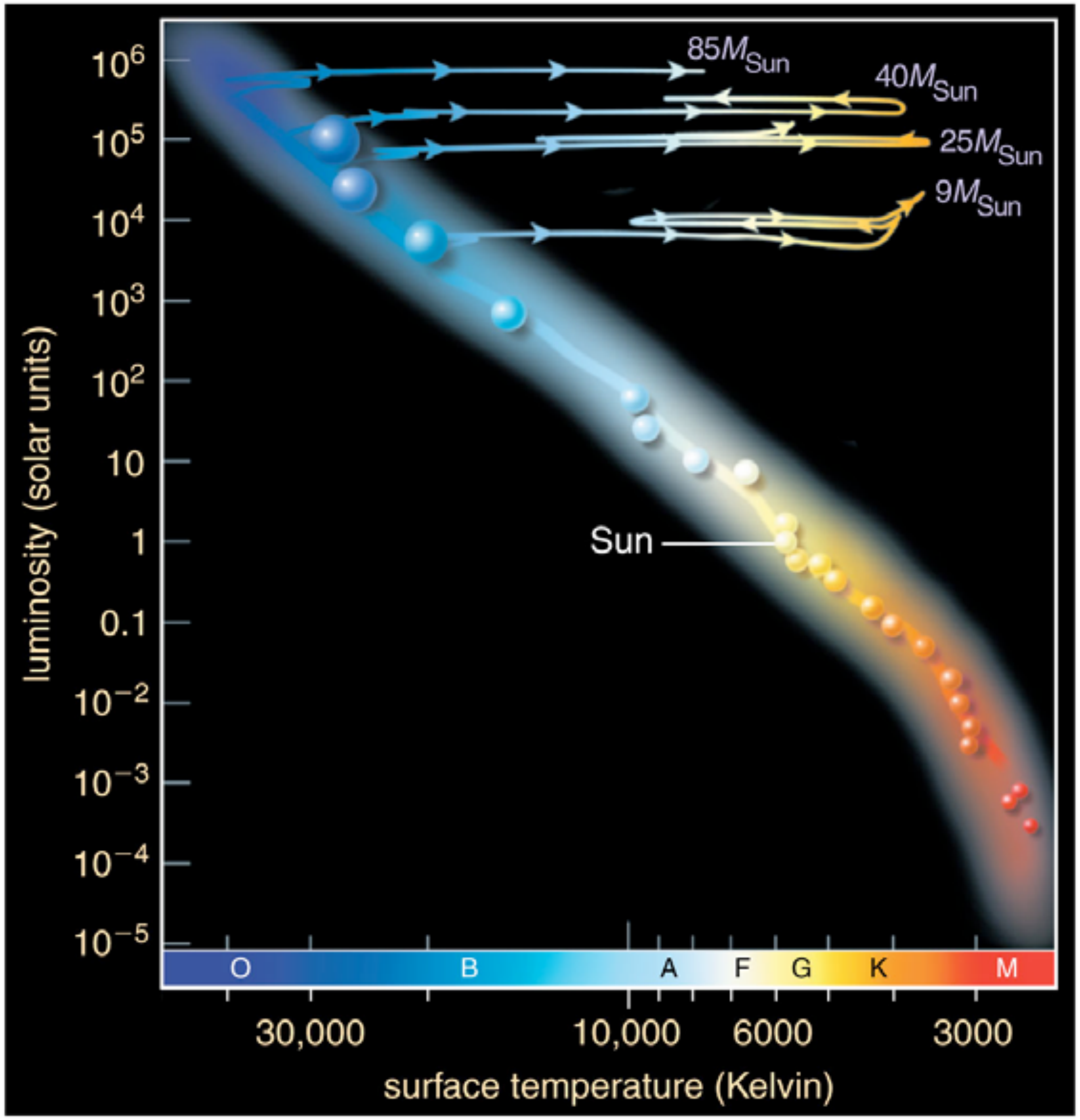
Evolution of High-Mass Stars on the HR diagram

After leaving the MS:

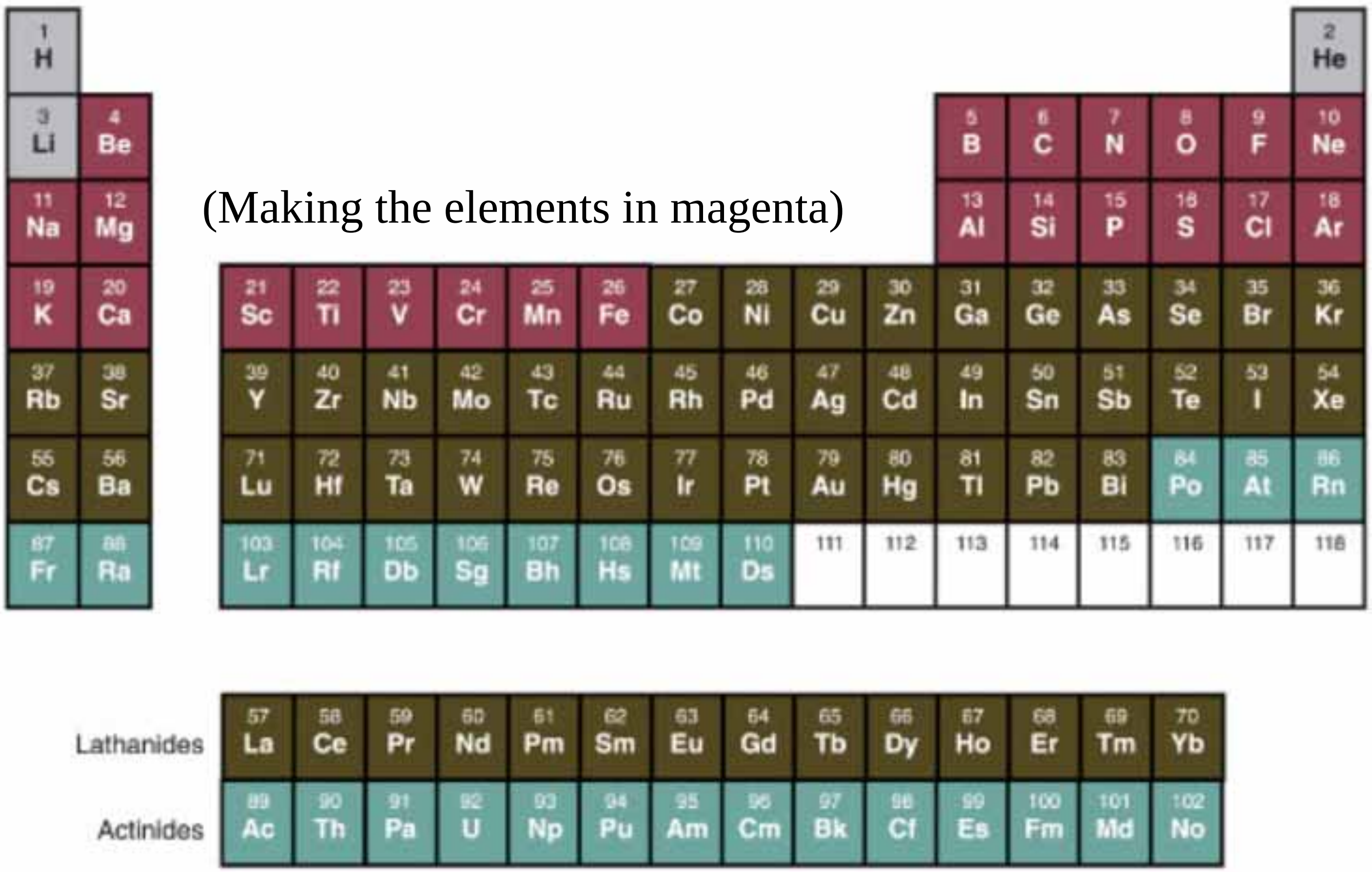
Massive stars shine with ~constant luminosity

Their radii grow and shrink repeatedly (how does color/temp change?)

Fusion of heavier elements (beyond He) occurs within the star



Post-Helium Burning Evolution



(Making the elements in magenta)

Overview of nuclear fusion in massive stars: post-helium burning

- Massive stars ($M > 8\ M_\odot$) are able to fuse elements heavier than helium (all the way to iron) to generate energy

- The products of one type of fusion become the reactants of the next fusion process
    - e.g., carbon → neon, followed by neon → oxygen

- Each successive fusion process requires higher central temperatures and densities in order to operate

- Each successive fuel source releases less energy than the previous fuel, meaning the star burns through it faster
    - Also, lots of neutrinos are produced and carry away large amounts of energy, shortening the burning timescale even more

SUMMARY

Advanced Nuclear Burning Stages

(e.g., 20 solar masses)

Fuel	Main Product	Secondary Products	Temp (10^9 K)	Time (yr)
H	He	^{14}N	0.02	10^7
He	C,O	^{18}O, ^{22}Ne s- process	0.2	10^6
C	Ne, Mg	Na	0.8	10^3
Ne	O, Mg	Al, P	1.5	3
O	Si, S	Cl, Ar K, Ca	2.0	0.8
Si	Fe	Ti, V, Cr Mn, Co, Ni	3.5	1 week

Lifetimes decrease dramatically

Massive stars are the ultimate "recyclers"

Why the big speed up?
Pair Neutrino Losses

After helium burning the core contracts and the temperature rises. The most abundant fuel with the lowest charge is carbon (^{12}C). In order to get two carbons to fuse, a temperature of almost a billion K is required (actually 0.8 billion).

At such high temperatures, a new energy loss mechanism comes into play: neutrinos.

$$\text{Gamma rays } (\gamma) \rightleftarrows e^+ + e^-$$

$$\text{Very rarely though} \quad e^+ + e^- \rightarrow \nu_e + \bar{\nu}_e$$

For $T \sim 10^9$ K, $kT = 86$ keV

$m_e c^2 = 511$ keV

number $e^+ \sim$ number $e^- \sim T^4$

The neutrinos carry away energy very efficiently, since they can pass through the star without interacting

Because the number of electron-positron pairs is very sensitive to the temperature, the energy loss rate due to neutrino losses also depends on a high power of the temperature.

For a temperatures over about 2×10^9 K

$$\varepsilon_{\nu,\mathrm{pair}} \approx -\frac{2\times10^{15}}{\rho}\left(\frac{T}{10^9\,K}\right)^9 \ \mathrm{erg\ g^{-1}\ s^{-1}}$$

For carbon burning and other later burning stages, these losses greatly exceed those due to radiative diffusion and convection.

Because the amount of energy released by each stage is roughly constant, the lifetime at each stage goes down very roughly as $1/T^9$. A higher T is required to burn each fuel.

CARBON BURNING

At a temperature $T \approx 8 \times 10^8$ K and a density $\rho \approx 10^5$ g cm^{-3}, carbon fusion provides energy at a rate that balances losses due to neutrinos.

A little bit of extra energy powers convection and keeps the core hot. Simply carbon $\rightarrow$ neon and magnesium

but in greater detail, the chief reaction is the fusion of two ^{12}C nuclei to produce isotopes of neon, sodium and magnesium

$$^{12}C + {}^{12}C \rightarrow {}^{23}Na + p + 2.24 \text{ MeV}$$

$$^{12}C + {}^{12}C \rightarrow {}^{20}Ne + \alpha + 4.62 \text{ MeV} \qquad (\alpha \equiv {}^4\text{ He})$$

$$^{12}C + {}^{12}C \rightarrow {}^{23}Mg + n - 2.63 \text{ MeV} \qquad \text{(rarely)}$$

CARBON BURNING

The neutrons, protons and alpha-particles (helium nuclei) react with other species that are there so that following the composition becomes complicated (but calculable)

$$^{23}\text{Na} + \text{p} \rightarrow {}^{24}\text{Mg} + \gamma \qquad {}^{23}\text{Na} + \alpha \rightarrow {}^{27}\text{Al} + \gamma$$

$$^{20}\text{Ne} + \alpha \rightarrow {}^{24}\text{Mg} + \gamma \qquad {}^{23}\text{Mg} + \text{n} \rightarrow {}^{24}\text{Mg} + \gamma$$

$$^{24}\text{Mg} + \text{n} \rightarrow {}^{25}\text{Mg} + \gamma \qquad \textit{etc.}$$

The net result is that $4 \times 10^{17}\ \Delta X_{12}$ erg g^{-1} are released and the most abundant isotopes of neon, sodium, magnesium and aluminum are created. Oxygen also survives with a slightly increased abundance. $\Delta X_{12} \approx 0.2$

Note the gradual decrease in energy yield from 6×10^{18} erg g^{-1} for hydrogen burning to about 1×10^{18} erg g^{-1} for helium burning to a lower value for carbon burning.

OXYGEN BURNING

- Similar to carbon burning; at $T \sim 2.0 \times 10^9$K, $\rho \sim 10^6$ g cm^{-3}

$$^{16}\mathrm{O} + {}^{16}\mathrm{O} \rightarrow {}^{28}\mathrm{Si} + \alpha$$

$$^{16}\mathrm{O} + {}^{16}\mathrm{O} \rightarrow {}^{31}\mathrm{P} + \mathrm{p}$$

$$^{16}\mathrm{O} + {}^{16}\mathrm{O} \rightarrow {}^{31}\mathrm{S} + \mathrm{n}$$

and a host of secondary reactions

- The net result is

^{16}O, ^{24}Mg $\rightarrow$ abundant isotopes of silicon, sulfur, chlorine, argon potassium and calcium. Most abundant ashes - ^{28}Si and ^{32}S

$$q_{\mathrm{nuc}} \approx 5.0 \times 10^{17} \; \Delta X_{16} \quad \mathrm{erg\ g^{-1}}$$

$$\varepsilon_{nuc} \propto \mathrm{T}^{33}$$

SILICON BURNING

- $T \approx 3.5 \times 10^9$ K, $\rho \approx 10^7$ g cm^{-3}.
- At the end of oxygen burning the lighest element is silicon.
- Nuclear reactions are complicated, but in the end

$$[Si, S, Cl, Ar, K, Ca] \rightarrow [Ti, V, Cr, Mn, Fe, Co, Ni]$$

- The most abundant nucleus produced is ^{56}Fe

$$q_{nuc} = 2 \times 10^{17} \text{ erg g}^{-1}$$

$$\epsilon_{nuc} \propto T^{47}$$

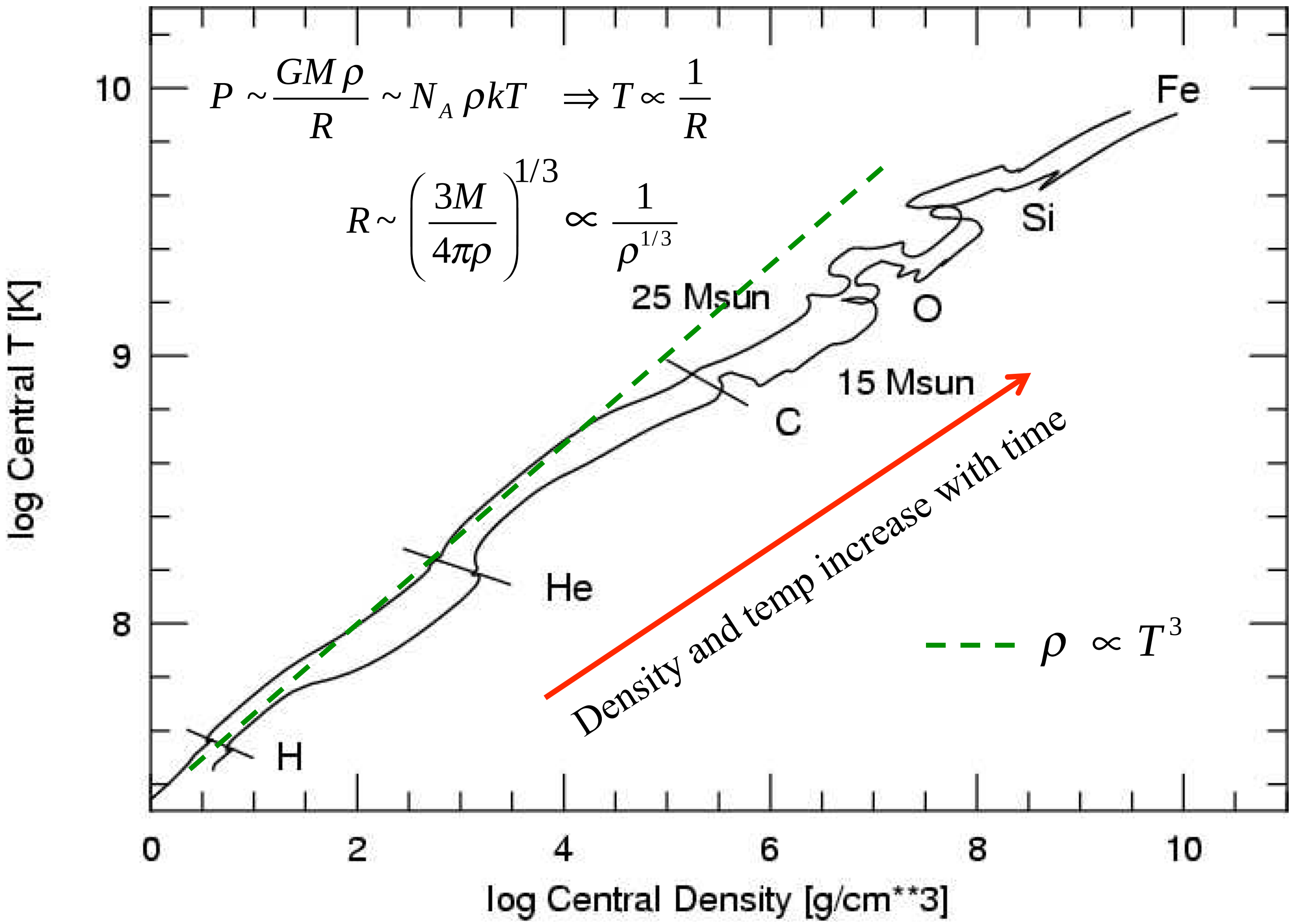
SUMMARY

Advanced Nuclear Burning Stages

(e.g., 20 solar masses)

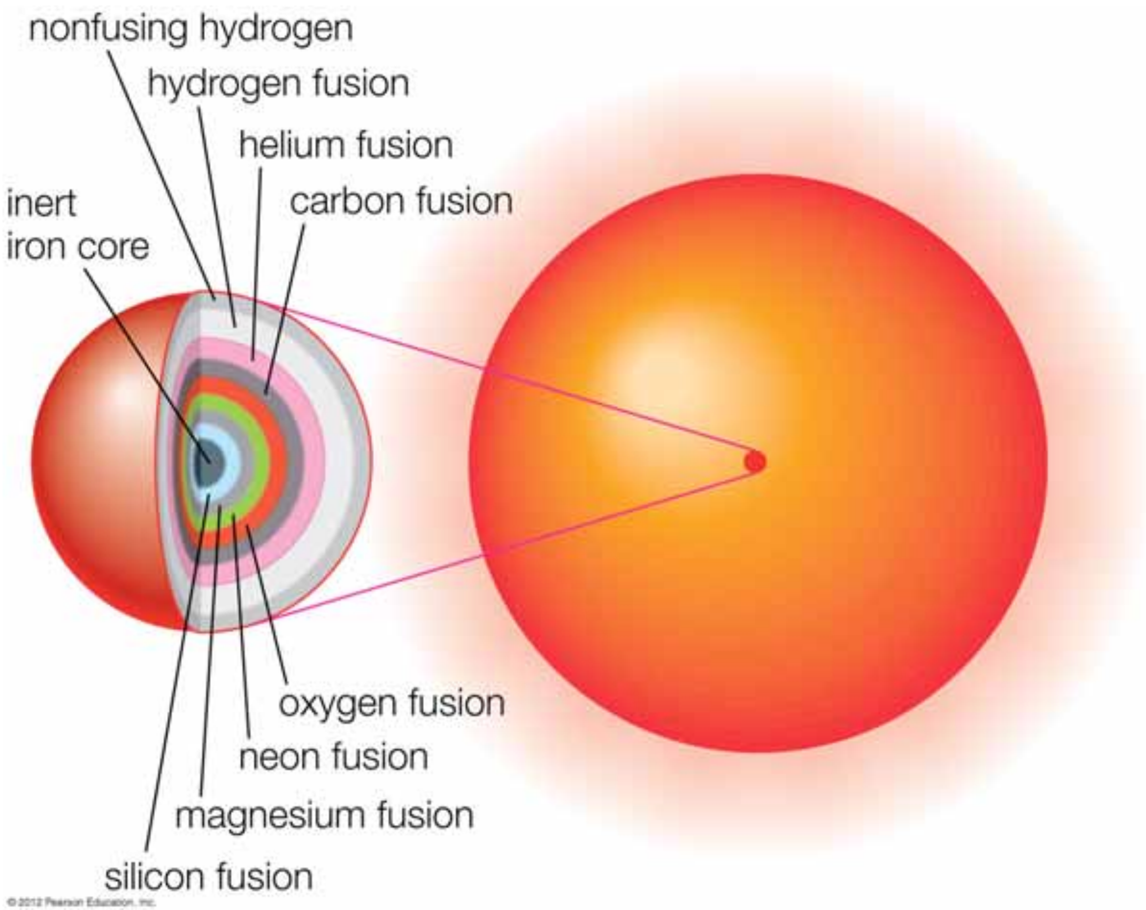
Fuel	Main Product	Secondary Products	Temp (10^9 K)	Time (yr)
H	He	^{14}N	0.02	10^7
He	C,O	^{18}O, ^{22}Ne s- process	0.2	10^6
C	Ne, Mg	Na	0.8	10^3
Ne	O, Mg	Al, P	1.5	3
O	Si, S	Cl, Ar K, Ca	2.0	0.8
Si	Fe	Ti, V, Cr Mn, Co, Ni	3.5	1 week

After each burning stage the core contracts, heats up and ignites another (heavier) fuel



The interior of an evolved high-mass star resembles an onion

A different fusion reaction is occurring in each shell, with heavier elements formed closer to the core



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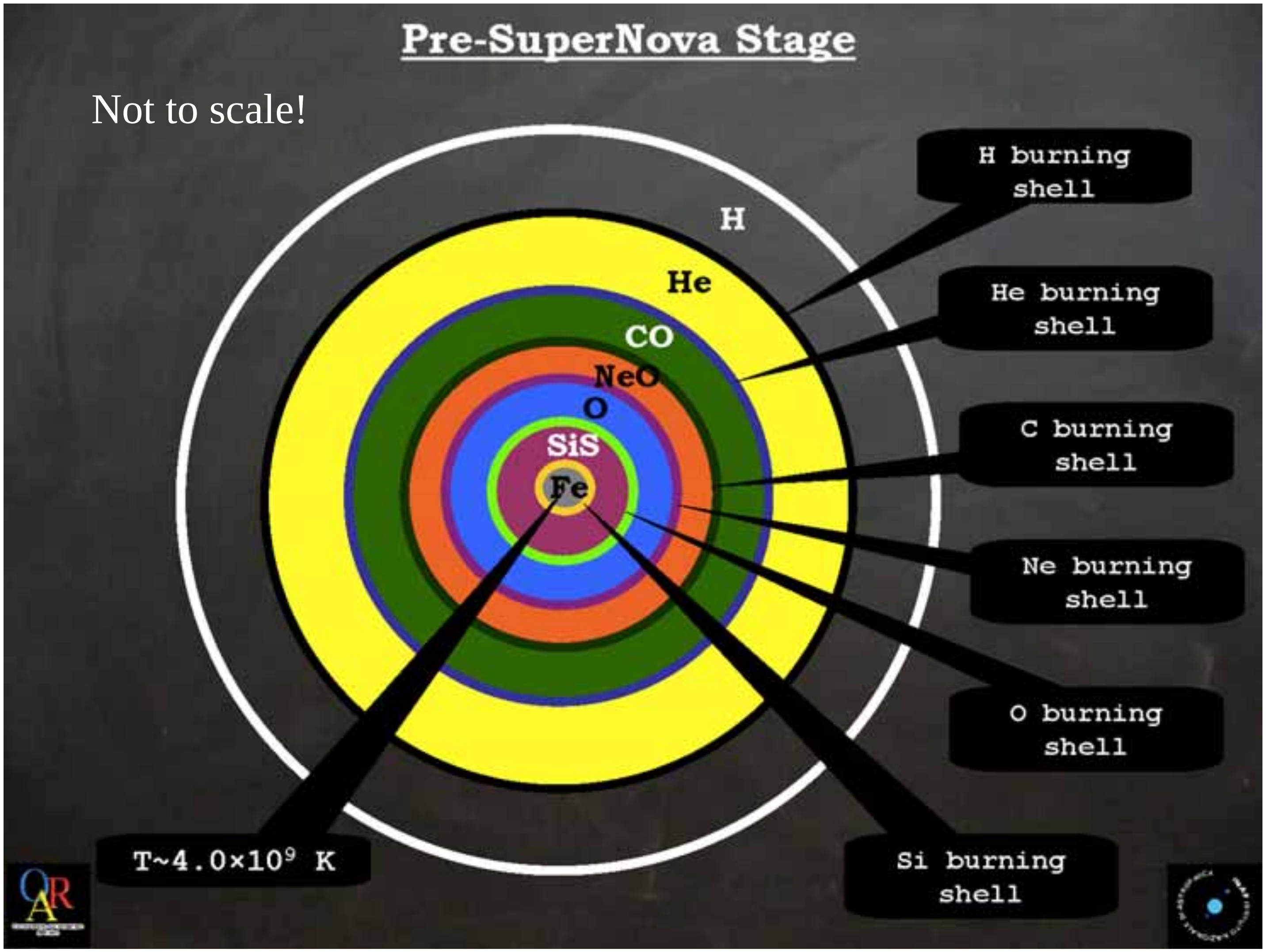
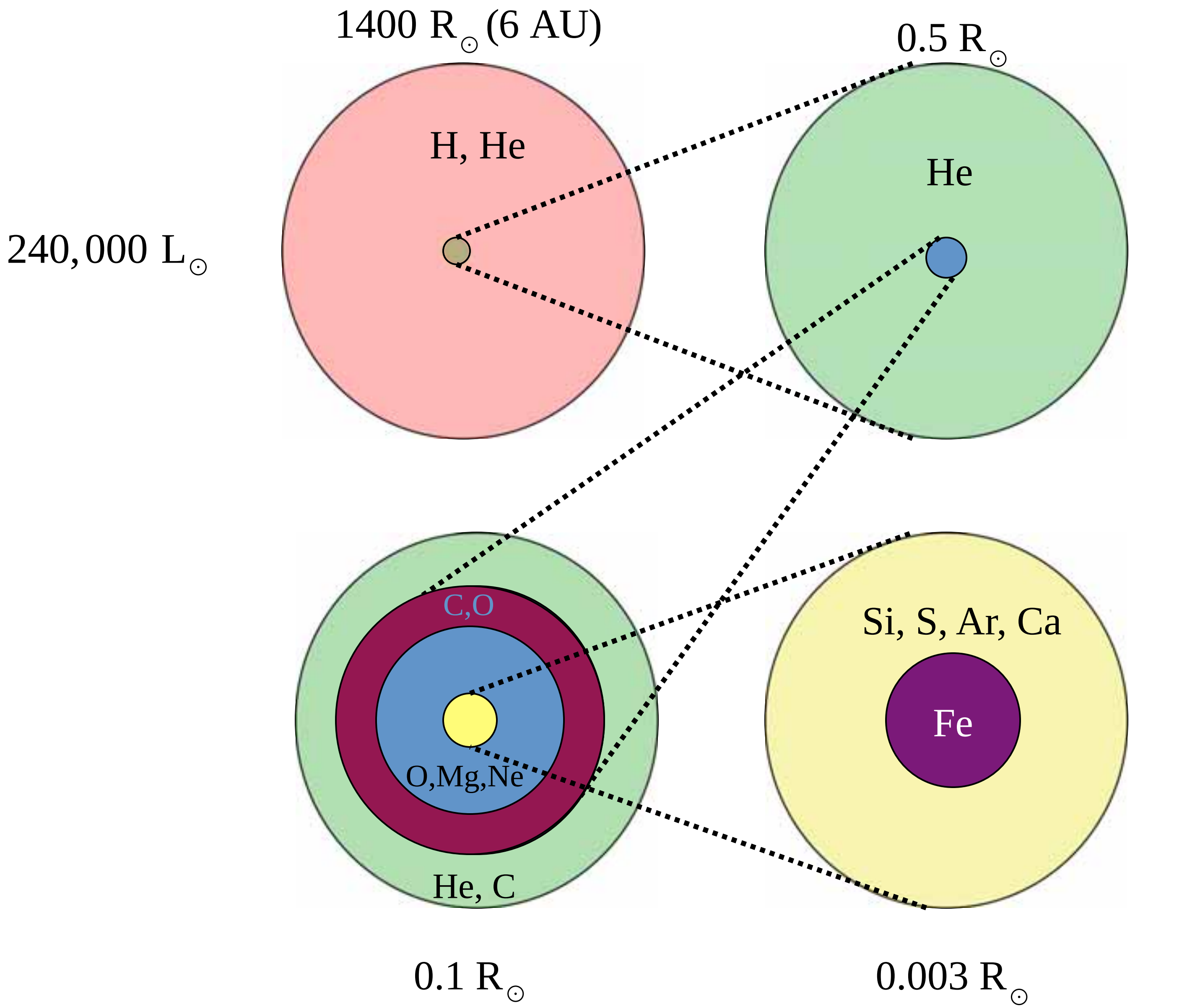


$25\ M_\odot$ Presupernova Star (typical for $9 - 130\ M_\odot$)



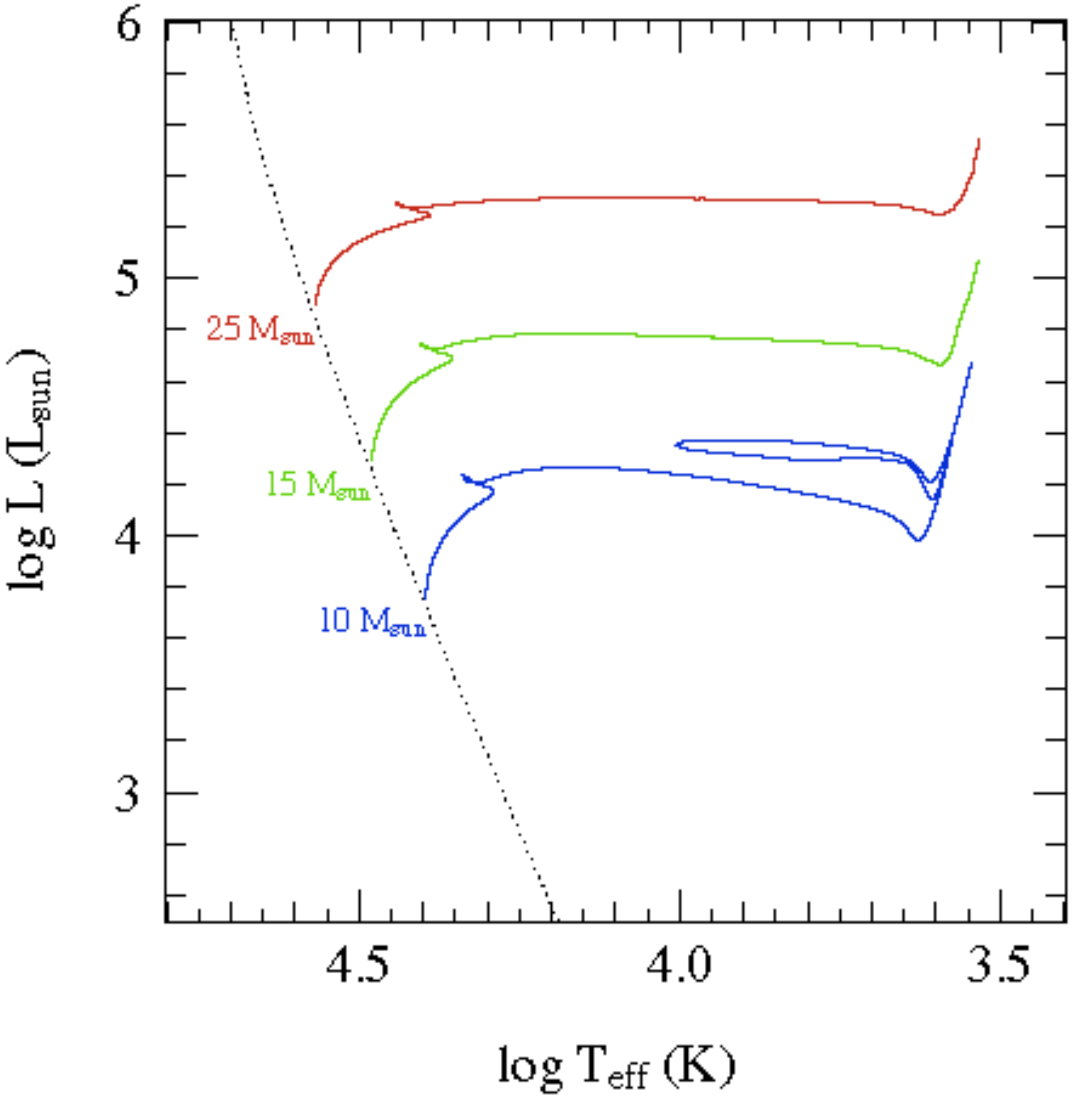
Actual – to scale

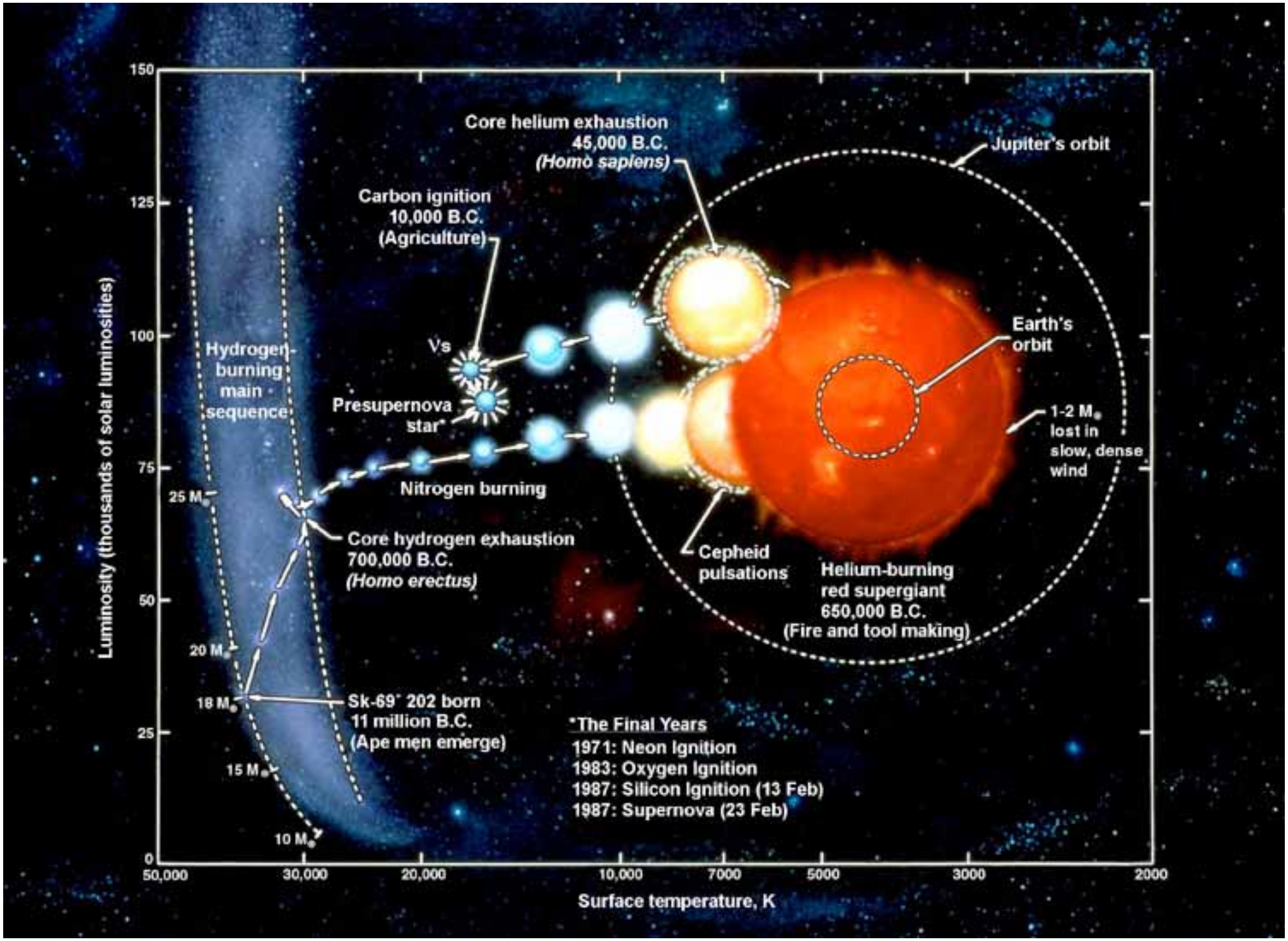
Neutrino emission dominates the energy budget after helium depletion in the center of the star…

Table 1 Burning stages in the evolution of a 20-$M_\odot$ star

Fuel	ρ_c (g cm^{-3})	T_c (10^9 K)	τ (yr)	L_{phot} (erg s^{-1})	L_ν (erg s^{-1})
Hydrogen	5.6(0)	0.040	1.0(7)	2.7(38)	—
Helium	9.4(2)	0.19	9.5(5)	5.3(38)	$<$1.0(36)
Carbon	2.7(5)	0.81	3.0(2)	4.3(38)	7.4(39)
Neon	4.0(6)	1.7	3.8(−1)	4.4(38)	1.2(43)
Oxygen	6.0(6)	2.1	5.0(−1)	4.4(38)	7.4(43)
Silicon	4.9(7)	3.7	2 days	4.4(38)	3.1(45)

In the HR diagram, massive stars evolve at nearly constant luminosity off the main sequence and eventually explode as red or blue supergiants





Most massive stars die as red supergiants. This one made a transition back to the blue just before dying

Critical Masses

$0.08\ M_\odot$

Contracting protostars below this mass do not ignite hydrogen burning on the main sequence. They become brown dwarfs or planets.

$0.50\ M_\odot$

Stars below this mass are completely convective on the main sequence

“ “ “ “ do not ignite helium burning

$2.0\ M_\odot$

Stars below this mass (and above .5) experience the helium core flash

Stars above this mass are powered by the CNO cycle (below by the pp-cycles)

Stars above this mass have convective cores on the main sequence (and radiative surfaces)

$8\ M_\odot$

Stars below this mass do not ignite carbon burning. They end their lives as planetary nebulae and white dwarfs. Stars above this mass make supernovae.

$\sim 150\ M_\odot$

Population I stars much above this mass pulse apart on the main sequence. No heavier stars exist.