Lecture 10

Nucleosynthesis During Helium Burning and the s-Process

A. Thermodynamic Conditions (Massive Stars)

For the most part covered previously
$\tau_{\mathrm{He}} \sim 10^6$ years (+- factor of three)

n=3

$$M_\alpha = 18\,\frac{\sqrt{1-\beta}}{\mu^2\beta^4} \qquad \mu = \frac{4}{3} \text{ (for pure helium)} \qquad \text{solve for } \beta$$

$$\mathrm{T_C} = 4.6\times10^6\,\mu\beta\left(\frac{M_\alpha}{M_\odot}\right)^{2/3}\rho_c^{\ 1/3} \quad \mathrm{K}$$

E.g., $\mathrm{M}_\alpha = 6$ (a 20 $\mathrm{M}_\odot$ main sequence star)

$$\beta = 0.83$$

$$\mathrm{T_c} = 1.7\times10^8\left(\frac{\rho_c}{1000\,\mathrm{g\ cm^{-3}}}\right)^{1/3} \mathrm{K}$$

need temperatures > 10^8 to provide significant energy generation by 3α

So, typical temperatures are 2 x 10^8 K (higher in shell burning later) when densities are over 1000 gm cm^{-3}. As the core evolves the temperature and density go up significantly. Note non-degenerate.

From Schaller et al (1992) $Z = 0.02$ and central helium mass fraction of 50%.

Main sequence Mass	Current Mass	Approximate M_α	$T_c/10^8$	$\rho_c/1000$ g cm^{-3}
12	11.66	3	1.76	1.42
15	14.24	4	1.83	1.12
20	18.10	6	1.92	0.831
25	20.40	8	1.99	0.674
40	20.7	12	2.11	0.470

At 1.9×10^8 K, the temperature sensitivity of the 3α rate is approximately T^{20}.

As discussed earlier during helium burning:

$$\frac{dY_\alpha}{dt} = -3\rho^2 Y_\alpha^3 \lambda_{3\alpha} - Y(^{12}C)Y_\alpha \rho \lambda_{\alpha\gamma}(^{12}C)$$

$$\frac{dY(^{12}C)}{dt} = \rho^2 Y_\alpha^3 \lambda_{3\alpha} - Y(^{12}C)Y_\alpha \rho \lambda_{\alpha\gamma}(^{12}C)$$

$$\frac{dY(^{16}O)}{dt} = Y(^{12}C)Y_\alpha \rho \lambda_{\alpha\gamma}(^{12}C)$$

Coulomb barrier and lack of favorable resonances inhibit alpha capture on ^{16}O.

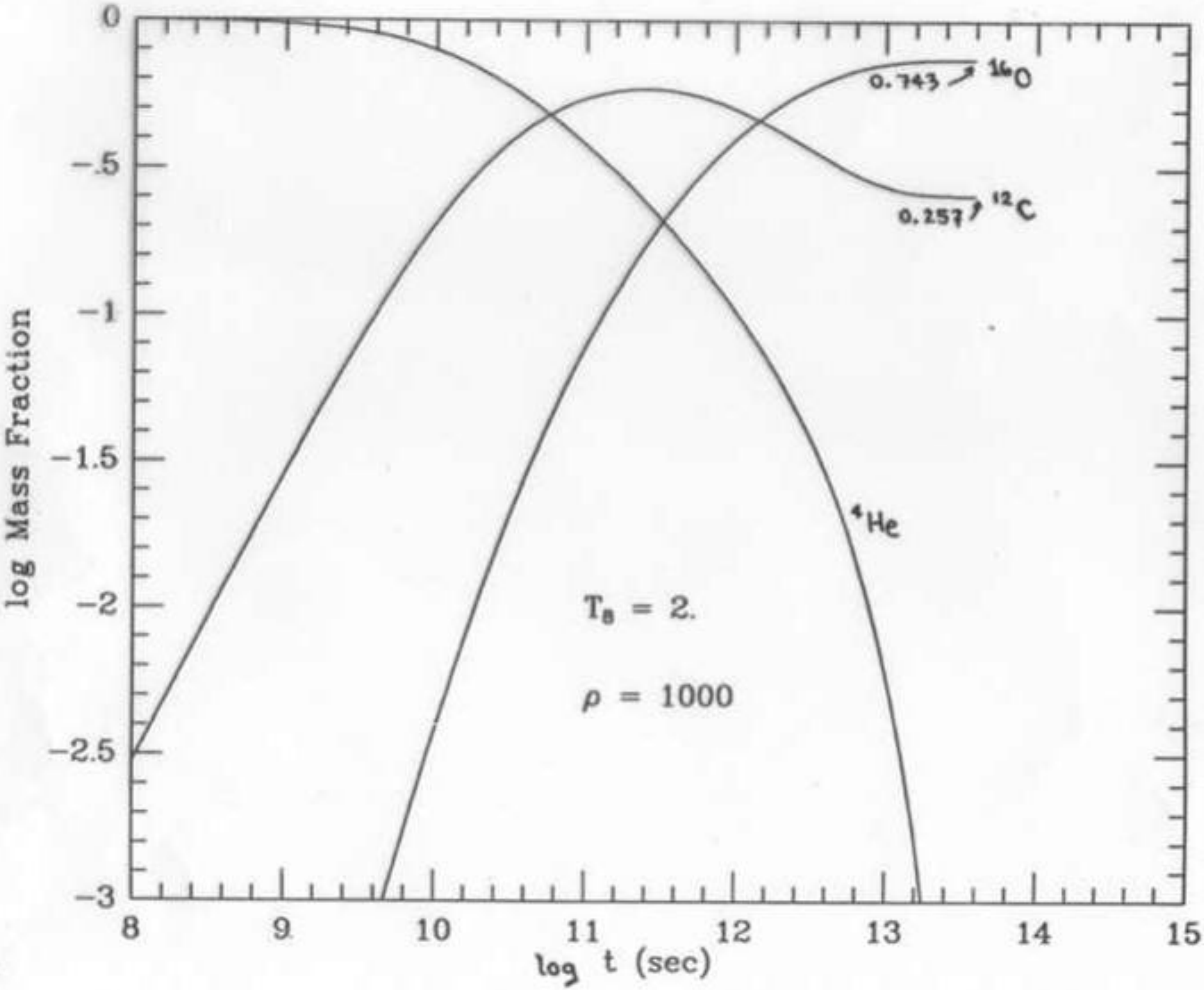
Several general features:

- ^{12}C production favored by large density; oxygen by lower density

- ^{12}C produced early on, ^{16}O later

- last few alpha particles burned most critical in setting ratio $^{12}C/^{16}O$

- Energy generation larger for smaller $^{12}C/^{16}O$.



0
-.5
-1
-1.5
-2
-2.5
-3
8
9
10
11
12
13
14
15
log Mass Fraction
log t (sec)
0.743
16O
0.257
12C
4He
T8 = 2.
ρ = 1000

B1. Principal Nucleosynthesis

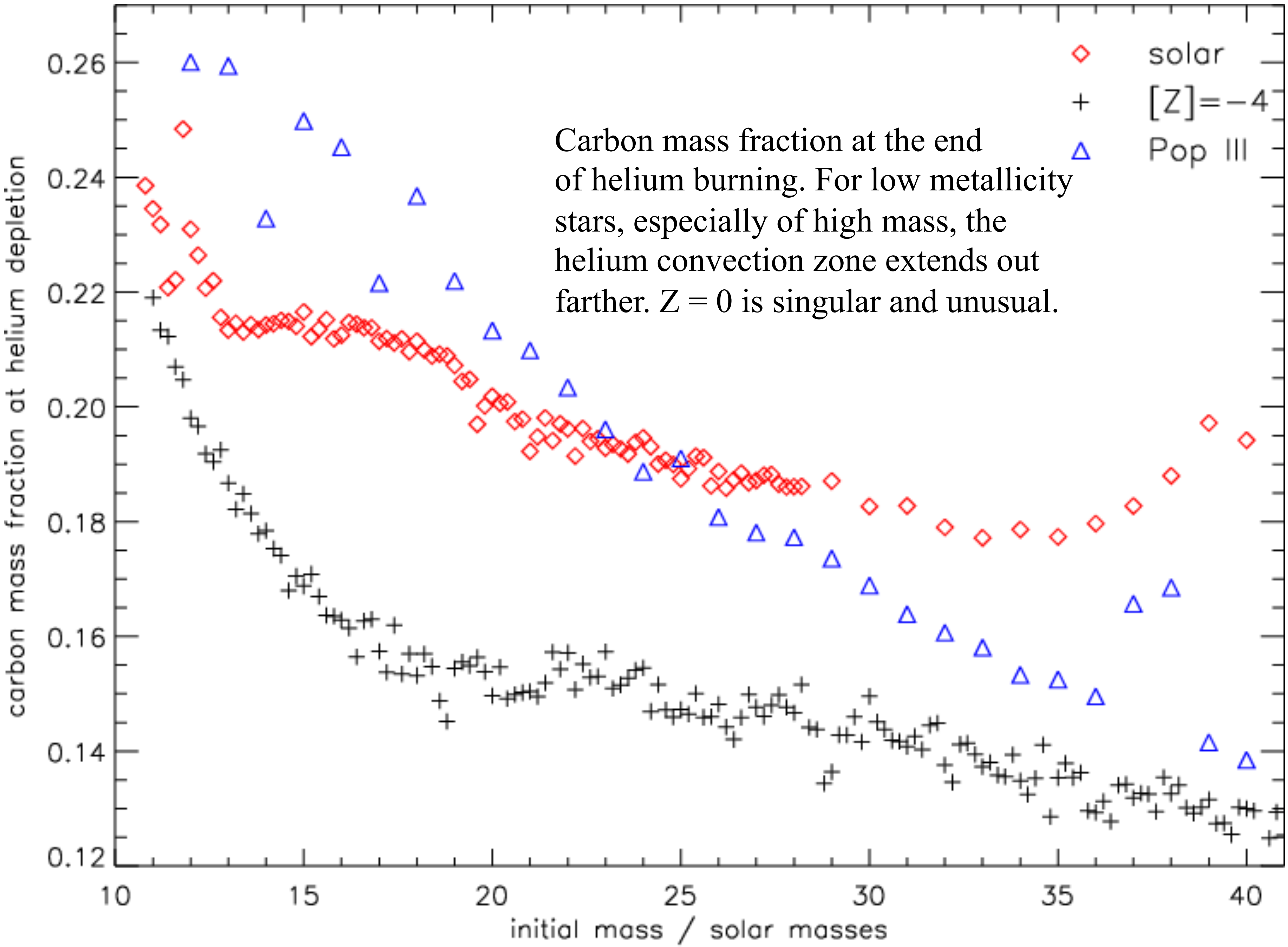
In massive stars after helium burning in the stars center (calculations included semiconvection).

In the sun, $^{12}C/^{16}O = 0.32$

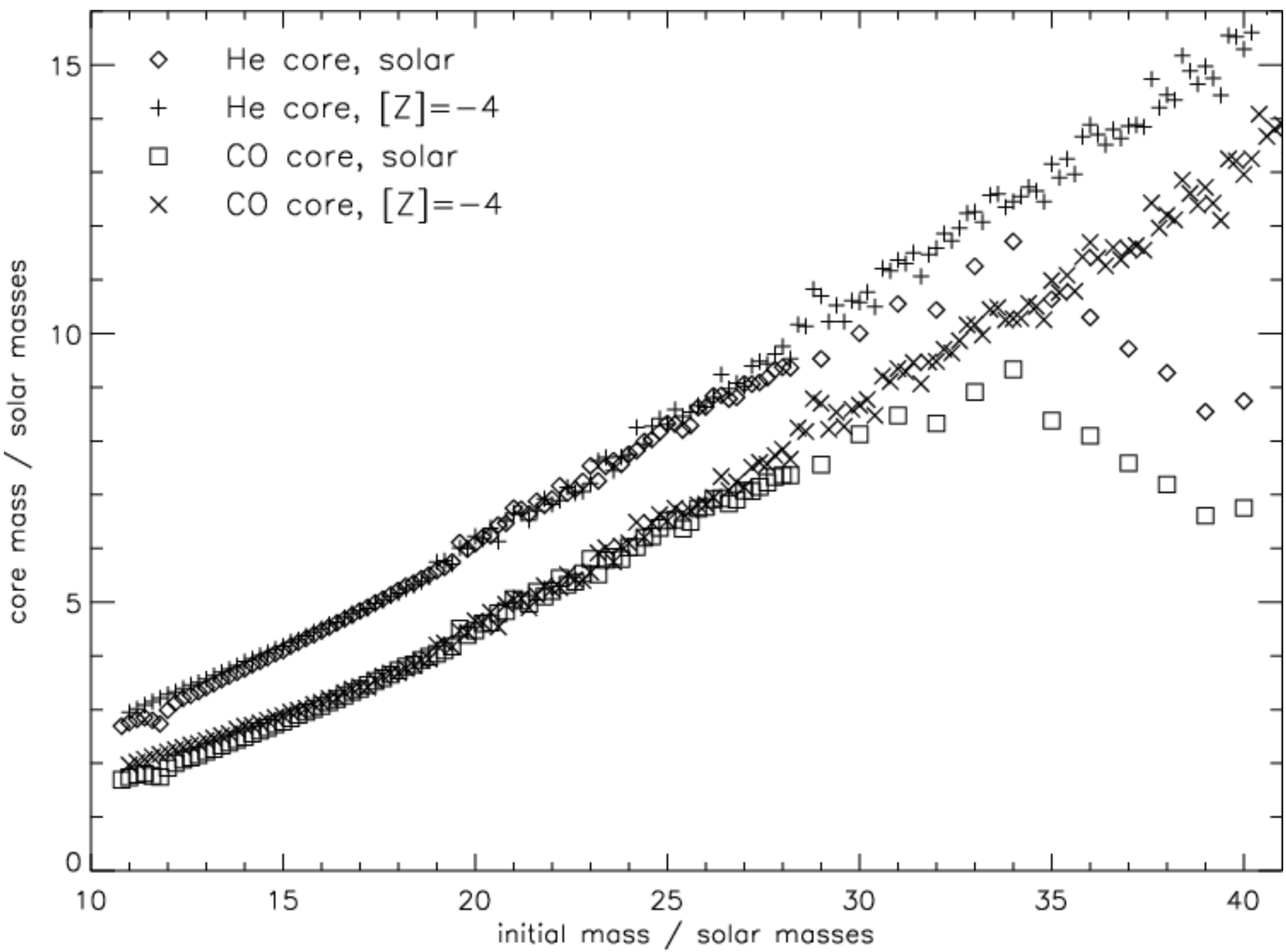
$M/M_\odot$	$X(^{12}C)$	$X(^{16}O)$	$X(^{20}Ne)$
12	0.247	0.727	1.91(-3)
15	0.188	0.785	3.02(-3)
19	0.153	0.815	6.44(-3)
25	0.149	0.813	1.16(-2)
35	0. 147	0.807	1.84(-2)

Buchman $^{12}C(\alpha,\gamma)^{16}O$ multiplied by 1.2.

If the star contains appreciable metals there is, as we shall see also ^{22}Ne and ^{18}O.



Carbon mass fraction at the end of helium burning. For low metallicity stars, especially of high mass, the helium convection zone extends out farther. Z = 0 is singular and unusual.


He core, solar
He core, [Z]=−4
CO core, solar
CO core, [Z]=−4
core mass / solar masses
initial mass / solar masses
0
5
10
15
10
15
20
25
30
35
40

Within uncertainties, helium burning in massive stars (over 8 solar masses) could be the origin in nature of ^{12}C. It is definitely the origin of ^{16}O

Complications:

- If the helium core grows just a little bit towards the end of helium burning, the extra helium convected in greatly decreases the ^{12}C synthesis.

- Mass loss from very massive WR stars can greatly increase the synthesis of both ^{12}C and ^{16}O in stars over 35 solar masses

- The uncertain rate for $^{12}C(\alpha,\gamma)^{16}O$

- $^{12}C/^{16}O$ ratio may be affected by post-helium burning evolution and by black hole formation above some critical main sequence mass. ^{16}O is made in the more massive (massive) stars.

B. Trace Element Nucleosynthesis - Charged Particles

During hydrogen burning CNO $\rightarrow$ ^{14}N so that at helium ignition

$$X(^{14}N) = 14\,(Y_i(^{12}C) + Y_i(^{14}N) + Y_i(^{16}O))$$
$$= 0.013\,(Z/Z_\odot)$$

Early during helium burning

$$\boxed{^{14}N(\alpha,\gamma)^{18}F(e^+\nu)^{18}O}$$

NB. From this point onwards have a net neutron excess

$$\boxed{\begin{aligned}\eta &= 1 - 2Y_e \\ &= \sum (N_i - Z_i)\,Y_i\end{aligned}}$$

$-1 \leq \eta \leq 1$

$\eta = 0$ if $Z = N$

Before the above reaction the composition was almost entirely ^{4}He and ^{14}N, hence $\eta \approx 0$ (actually a small posive value exists because of ^{56}Fe and the like).

After this reaction

$$\boxed{\eta = 0.0019\,\frac{Z}{Z_{\odot}}}$$

During helium core burning, ^{18}O is later mostly destroyed by $^{18}\mathrm{O}(\alpha,\gamma)^{22}\mathrm{Ne}$.

During helium shell burning which does not go to completion in massive stars, much of ^{18}O remains undestroyed and this is the source of ^{18}O in nature. Convection helps to preserve it.

Lifetimes (yr) $X_\alpha = 1/2$ $\rho = 1000.$

	$T_8 = 1.8$	$T_8 = 2$	
$^{16}O(\alpha,\gamma)$	3.5 (10)	2.3 (9)	*survives*
$^{18}O(\alpha,\gamma)$	2700	98	*destroyed unless preserved by convection*
$^{14}N(\alpha,\gamma)$	15	1.1	*destroyed*
$^{22}Ne(\alpha,n)$ + (α,γ)	4.1 (6)	2.2 (5)	*partly destroyed*

So one expects that, depending on mass, some but not all of the ^{22}Ne will burn towards the end of helium burning when the temperature goes up.

The following table gives the temperature at the center of the given model and the mass fractions of ^{22}Ne, ^{25}Mg, and ^{26}Mg each multiplied by 1000, when the helium mass fraction is 1% and zero

Woosley, and Heger (2007)

M	T_C	^{22}Ne	^{25}Mg	^{26}Mg
12	2.42	13.4	0.51	0.61
		12.3	1.17	1.05
15	2.54	12.7	0.98	0.91
		11.1	1.99	1.80
19	2.64	11.5	1.73	1.54
		9.4	2.90	2.87
25	2.75	9.8	2.67	2.59
		6.96	4.05	4.54
35	2.86	7.37	3.87	4.22
		4.41	5.18	6.39

The remainder of the ^{22}Ne will burn early during carbon burning, but then there will be more abundant "neutron poisons".

These numbers are quite sensitive to the uncertain reaction rate for ^{22}Ne(α,n)^{25}Mg and may be lower limits to the ^{22}Ne consumption.

C. The s-Process in Massive Stars

Late during helium burning, when the temperature rises to about 3.0×10^8 K, ^{22}Ne is burned chiefly by the reaction $^{22}\mathrm{Ne}(\alpha,n)^{25}\mathrm{Mg}$ (with some competition from $^{22}\mathrm{Ne}(\alpha,\gamma)^{26}\mathrm{Mg}$).

Where do the neutrons go?

Some go on ^{56}Fe but that fraction is only:

$$f = \frac{\sigma_{56} Y_{56}}{\sum \sigma_i Y_i}$$

^{16}O	$Y_{16} \approx \dfrac{0.5}{16} = 3.1 \times 10^{-2}$	$Y_{16}\,\sigma_{16} \approx 1.2 \times 10^{-3}$
^{22}Ne	$Y_{22} \approx \dfrac{0.005}{22} = 2.3 \times 10^{-4}$	$Y_{22}\,\sigma_{22} \approx 1.3 \times 10^{-5}$
^{25}Mg	$Y_{25} \approx \dfrac{0.005}{25} = 2 \times 10^{-4}$	$Y_{25}\,\sigma_{25} \approx 1.3 \times 10^{-3}$
^{56}Fe	$Y_{56} \approx \dfrac{0.0013}{56} = 2.3 \times 10^{-5}$	$Y_{56}\,\sigma_{56} \approx 2.7 \times 10^{-4}$

But, $^{17}\mathrm{O}(\alpha,n)^{20}\mathrm{Ne}$ destroys the ^{17}O and restores the neutron

For σ in mb

30 keV neutron capture cross sections

(mostly from Bao et al, ADNDT, 2000)

Nucleus	σ (mb)	Nucleus	σ (mb)	
^{12}C	0.0154	^{54}Fe	27.6	
* ^{16}O	0.038	^{56}Fe	11.7	
^{20}Ne	0.119	^{57}Fe	40.0	
^{22}Ne	0.059	^{58}Fe	12.1	
^{24}Mg	3.3	^{58}Ni	41.0	
^{25}Mg	6.4	^{64}Zn	59	
^{26}Mg	0.126	^{65}Zn	162	
^{28}Si	2.9	^{66}Zn	35	
		^{88}Sr	6.2	(closed shell)

The large cross section of ^{25}Mg is particularly significant since it is made by $^{22}Ne(\alpha,n)$ ^{25}Mg.

* Igashira et al, *ApJL*, **441**, L89, (1995); factor of 200 upwards revision

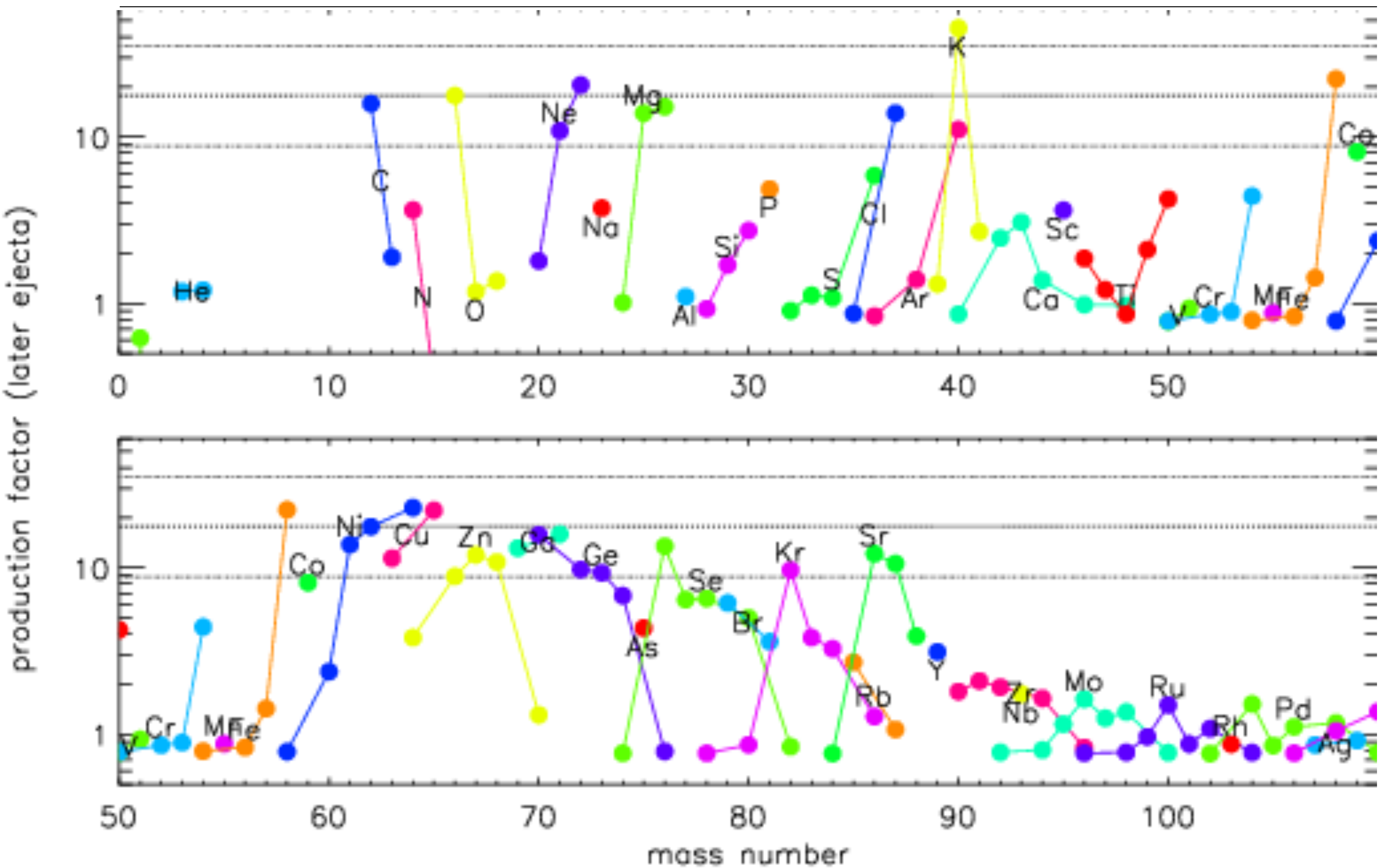
So a fraction $\frac{Y_{56}\sigma_{56}}{Y_{25}\sigma_{25}} \sim 10-20\%$ capture on iron. How many neutrons is this?

$$Y_n / Y_{Fe} \sim Y_{22} / Y_{56} = 42$$

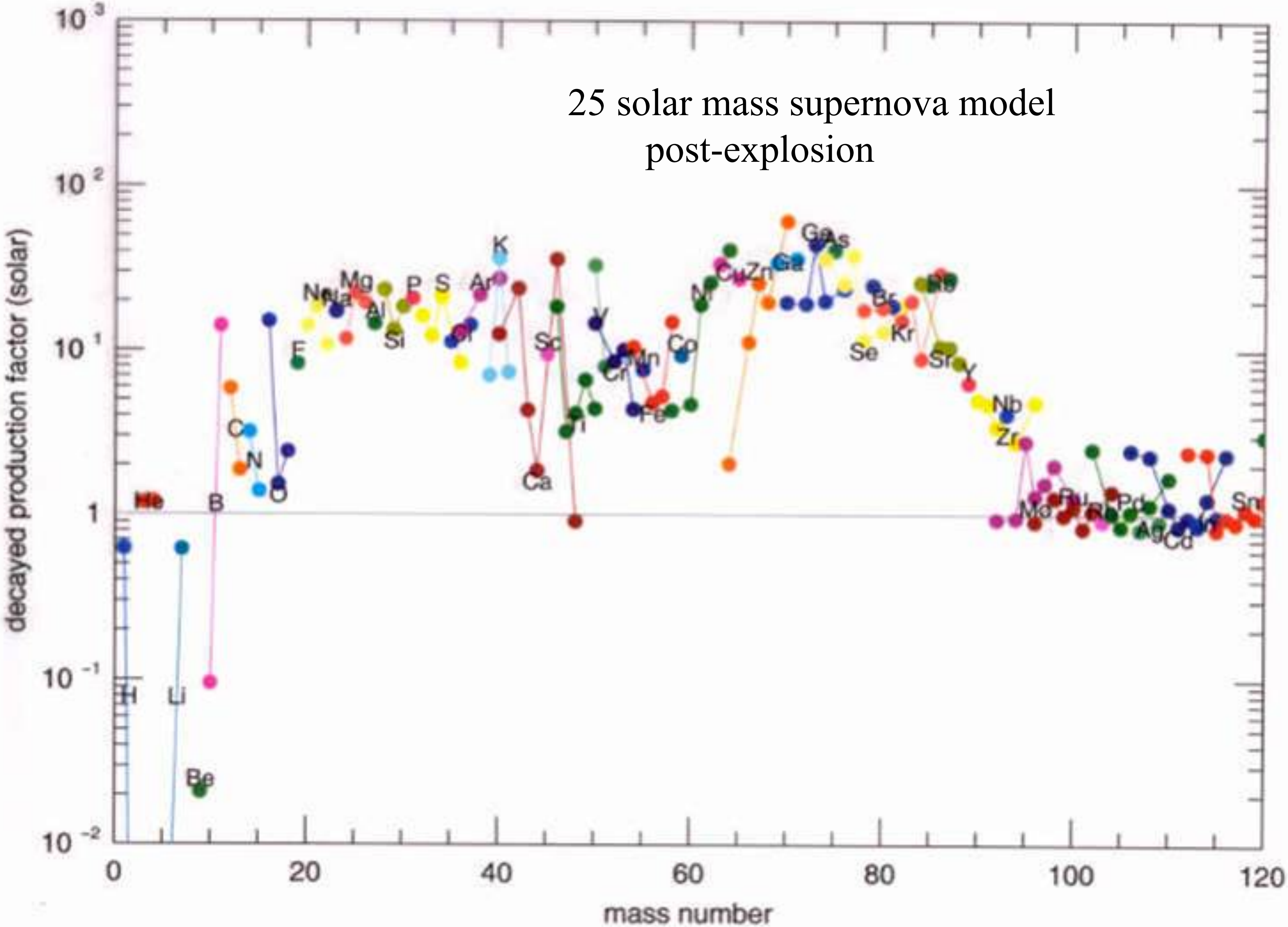
where we have assumed a mass fraction of 0.02 for ^{22}Ne and 0.0013 for ^{56}Fe and that all ^{22}Ne burns by (α,n).

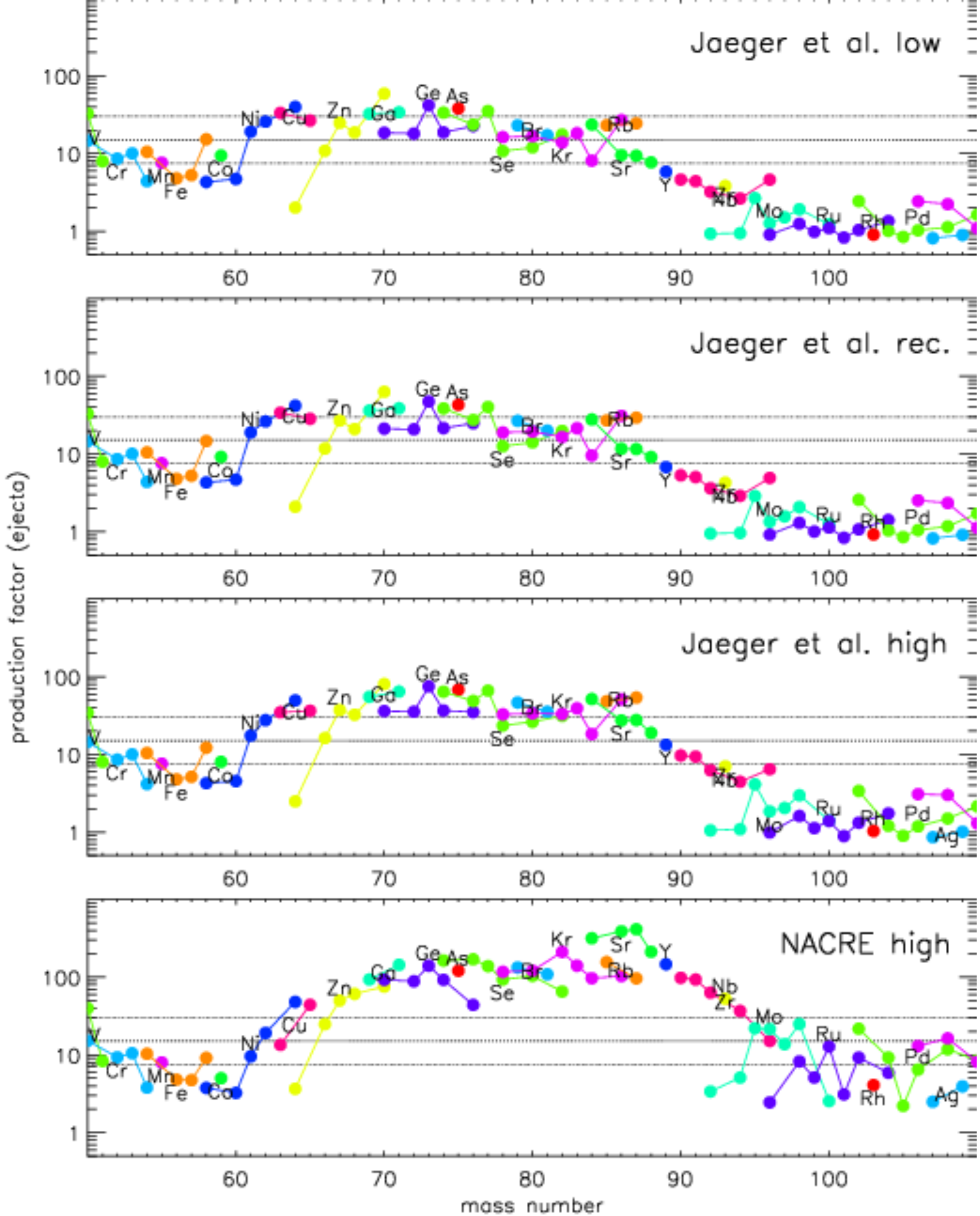
This is about 4 - 8 *neutrons per iron* and obviously not nearly enough to change e.g., Fe into Pb, but the neutron capture cross sections of the isotopes generally increase above the iron group and the solar abundances decrease. A significant s-process occurs that produces significant quantities of the isotopes with A < 90.

Composition of a 25 solar mass star at the end of helium burning compared with solar abundances (Rauscher et al 2001)



At the end ....

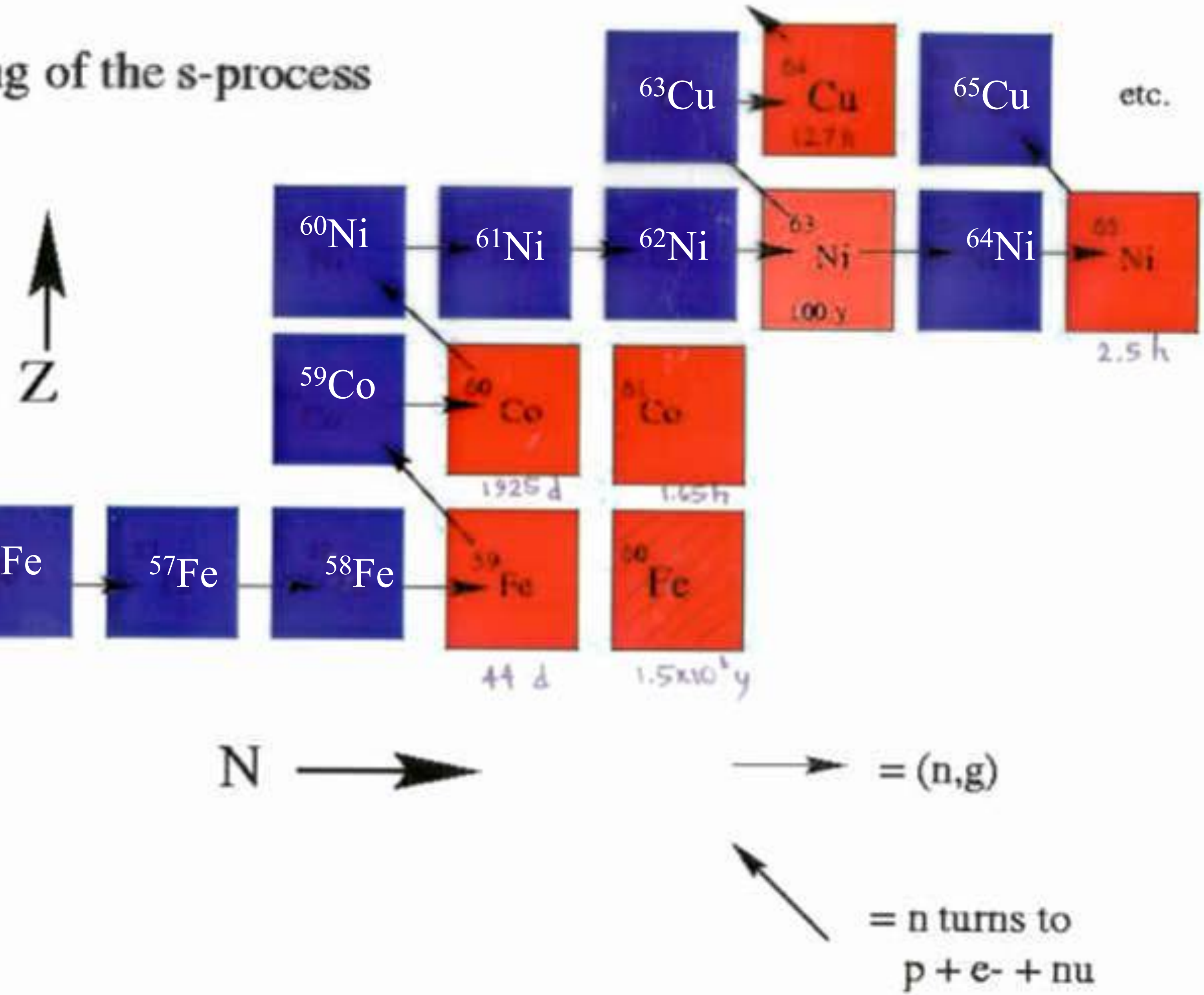

25 solar mass supernova model
post-explosion
decayed production factor (solar)
mass number
10^3
10^2
10^1
1
10^-1
10^-2
0
20
40
60
80
100
120
H
He
Li
Be
B
C
N
O
F
Ne
Na
Mg
Al
Si
P
S
Cl
Ar
K
Ca
Sc
Ti
V
Cr
Mn
Fe
Co
Ni
Cu
Zn
Ga
Ge
As
Se
Br
Kr
Rb
Sr
Y
Zr
Nb
Mo
Ru
Rh
Pd
Ag
Cd
In
Sn



Results depend most sensitively upon the reaction rate for $^{22}Ne(\alpha,n)^{25}Mg$.

If ^{22}Ne does not burn until later (i.e., carbon burning there are much more abundant neutron poisons

Beginning of the s-process



Here "s" stands for "slow" neutron capture

$$\boxed{\tau_\beta \ll \tau_{n\gamma}} \qquad \tau_\beta = \frac{1}{\lambda_\beta} = \frac{\tau_{1/2}}{\ln 2}$$

$$\frac{dY_i}{dt} = -Y_i Y_n \rho \lambda_{n\gamma}(i) + \ldots$$

$$\tau_{n\gamma} = \left(\rho Y_n \lambda_{n\gamma}\right)^{-1}$$

$$\tau_{n\gamma} = \left(\frac{1}{Y_i}\frac{dY_i}{dt}\right)^{-1}$$

This means that the neutron densities are relatively small

E.g. for a ^{22}Ne neutron source

$$\frac{dY_n}{dt} \approx 0 \approx \rho Y_\alpha Y(^{22}Ne)\lambda_{\alpha n}(^{22}Ne)$$

$$- \quad \rho Y_n Y(^{25}Mg)\lambda_{n\gamma}(^{25}Mg)$$

$$\rho \approx 1000,\ X_\alpha \approx 0.5,\ X(^{22}Ne) \approx 0.005,\ X(^{25}Mg) \approx 0.005$$

$$Y_n \approx \frac{Y_\alpha Y(^{22}Ne)\lambda_{\alpha n}(^{22}Ne)}{Y(^{25}Mg)\lambda_{n\gamma}(^{25}Mg)} \sim \frac{Y_\alpha \lambda_{\alpha n}(^{22}Ne)}{\lambda_{n\gamma}(^{25}Mg)}$$

$$Y_\alpha \sim 0.02/4;\ Y(^{22}\mathrm{Ne}) \sim 2\ Y(^{25}\mathrm{Mg});\ \rho = 1000$$

T_8	$\lambda_{\alpha n}(^{22}Ne)$	$\lambda_{n\gamma}(^{25}Mg)$	$n_n = \rho N_A Y_n$	$\lambda_{n\gamma}(^{56}Fe)$
2.0	9.1(–17)	1.2(6)	*negligible*	
2.5	1.5(–13)	1.1(6)	$\sim 10^6$	1.9(6)
3.0	2.6(–11)	1.0(6)	$\sim 10^8$	1.9(6)

Most of the s-process takes place around T_8 = 2.5 - 3, so the neutron density is about 10^6 - 10^8 cm^{-3} (depends on uncertain rate for (α,n) on ^{22}Ne and on how much ^{22}Ne has burned).

At these neutron densities the time between capture, even for heavy elements with bigger cross sections than iron, is days. For ^{56}Fe itself it is a few years

Eg. at $T_8 = 2.5$ ($n_n \sim 10^6$), the lifetime of ^{56}Fe is about

$$\tau(^{56}Fe) = \left(\frac{1}{Y(^{56}Fe)} \frac{dY(^{56}Fe)}{dt} \right)^{-1} = \left(\rho Y_n \lambda_{n\gamma}(^{56}Fe) \right)^{-1}$$

$$\approx 10^4 \text{ year (less at higher temperatures)}$$

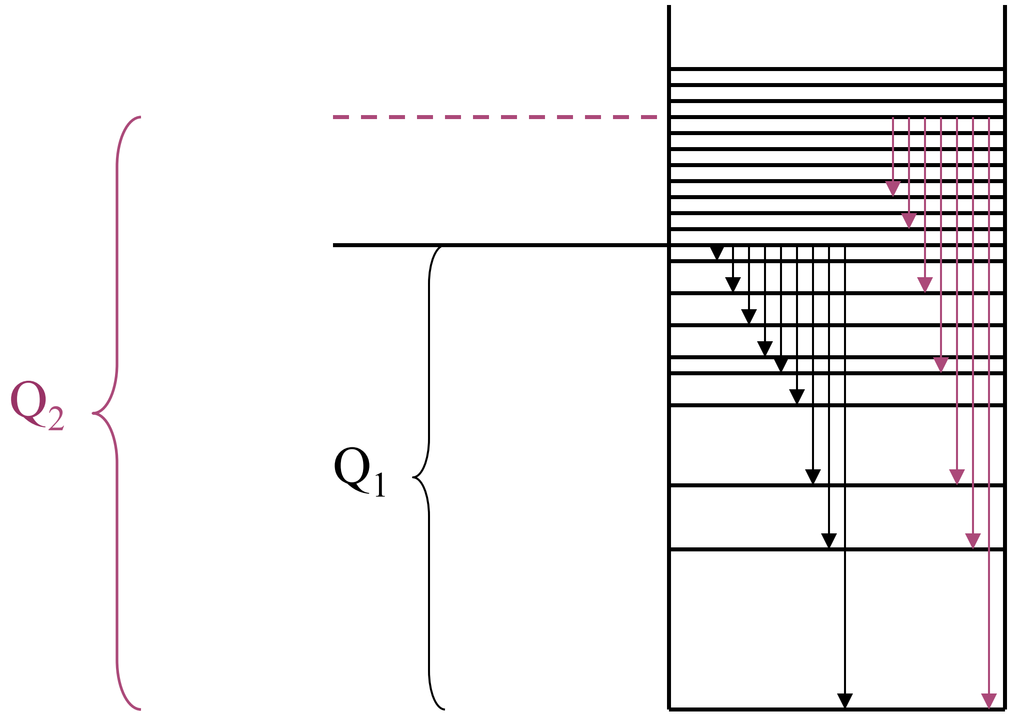
The s-process in massive stars only goes on during a brief period at the end of helium burning. The time scale is lengthened by convection.

Reaction Rates (n,γ):

Either measured (Bao et al, ADNDT, 76, 70, 2000)
or calculated using Hauser-Feshbach theory
(Woosley et al., ADNDT, 22, 371, (1976)
Holmes et al., ADNDT, 18, 305, (1976);
Rauscher et al. ADNDT, 75, 1, (2000))

The calculations are usually good to a factor of two. For heavy nuclei within $kT \sim 30$ keV of Q_{ng} there are very many resonances.

Occasionally, for light nuclei or near closed shells, direct capture is important: e.g., ^{12}C, $^{20,22}Ne$, ^{16}O, ^{48}Ca



More levels to make transitions to at higher Q and also, more phase space for the outgoing photon.

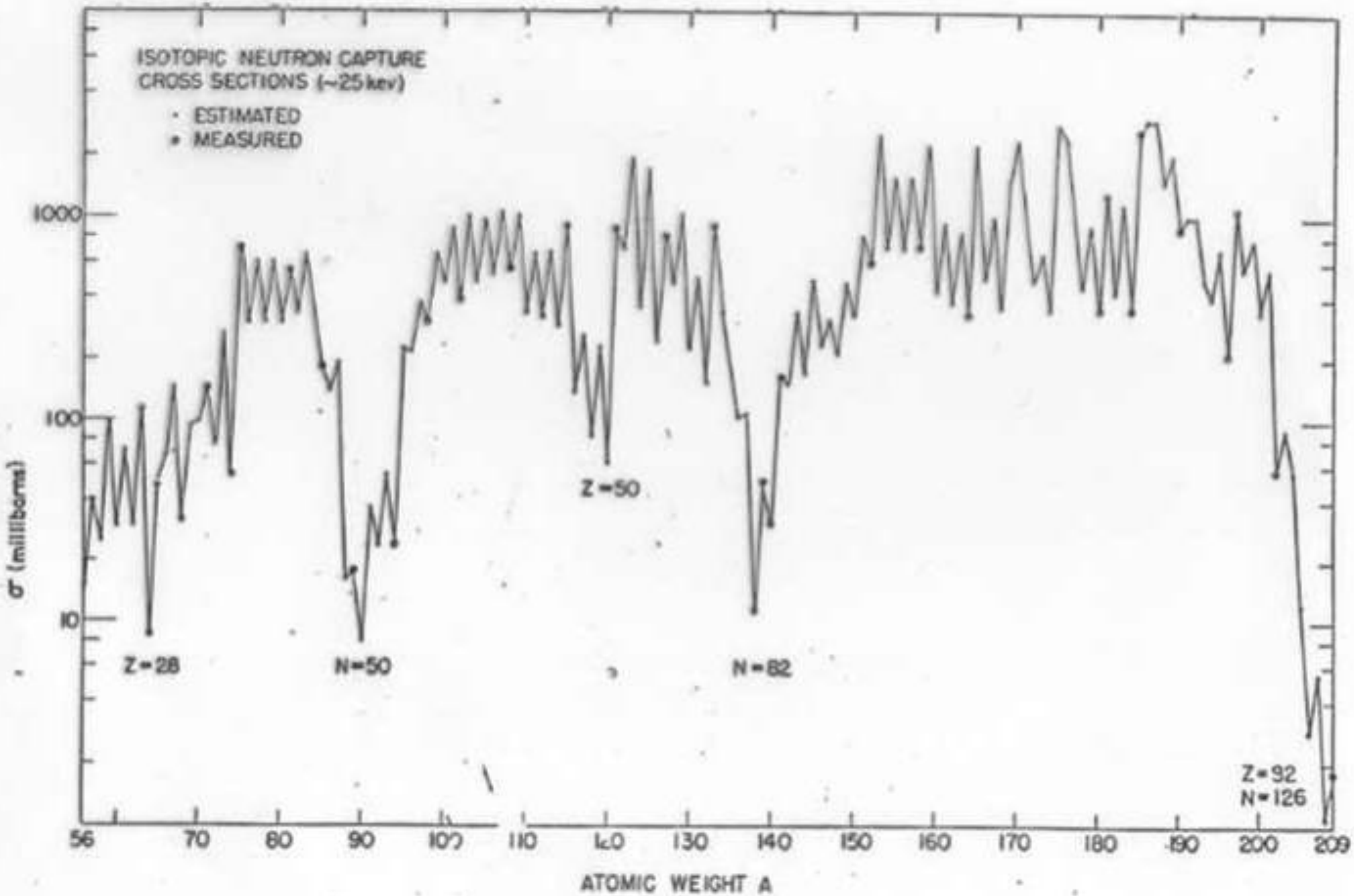
E_γ^3 for electric dipole

$$\boxed{T_\gamma(Q_2) > T_\gamma(Q_1)}$$

and as a result

$$\sigma_{n\gamma} \propto \frac{T_n \; T_\gamma}{T_n + T_\gamma} \approx T_\gamma$$

is larger if Q is larger



ISOTOPIC NEUTRON CAPTURE
CROSS SECTIONS (~25 kev)
· ESTIMATED
• MEASURED
σ (millibarns)
1000
100
10
Z=28
N=50
Z=50
N=82
Z=92
N=126
56
70
80
90
110
130
140
150
160
170
180
190
200
209
ATOMIC WEIGHT A

Rate Equations: Their Solutions and Implications

Assume constant density, temperature, cross section, and neutron density and ignore branching (would never assume any of these in a modern calculation). Then

$$\frac{dY(^A Z)}{dt} \equiv \frac{dY_A}{dt} = -Y_A Y_n \rho \lambda_{n\gamma}(A) + Y_{A-1} Y_n \rho \lambda_{n\gamma}(A-1)$$

and since $n_n = \rho N_A \; Y_n$ and $\lambda_{\mathrm{n}\gamma} = N_A \left\langle \sigma_{n\gamma} v \right\rangle \approx N_A \sigma_A v_{thermal}$

defining $\tau \equiv \int \rho N_A Y_n \, v_{thermal} \, dt = \int n_n \, v_{thermal} \; dt$, one has

$$\boxed{\frac{dn_A}{d\tau} = -n_A \sigma_A + n_{A-1} \sigma_{A-1}}$$

Note that τ has units of inverse cross section (inverse area).

If there were locations where steady state is achieved then

$$\frac{dn_A}{d\tau} \approx 0 = n_A \sigma_A - n_{A-1} \sigma_{A-1}$$

$i.e.,$ σn is locally a constant, and n $\propto \dfrac{1}{\sigma}$

Attaining steady state requires a time scale longer than a few times the destruction lifetime of the species in the steady state group. One has "local" steady state because any flux that would produce, e.g., lead in steady state would totally destroy all the lighter s-process species.

The flow stagnates at various "waiting points" along the s-process path, particularly at the closed shell nuclei.

Eg., $n_n \sim 10^8 \text{ cm}^{-3} \Rightarrow \rho Y_n \sim n_n / N_A \sim 5 \times 10^{-17}$

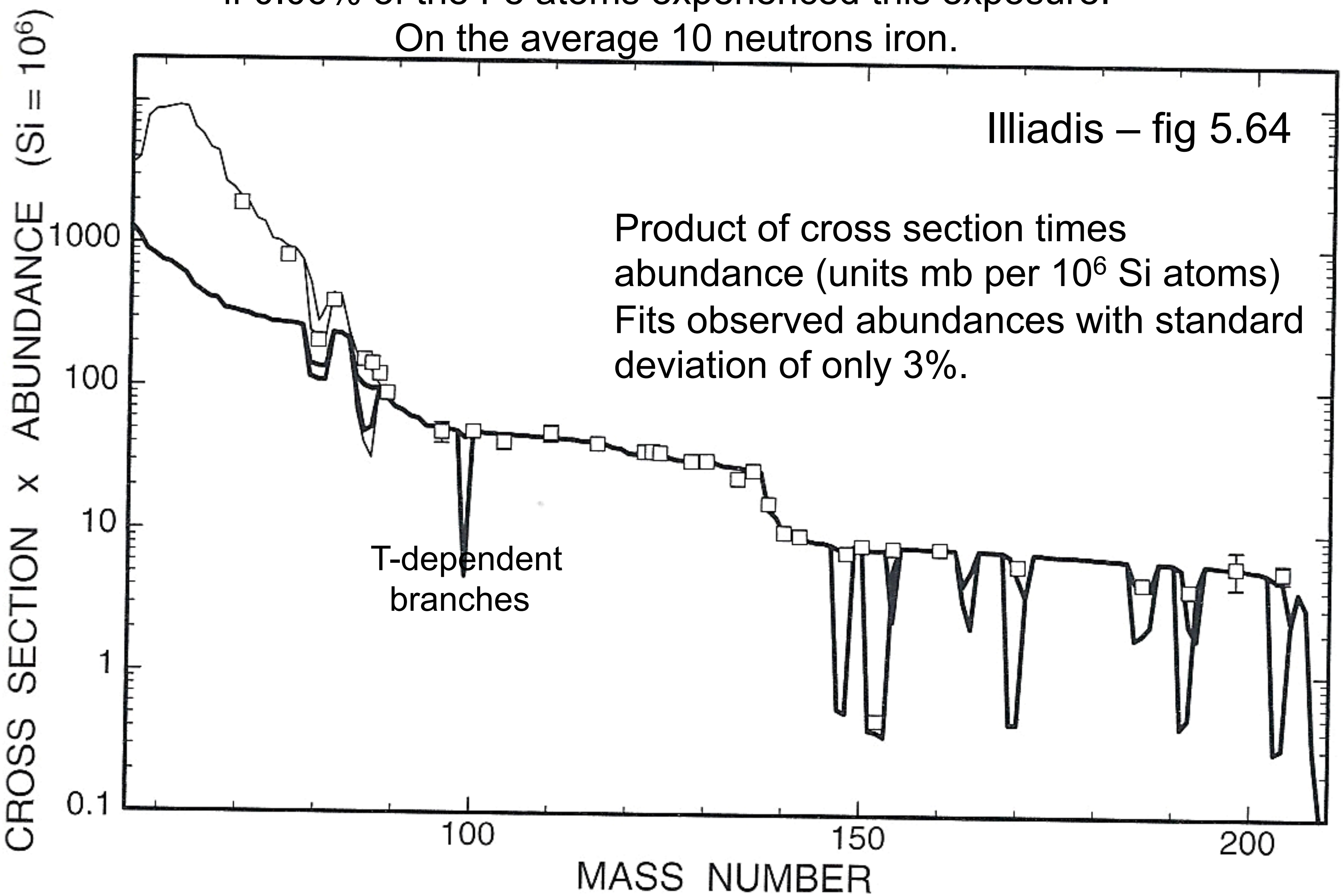
$\lambda_{n\gamma}$ experimentally at helium burning temperatures is $10^5 - 10^8$

$$\tau_{n\gamma} = \left(\rho Y_n \lambda_{n\gamma}\right)^{-1} = \left(\frac{d \ln Y_A}{dt}\right)^{-1} \sim 10 - 10^4 \text{ years}$$

This can be greatly lengthened in a massive star by convection.

As a result nuclei with large cross sections will be in steady state while those with small ones are not. This is especially so in He shell flashes in AGB stars where the time scale for a flash may be only a few decades.

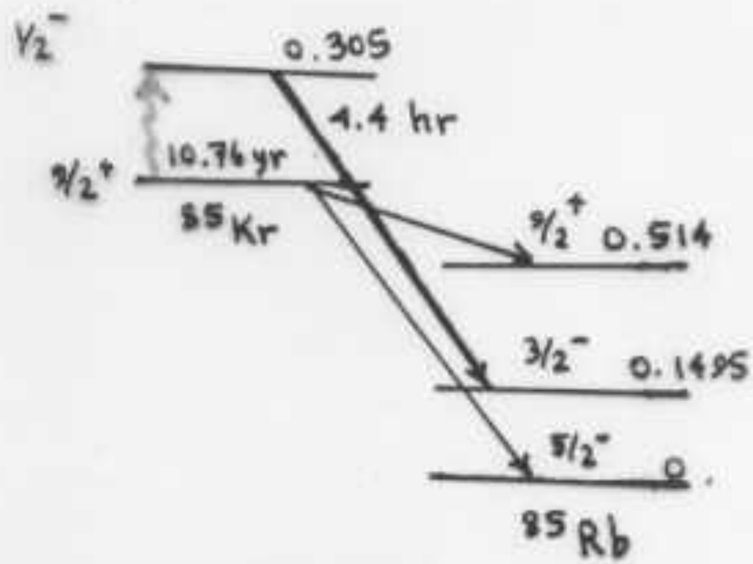
Can make the main s-process component in the sun
if 0.06% of the Fe atoms experienced this exposure.
On the average 10 neutrons iron.



The symbols are s-only nuclei. The solid lines are the model results for a standard (exponential) set of exposure strengths. Below A = 90 there is evidence for a separate additional s-process component.

The beta decay rates too are <u>usually</u> insensitive to T. However, occasionally

λ's about equal @ KT = 30 KeV



Thermally populating 1st excited state of ^{85}Kr may greatly accelerate its decay

Other T-sensitive branch points:

^{93}Zr, ^{113}Cd, ^{134}Cs, ^{151}Sm, ^{163}Dy, ^{170}Tm

^{176}Lu, ^{163}Ho, ^{79}Se, ^{80}Br, ^{152}Eu, ^{154}Eu, ^{185}W

These can be analysed to estimate the temperature of the s-process.

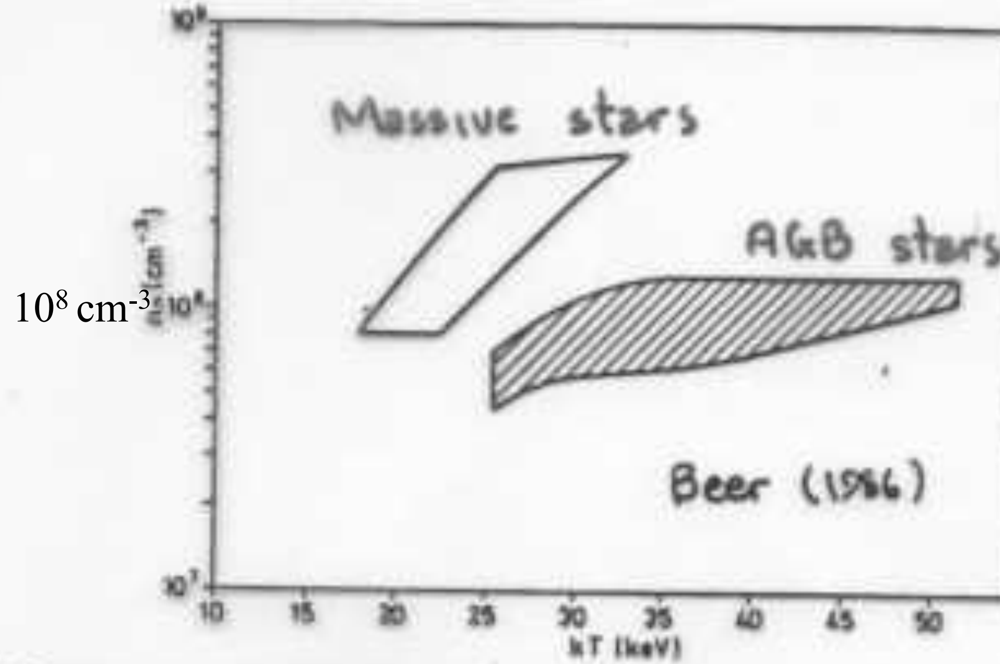


Fig.2 s-process neutron density as a function of s-process temperature. Consistent values for the main and weak component lie within the indicated areas [8]. The area of the weak component is hatched. The fact that there is no common range of values for the two components is interpreted as a strong evidence for their independance.

Implicit solution:

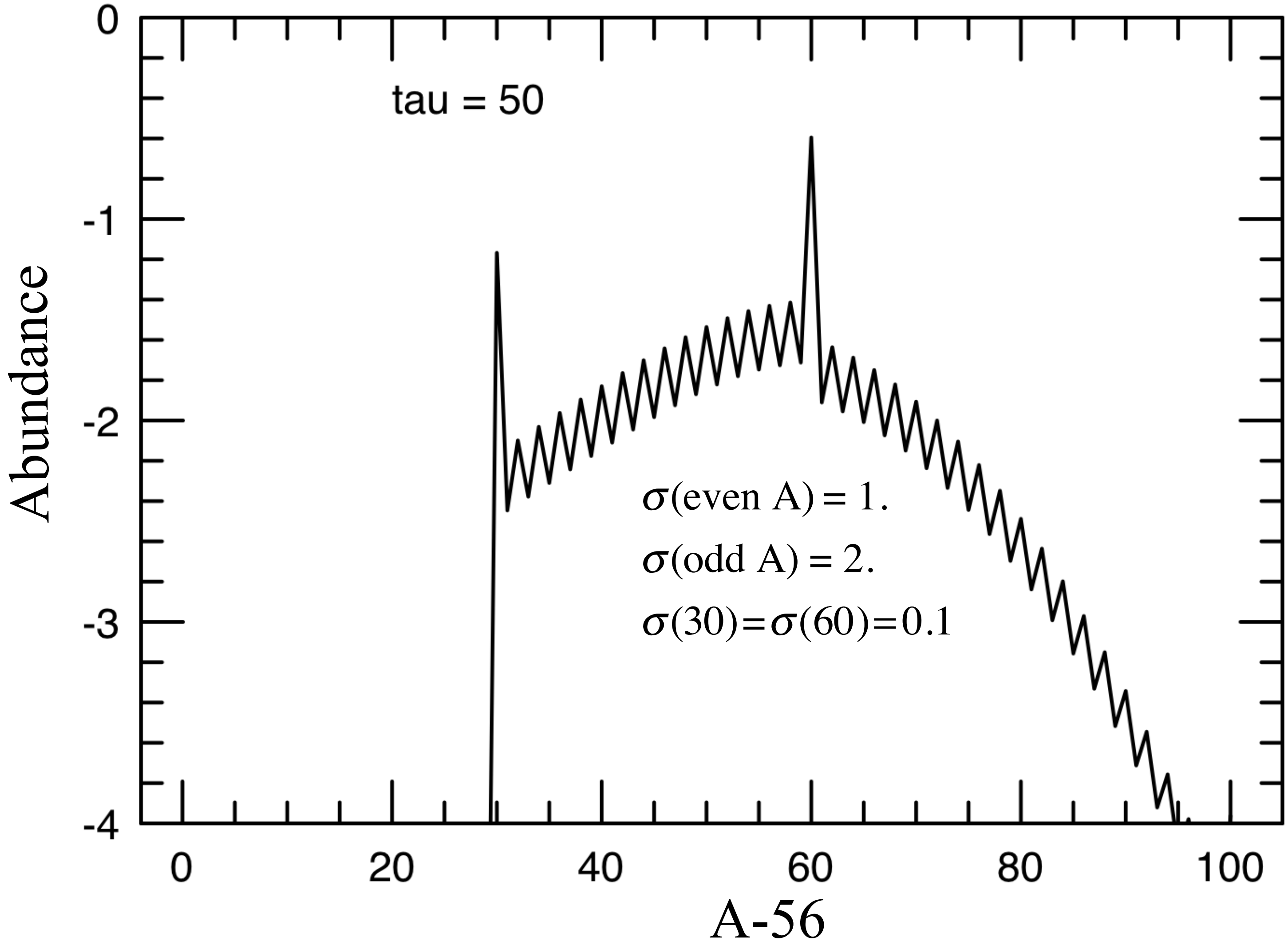
Assuming no flow downwards from A+1 and greater to A and below.

$$n_{new}(A) = \frac{n_{old}(A)/d\tau + n_{new}(A-1)\,\sigma_{n\gamma}(A-1)}{1/d\tau + \sigma_{n\gamma}(A)}$$

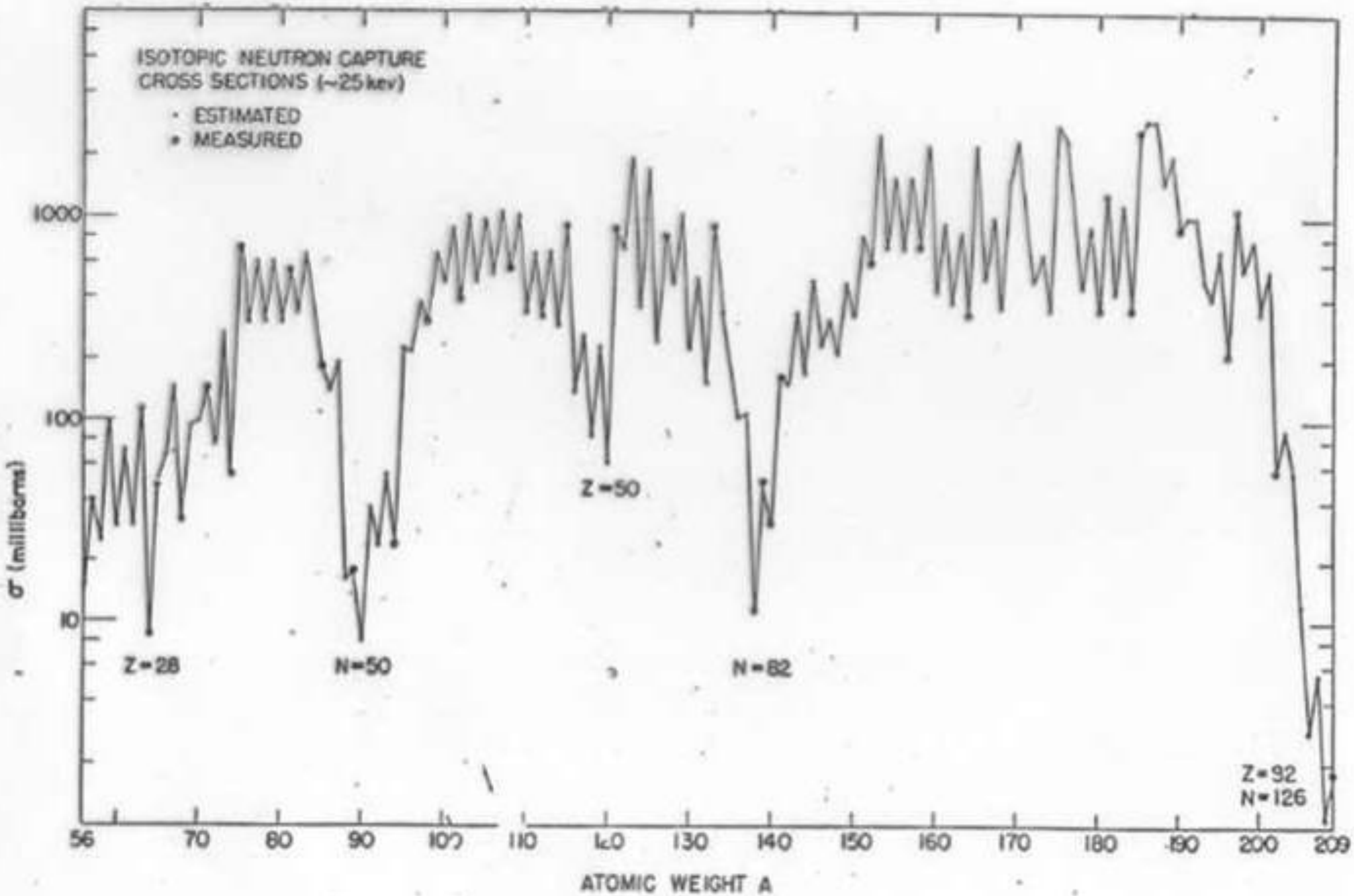
This works because in the do loop, $n_{new}(A-1)$ is updated to its new value before evaluating $n_{new}(A)$. Matrix inversion reduces to a recursion relation.

$$\frac{n_{new}(A) - n_{old}(A)}{d\tau} = -n_{new}(A)\sigma_A + n_{new}(A-1)\sigma_{A-1}$$

$n_{new}(A-1)$ known from previous step in recursion relation



Sample output from toy model code micros2.f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56
70
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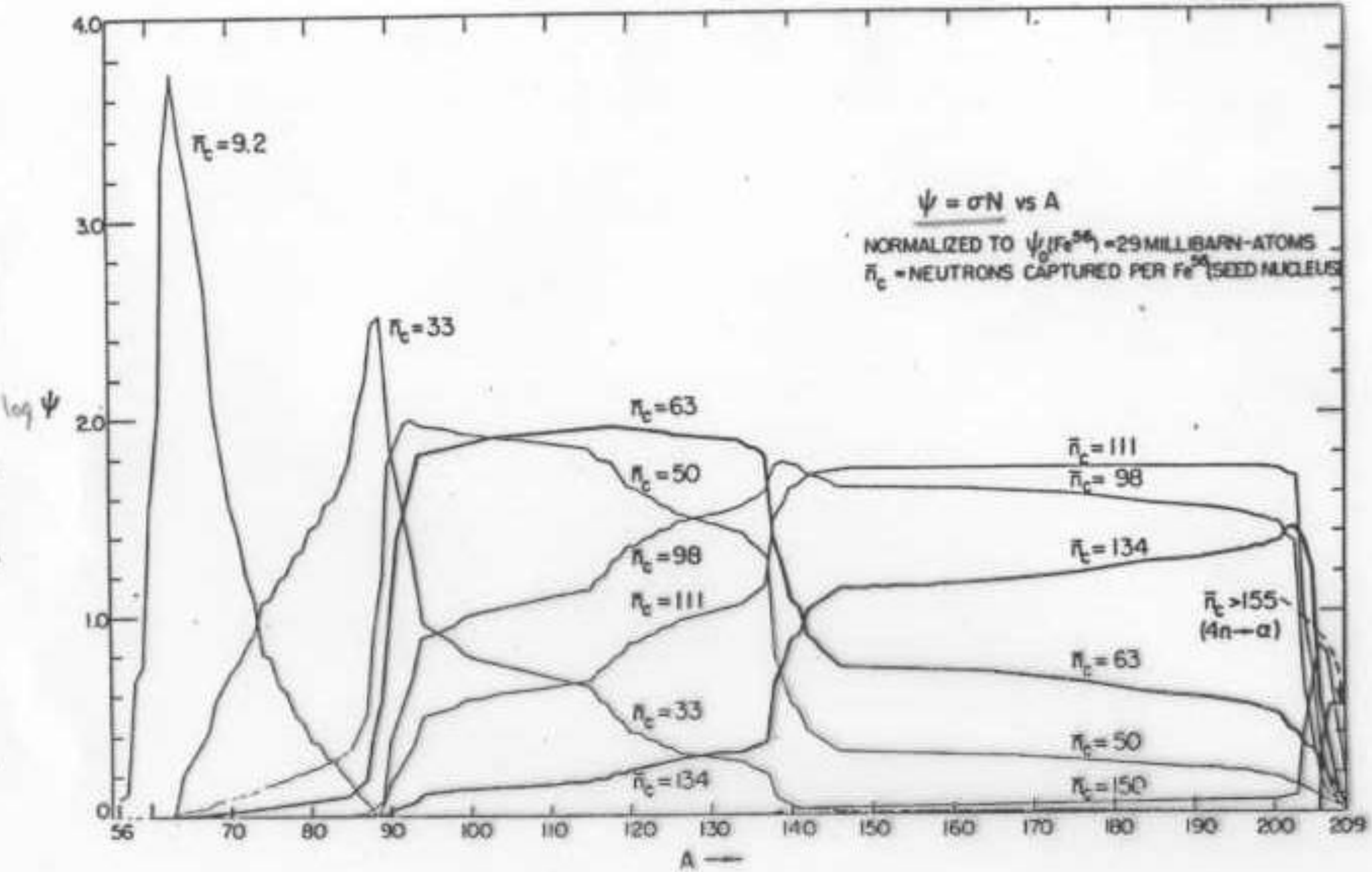
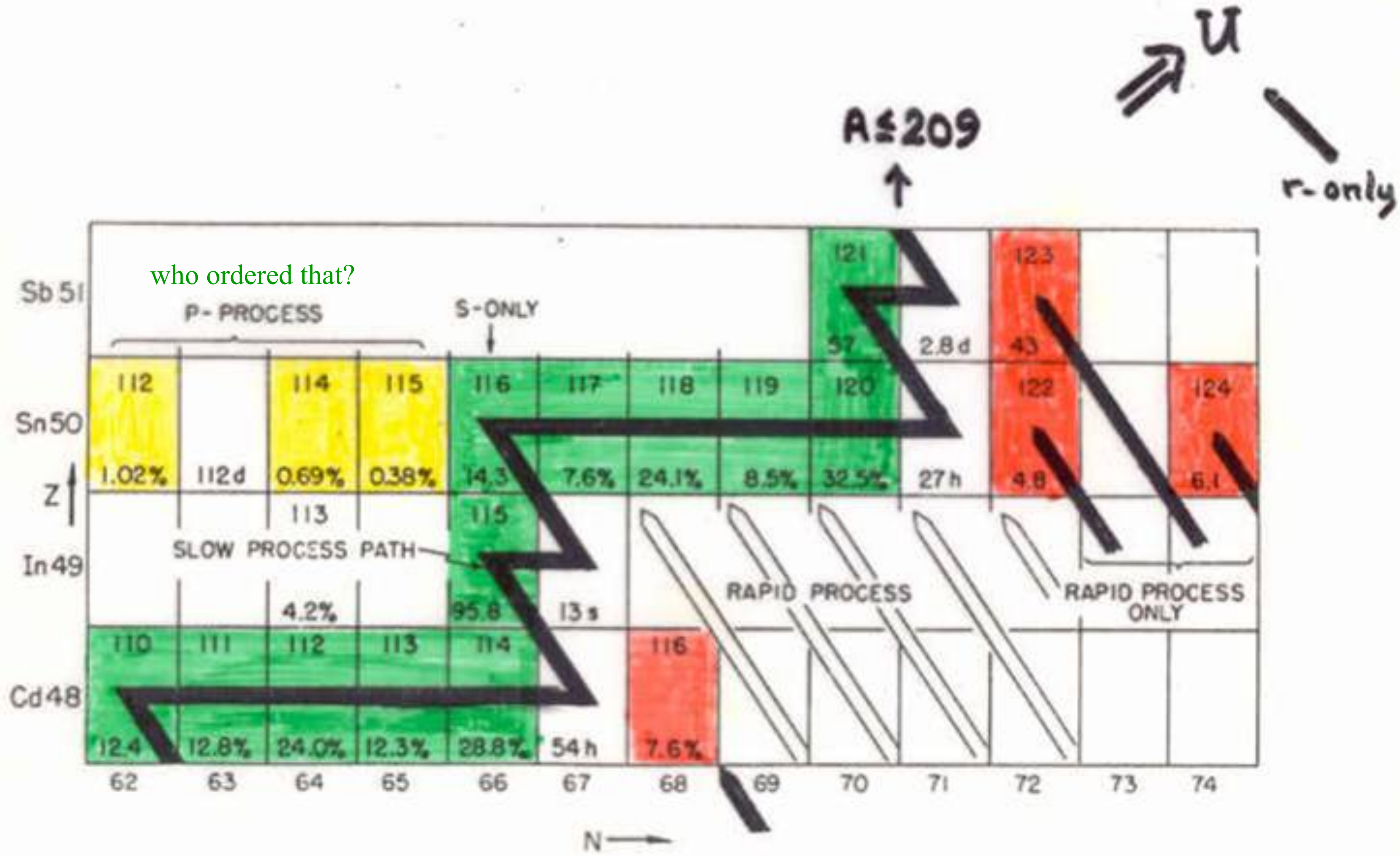
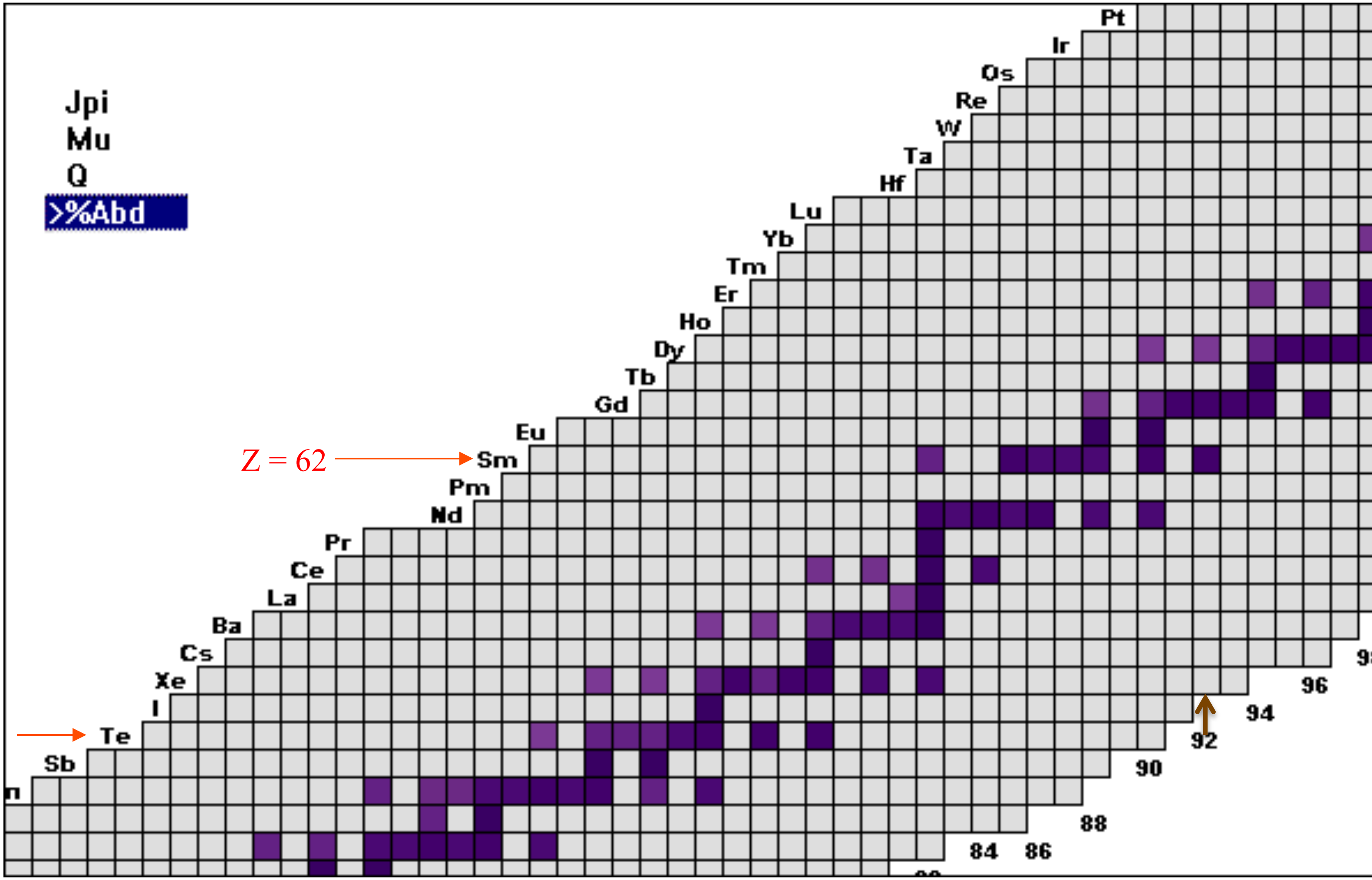
180
190
200
209
ATOMIC WEIGHT A

Fig 7-23 in Clayton gives $\bar{n}_c(\tau)$





e.g., ^{117}Sn, ^{118}Sn, ^{119}Sm, and ^{120}Sn are s,r isotopes. Sn is not a good place to look for s n = const though because it is a closed shell.



Jpi
Mu
Q
>%Abd
Z = 62
Pt
Ir
Os
Re
W
Ta
Hf
Lu
Yb
Tm
Er
Ho
Dy
Tb
Gd
Eu
Sm
Pm
Nd
Pr
Ce
La
Ba
Cs
Xe
I
Te
Sb
98
96
94
92
90
88
84
86

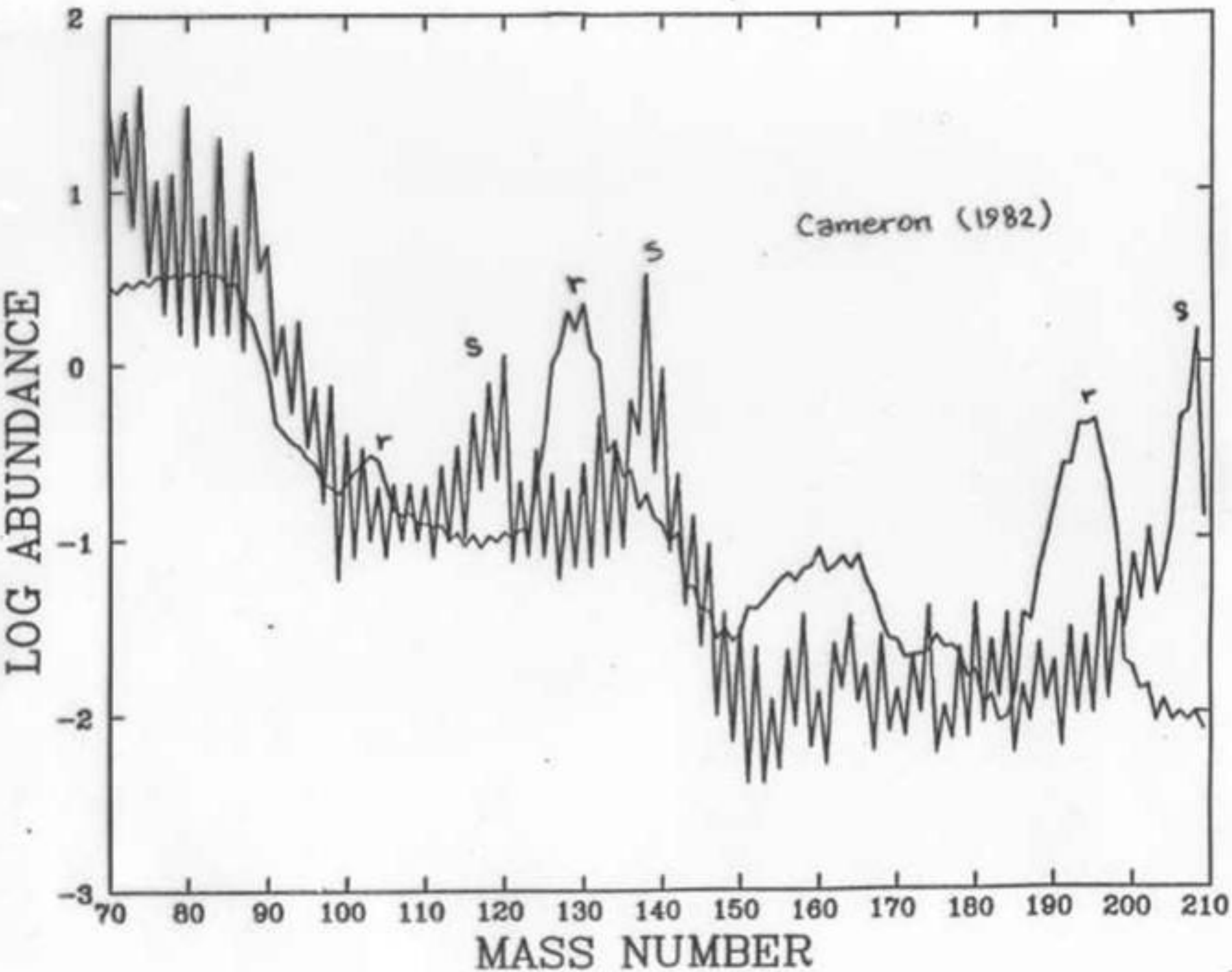
See http://www.nndc.bnl.gov/chart/

similarly for Sm

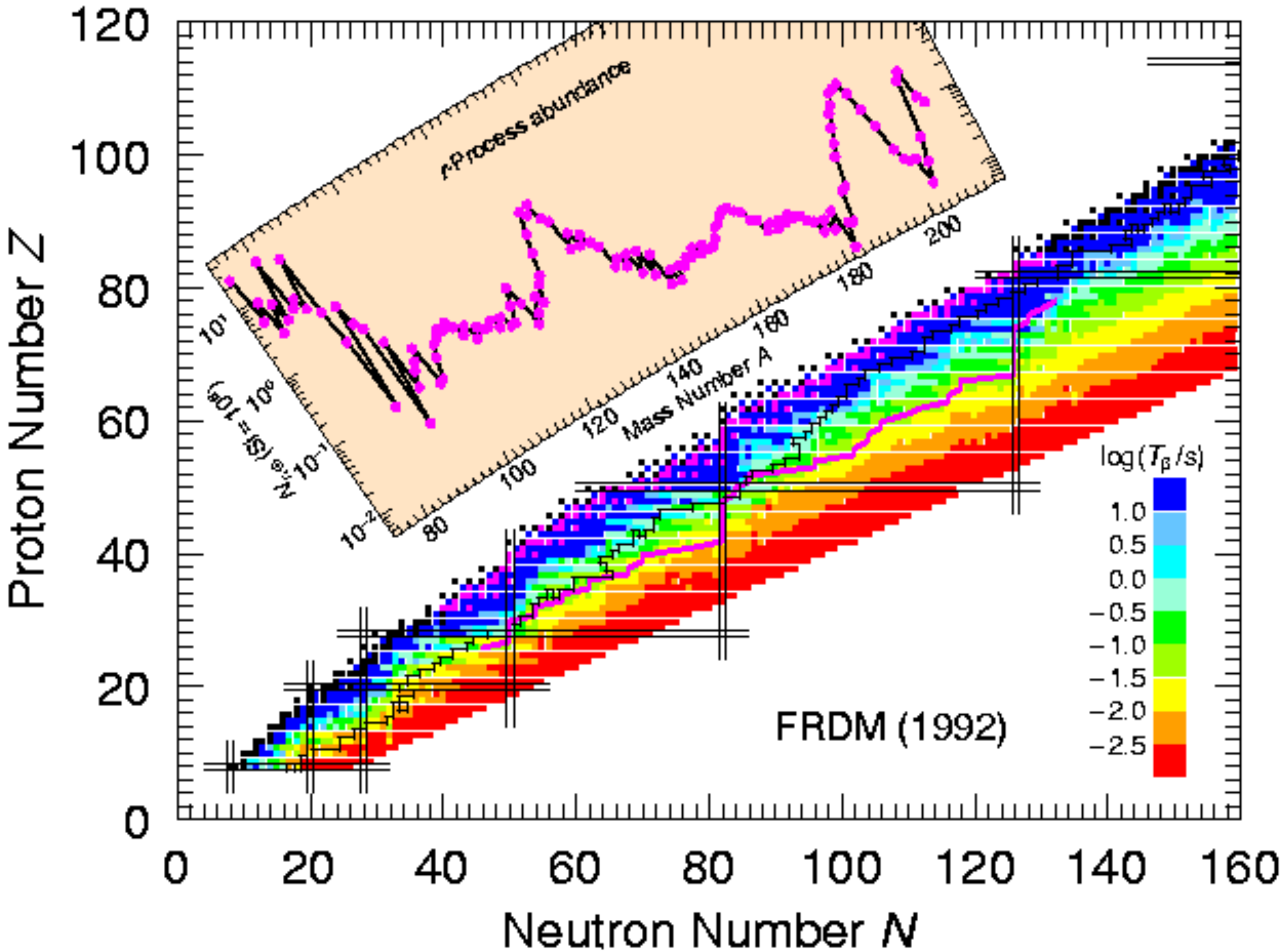
			σ(mb)	n	σn
	^{144}Sm	p	92 ± 9	7.42(-3)	0.68
	^{147}Sm	r, s	1000 ± 100	3.71(-2)	37 ± 4
Z = 62	^{148}Sm	s	267 ± 12	2.70(-2)	7.2 ± 0.3 ←
	^{149}Sm	r, s	1454 ± 66	3.32(-2)	48.3 ± 2.2
	^{150}Sm	s	447 ± 26	1.79(-2)	8.0 ± 0.5 ←
	^{152}Sm	s, r	375 ± 23	6.41(-2)	24.2 ± 1.5
	^{154}Sm	r	293 ± 19	5.45(-2)	16.0 ± 1.0

$\tau_{1/2}(^{151}\mathrm{Sm}) = 90$ years; $\tau_{1/2}(^{153}\mathrm{Sm}) = 46$ hours

Based upon the abundances of s-only isotopes and the known neutron capture cross sections one can subtract the s-portion of s,r isotopes to obtain the r- and s-process yields separately.



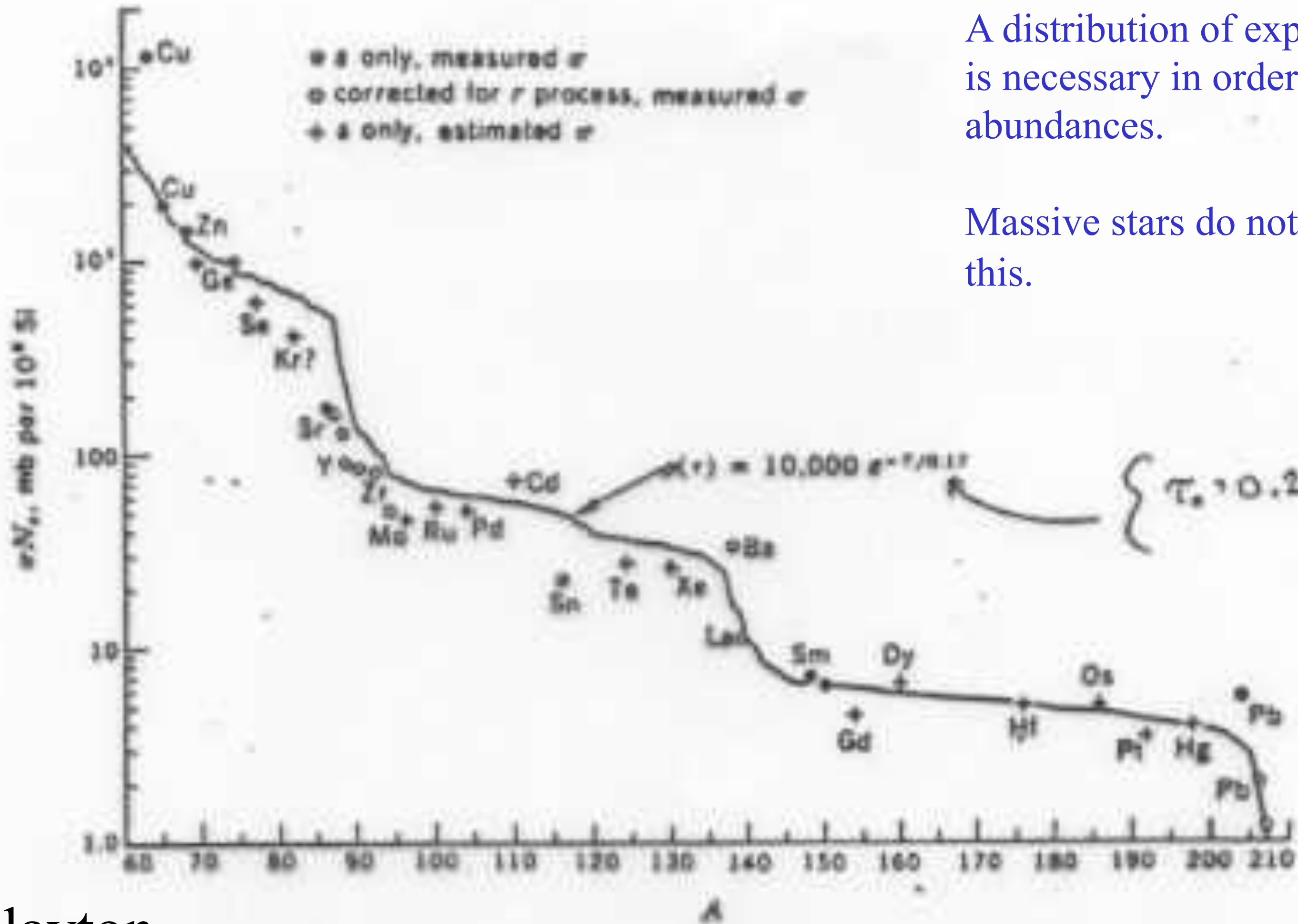
Cameron (1982)
s
r
r
s
s
r
s
LOG ABUNDANCE
2
1
0
−1
−2
−3
70
80
90
100
110
120
130
140
150
160
170
180
190
200
210
MASS NUMBER



r-Process abundance
Mass Number A
$N_{r,\odot}$ (Si = 10^6)
10^1
10^0
10^{-1}
10^{-2}
80
100
120
140
160
180
200
Proton Number Z
0
20
40
60
80
100
120
Neutron Number N
0
20
40
60
80
100
120
140
160
log (T_β/s)
1.0
0.5
0.0
−0.5
−1.0
−1.5
−2.0
−2.5
FRDM (1992)

A distribution of exposure strengths is necessary in order to get the solar abundances.

Massive stars do not naturally give this.

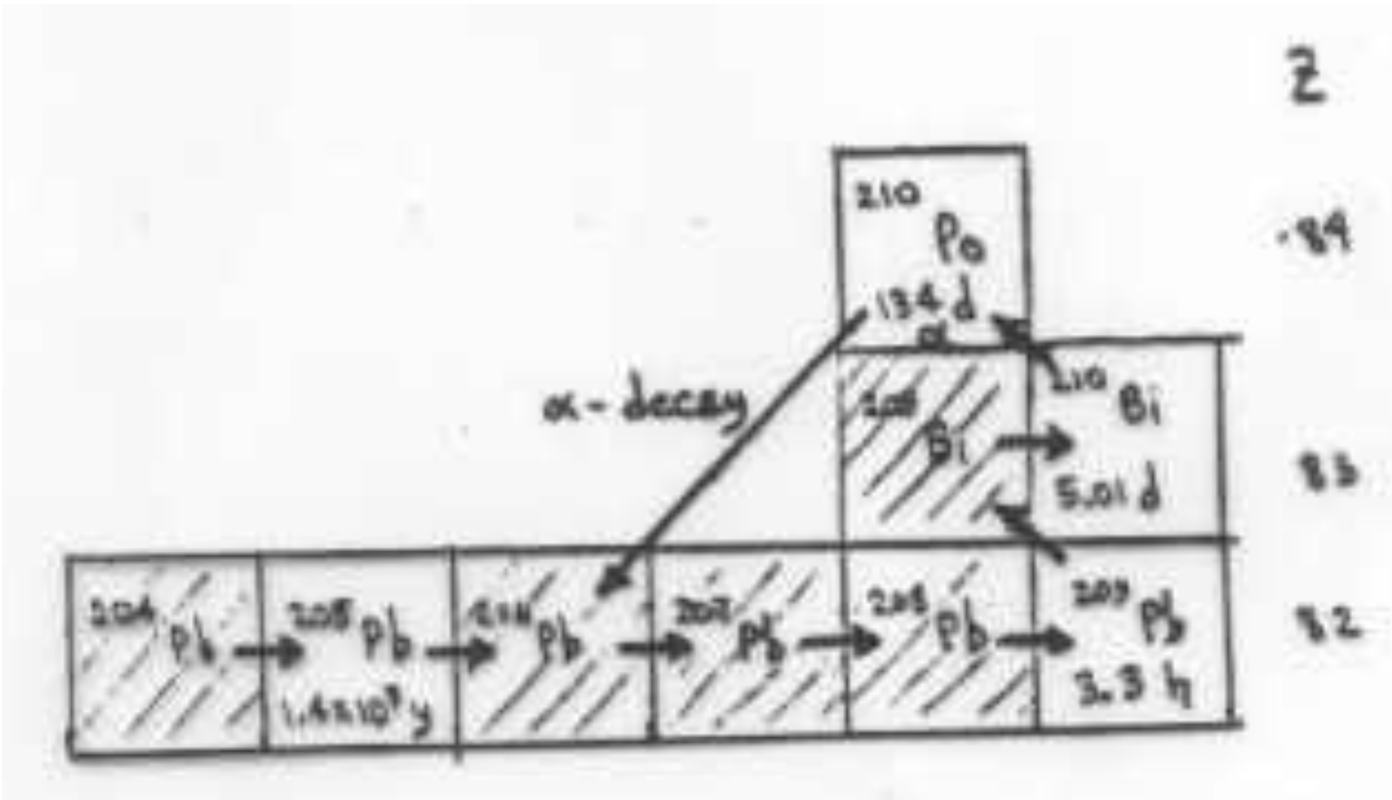


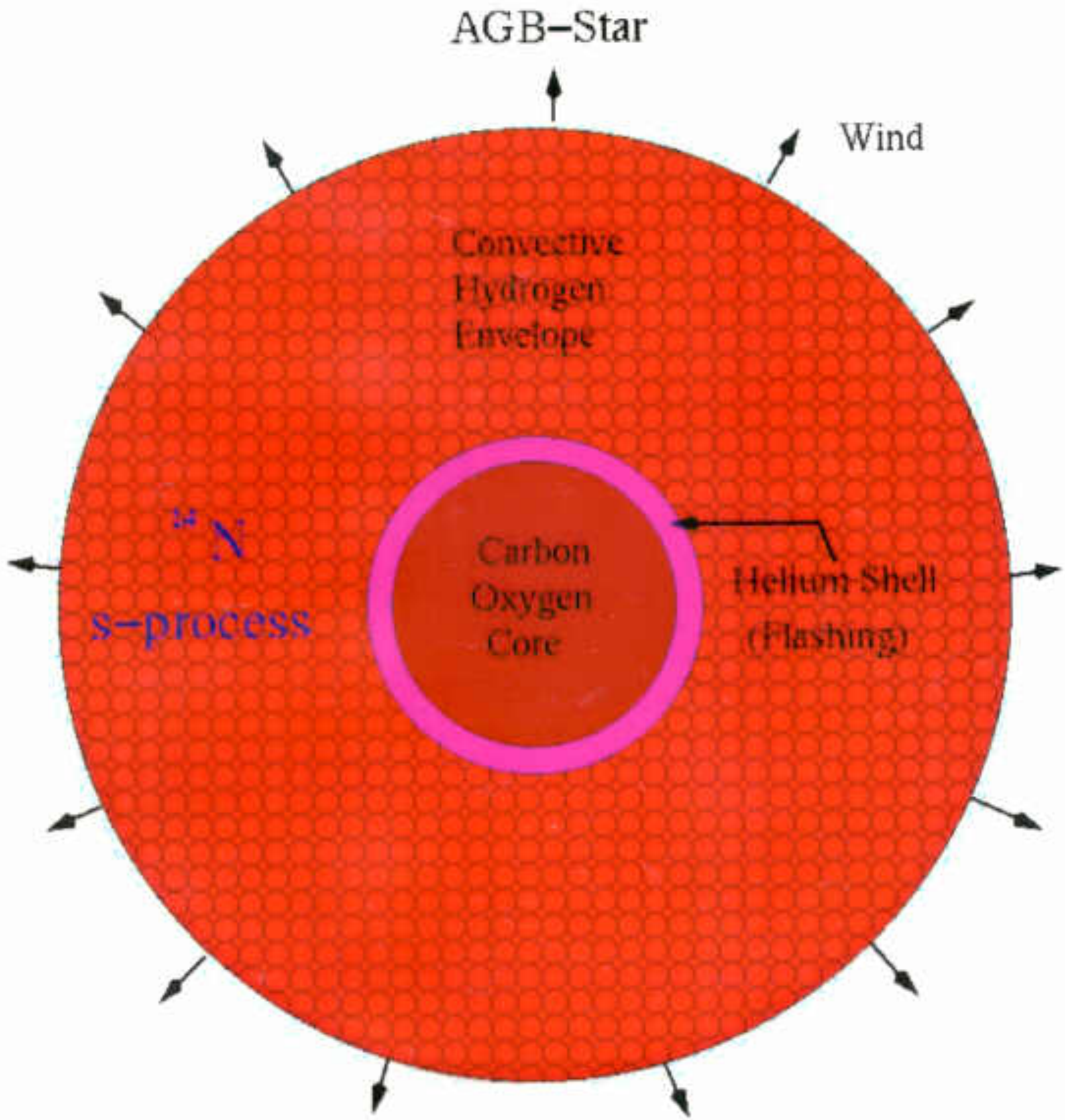
$\tau_0 \approx 0.25 \text{ mb}^{-1}$ a little better

Clayton

Fig. 7-20 The solar-system σN_s curve. The product of the neutron-capture cross sections for $kT = 30$ kev times the nuclide abundance per 10^6 silicon atoms is plotted versus the atomic mass number A. The solid curve is the calculated result of an exponential distribution of neutron exposures. [*P. A. Seeger, W. A. Fowler, and D. D. Clayton, Astrophys. J. Suppl.*, 11:121 (1965). *By permission of The University of Chicago Press. Copyright* 1965 *by The University of Chicago.*]

Termination of the s-Process




AGB−Star
Wind
Convective
Hydrogen
Envelope
^{14}N
s−process
Carbon
Oxygen
Core
Helium Shell
(Flashing)

Formation of an AGB star - after core helium burning in a star lighter than about 8 solar masses

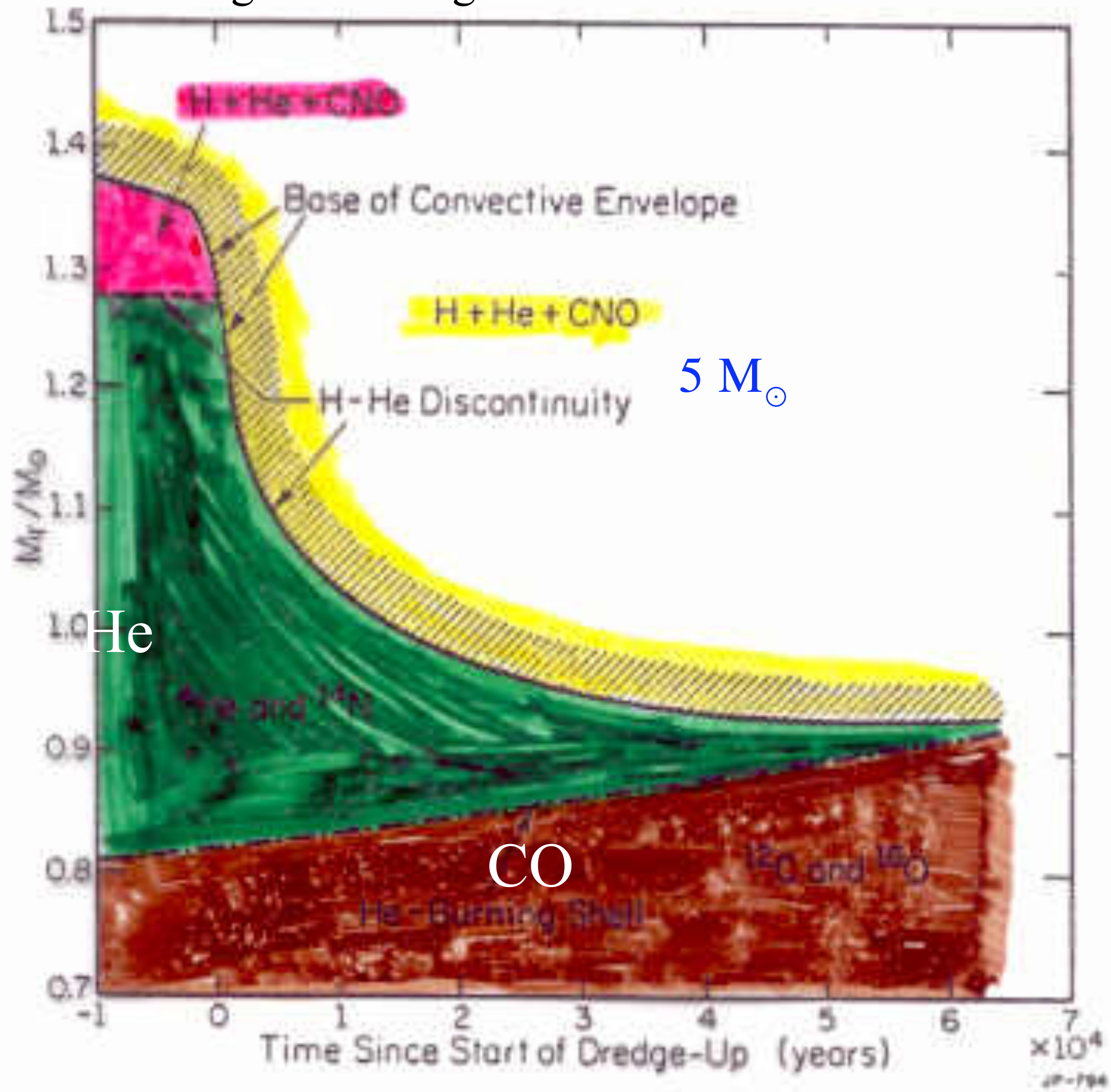


Figure 1 The variation with time of the location (in mass) of the base of the convective envelope and of the center of the helium-burning shell during the second dredge-up phase in a model of mass $5\ M_\odot$ and initial composition $(Y, Z) = (0.28, 0.001)$.

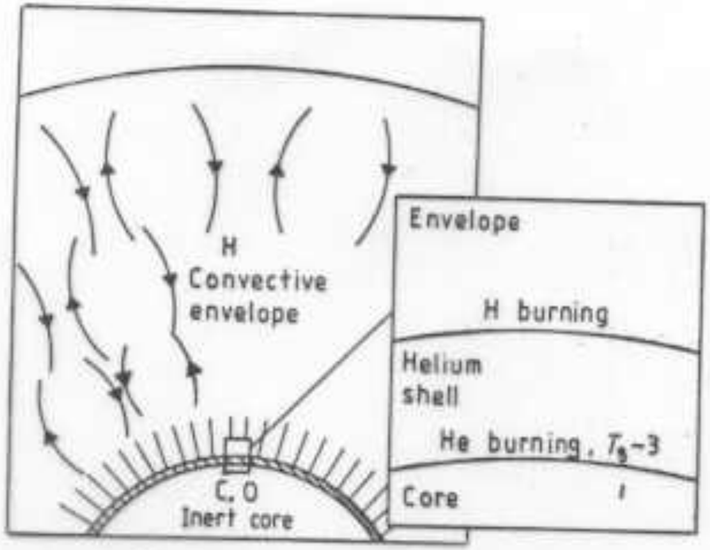
$$\frac{dP}{dm} = \frac{GM(R)}{4\pi R^4}$$

Thin shell instability:

If a shell is sufficiently thin, its pressure is set by the gravitational potential in which it rests. During a flash, this does not change. Burning raises the temperature but does not change the pressure so the density in the shell goes down. But the energy generation is very sensitive to the temperature and continues to go up. Finally sufficient burning occurs to cause enough overall expansion to reduce the pressure. This instability tends to happen at the edges of compact objects where the burning shells are quite thin compared with the radius of the core and the gravitational potential is thus constant.

Schwarzshild & Harm (1965) and Weigert (1966),
Yoon, Langer and van der Sluys, *A&A*, **425**, 207 (2004)

Models by Icko Iben

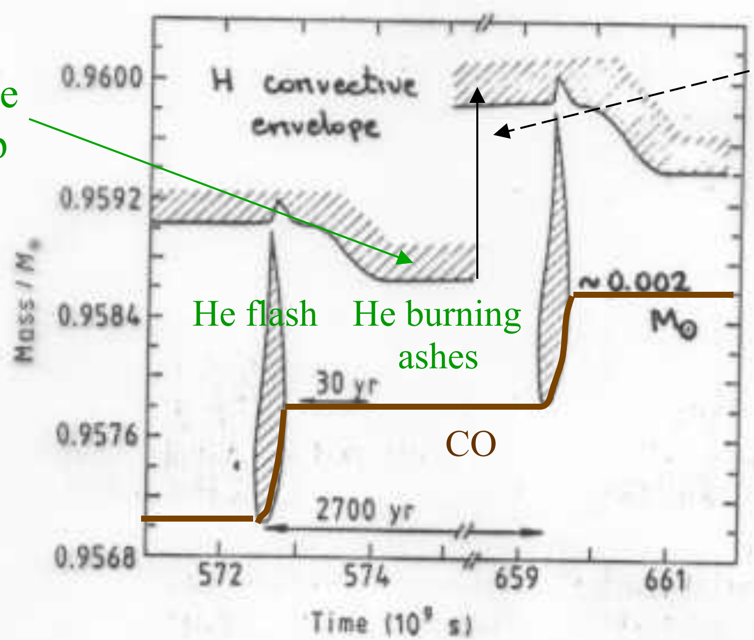


Schematic structure of a $7M_\odot$ mass star during shell burning. For illustration, the helium shell is greatly exaggerated. Helium shell burning occurs in pulses, which are triggered when the CNO cycle at the bottom of the envelope has produced a critical amount of helium.

Part of the ashes of each helium shell flash are incorporated into the fuel for the next flash.

This naturally gives a power law distribution of exposures

Convective dredge-up



H-shell burning extends helium layer in preparation for the next flash

During each flash have a mixture of He, C, ^{13}C or ^{22}Ne, new seed nuclei and s-process from prior flashes

Important nuclear physics modification:

s-process giants derived from AGB stars in the solar neighborhood do not show the large ^{26}Mg excesses one would expect if the neutron source were ^{22}Ne(a,n)^{25}Mg [as it surely is in massive stars]. Moreover these stars are too low in mass for ^{22}Ne(a,n)^{25}Mg to function efficiently. A different way of making neutrons is required. Probably

$$^{4}He(2\alpha,\gamma)^{12}C(p,\gamma)^{13}N(e^{+}\nu)^{13}C$$

$$^{13}C(\alpha,n)^{16}O$$

with the protons coming from mixing between the helium burning shell and the hydrogen envelope. Each p mixed in becomes an n.

McWilliam and Lambert, *MNRAS*, **230**, 573 (1988) and
Malaney and Boothroyd, *ApJ*, **320**, 866 (1987)
Hollowell and Iben, *ApJL*, **333**, L25 (1988); *ApJ*, **340**, 966, (1989)
many more since then

The metallicity history of the s-process can be quite complicated.

In the simplest case in massive stars with a ^{22}Ne neutron source it is independent of metallicity until quite low values of Z.

At very low Z things can become complicated because of the effect of neutron poisons, ^{12}C and ^{16}O. In AGB stars the mixing between H and He shells is Z dependent. Some metal poor stars are actually very s-process rich. Cristallo et al (ApJ, 696, 797 (2009)).